OBSERVATIONAL ASTRONOMY

FALL 2026

Vitaly Neustroev

Contact details

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Course content

- Observational Techniques in Gamma, X-ray, UV, **Optical** and Infra-Red astronomy
- Methods of Observations with modern Space- and Ground-based Telescopes
- Observational Experiments, Calibrations and Data Reductions
- Data Analysis

Text Books

- **Astrophysical techniques** (5th Edition – 2008) – C.R. Kitchin: Taylor & Francis / CRC Press. ISBN 978-1-4200-8243-2.
- **Observational Astrophysics** (3rd Edition – 2012) – P. Léna, D. Rouan, F. Lebrun, F. Mignard, D. Pelat, Translated by S. Lyle: Springer. ISBN 978-3-642-21815-6.

- **An Introduction to Astronomical Photometry Using CCDs** (October 22, 2006) – W. Romanishin: can be downloaded from the course web-page.
- ***Observational Astronomy*** (Second Edition - 2006) – D. Scott Birney, Guillermo Gonzalez & David Oesper: Cambridge Univ. Press. ISBN 978-0-521-85370-5
- *ISSI Scientific Report Volume 9 (SR-009):* ***Observing Photons in Space*** (2010) – Edited by M.C.E. Huber, A. Pauluhn, J. Len Culhane, J. G. Timothy, K. Wilhelm and A. Zehnder. ISBN 978-92-9221-938-3.

Home exercises

- Compulsory homework sets will be assigned (return by the deadline).
  You are urged not to start these the night before they are due.
  For late exercises only one-half points will be given.
- The most important thing you can learn from homework is how to solve problems yourself. This is what you need to do to succeed in the real world.
  Therefore, please try each problem for at least 1 hour before discussing it with anyone else.
- Please, write your homework solutions in an **extremely** clear manner.
  It will not be possible to give credits for work that is not clearly explained.
  Please, show your work since this will allow partial credit to be given if you cannot solve the whole problem.
- When it is relevant, use general formulae as long as possible and only plug in numbers at the end of a problem.

Data reduction / analysis

- Multicolour photometry of a star.
  The project consists of three steps:

1. Find the optimal aperture size which allows measuring instrumental magnitudes with the highest accuracy
2. Using this aperture size, measure the instrumental magnitudes of a star and 3 comparison stars in all five fiters (*B, V, R, I, z*).
   Find the apparent magnitudes of your target.
3. Convert the magnitudes to fluxes and plot the Spectral Energy Distribution (SED).

Exam

- The exam will consist of a few questions, some of them requiring a detailed answer, while others requiring a few sentence answers, plus four problems similar to (or just from) the home exercises.

- No help from the books, lecture notes, or any other material is allowed during the exam. A standard non-programmable calculator could be used.

Assessment

Your grade will be based on:

- 50% Exam on lecture course
- 30% Homeworks (problems)
- 20% Data reduction / analysis

Introduction

Seeing the Night Sky with Our Naked Eyes

Viewing objects in wavebands other than optical

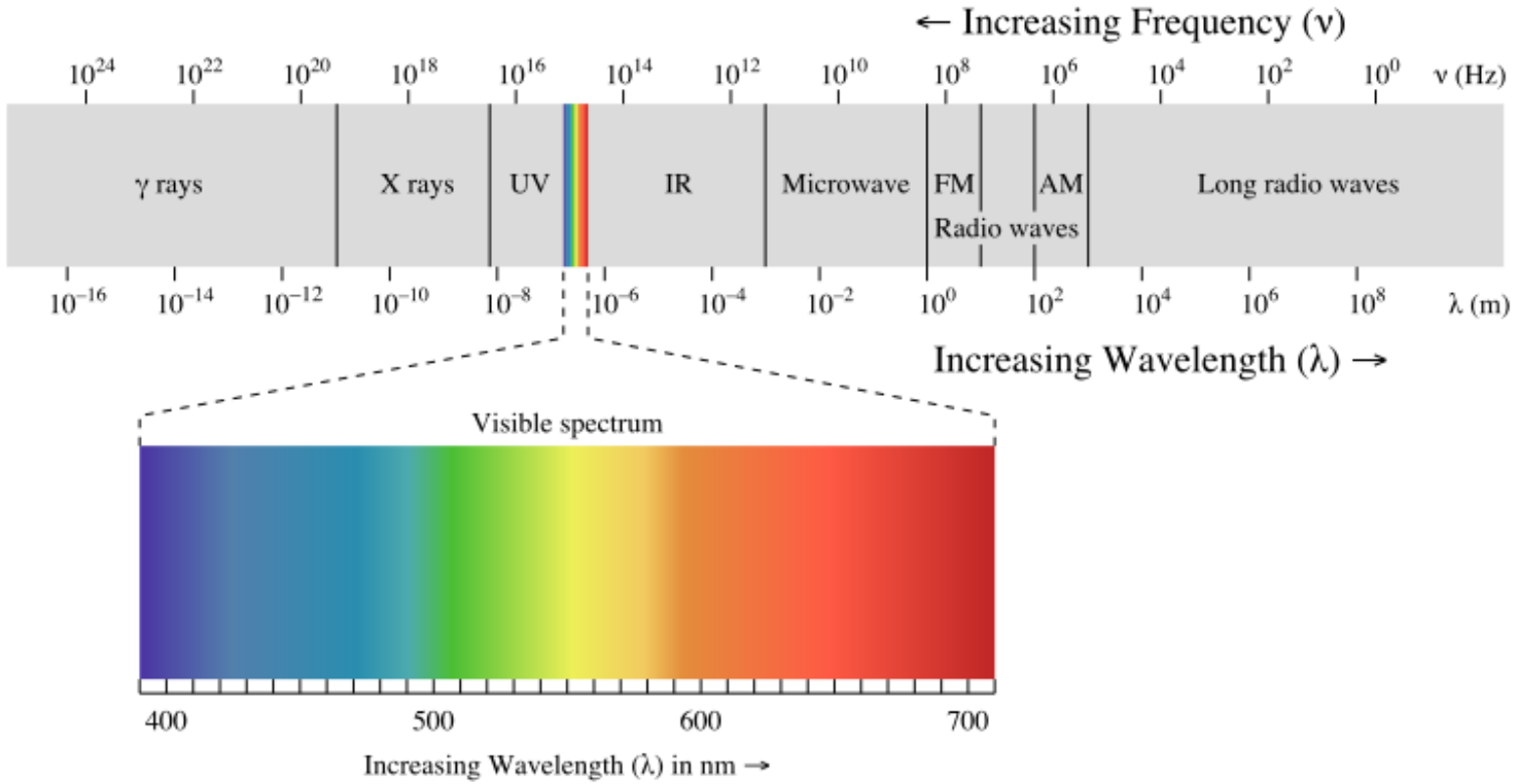
The Effect of the Earth's Atmosphere - I

Continuum Radiation Processes

Eye as a detector

Seeing the Night Sky with our naked Eyes

- With an unaided eye we view our universe through a very small window, from blue light to red light.
- The "big" window is the electromagnetic spectrum. This band, or range of electromagnetic energy exists from extremely long radio waves (kilometres), to the extremely energetic gamma rays.

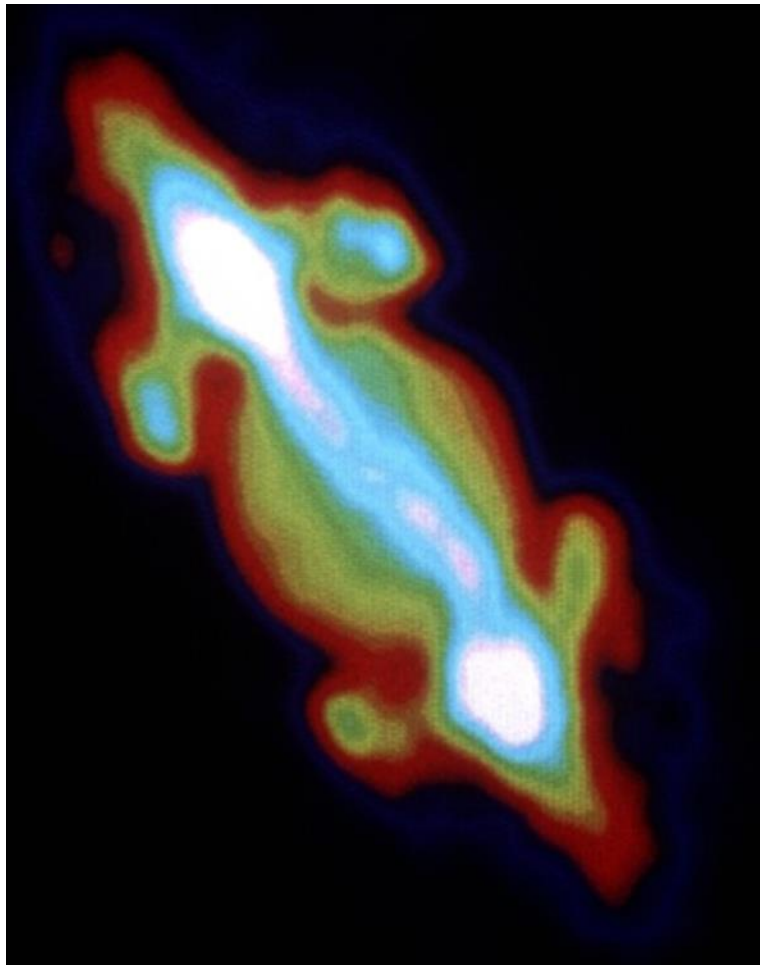


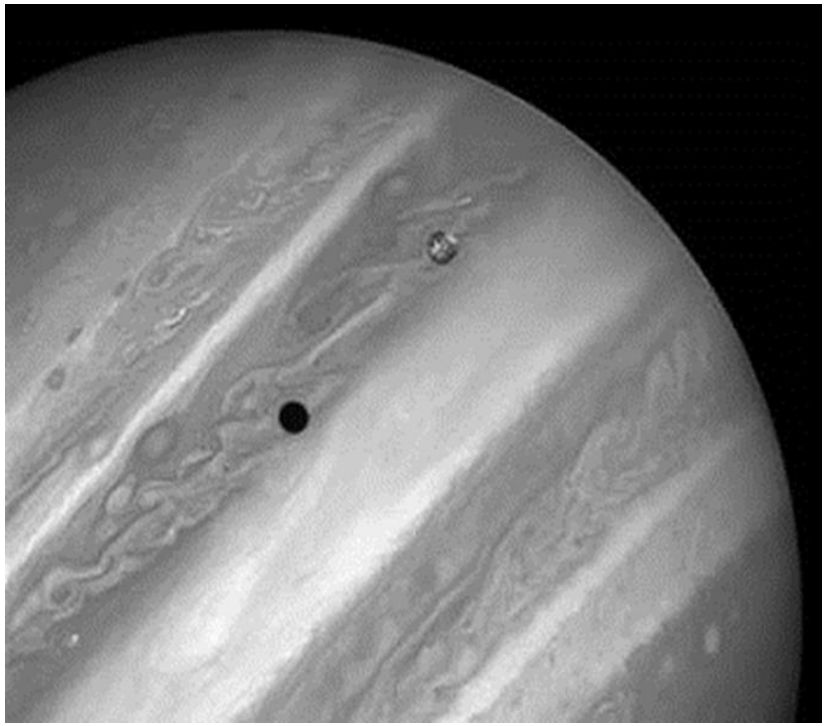
Viewing objects in wavebands other than optical

We have a very limited view into the universe through our visual senses but with the aid of modern technology we can see things that are far beyond that. Being able to see images from radio waves to X-rays and beyond has opened up the universe for our study.



Jupiter

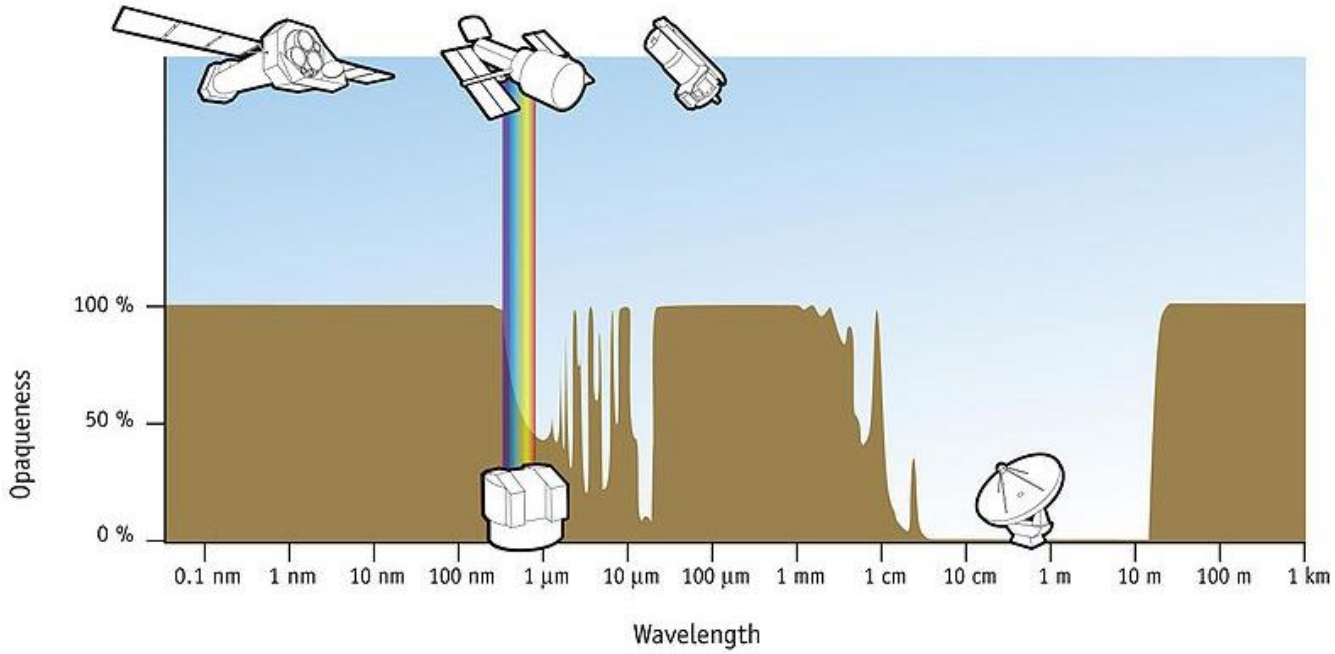




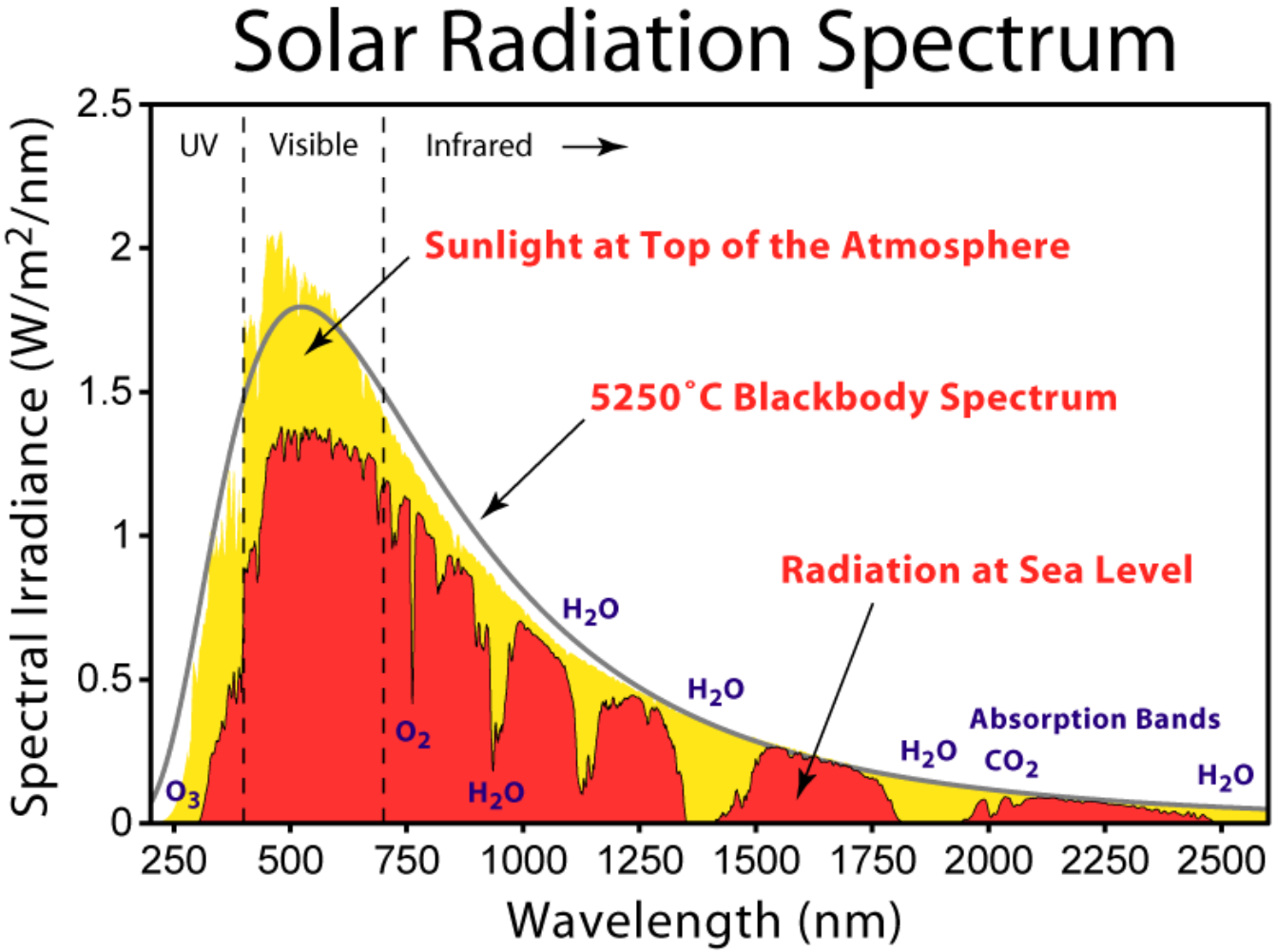
Left is in radio, synchrotron (high magnetic field); right is optical which reflected sunlight.

The Effect of the Earth's Atmosphere - Ia

- Ground based astronomy is heavily influenced by the Earth's atmosphere. In addition to the interruption of observations by clouds, the atmosphere affects the resolution that can be achieved, the accuracy of position and brightness measurements.
- The optical window and the radio window – the only two wavelength ranges of the electromagnetic spectrum that are not severely attenuated by the atmosphere.



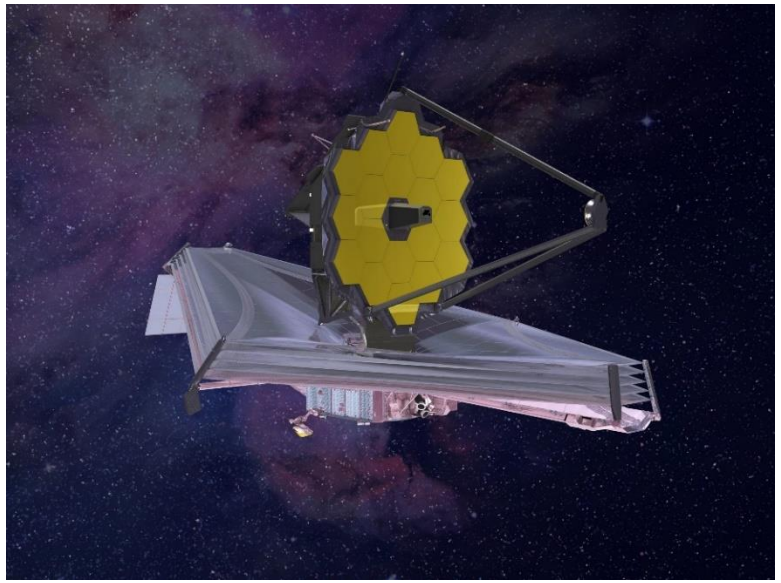
The Effect of the Earth's Atmosphere - Ib



The Effect of the Earth's Atmosphere - Ic

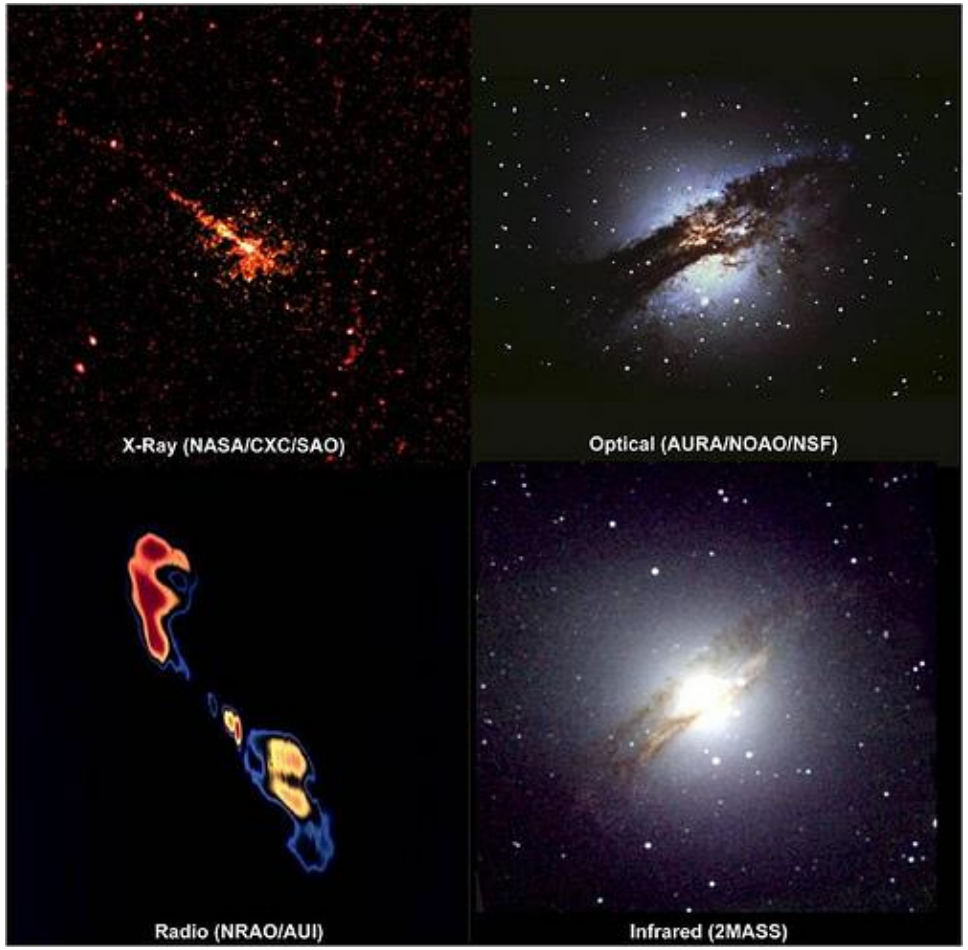
- The Earth's atmosphere is opaque to ultraviolet radiation, X-rays, and gamma rays and is partially opaque to infrared radiation so observations in these portions of the electromagnetic spectrum are best carried out from a location above the atmosphere of our planet.
- Another advantage of space-based telescopes is that, because of their location above the Earth's atmosphere, their images are free from the effects of atmospheric turbulence that distorts ground-based observations.







Viewing objects in wavebands other than optical

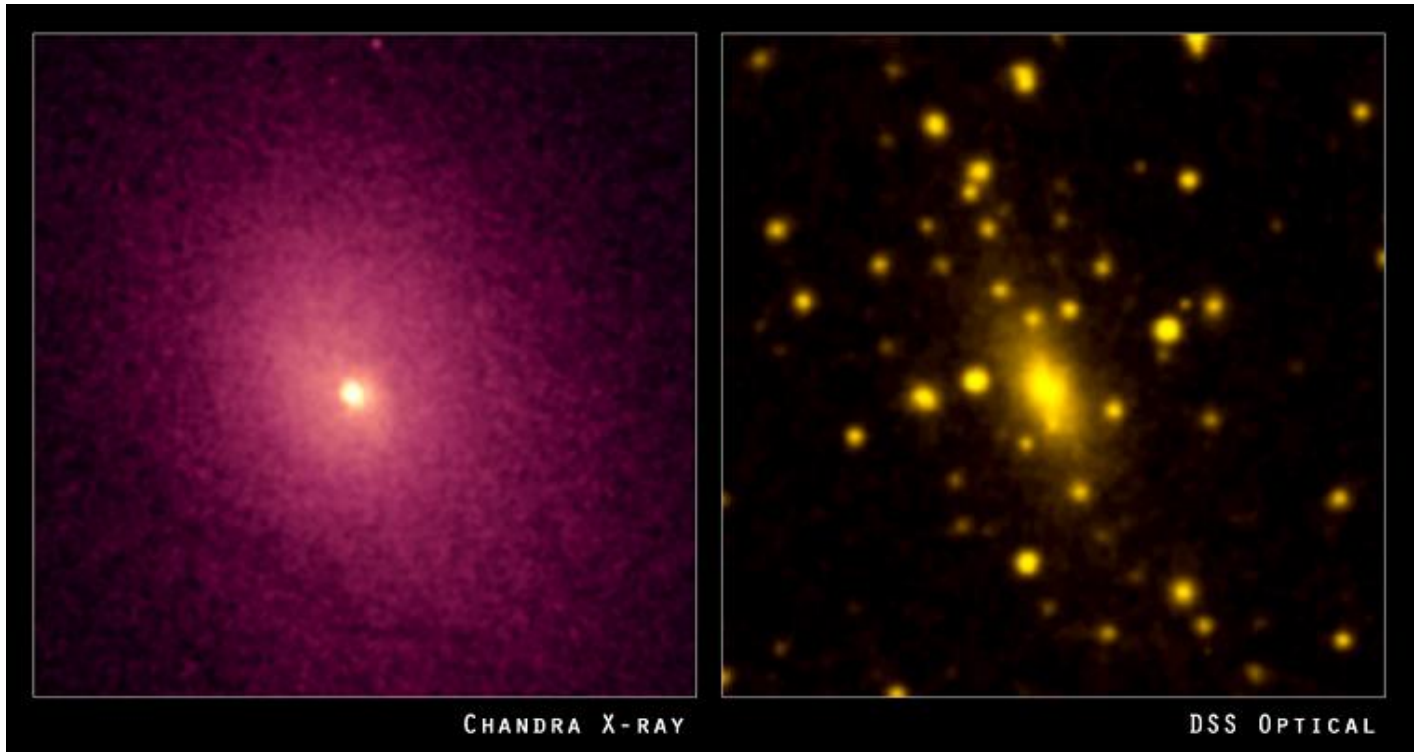


Radio Galaxy Centauri A - one of the most powerful sources of radio waves on the sky.

In **IR** one can very clearly see through the dust hiding the center of the galaxy in **optical** waveband.

X-ray and **radio** pictures show two nice jets of hot gas perpendicular to the galaxy's plain (one has no idea about them just taking the optical/IR/UV image).

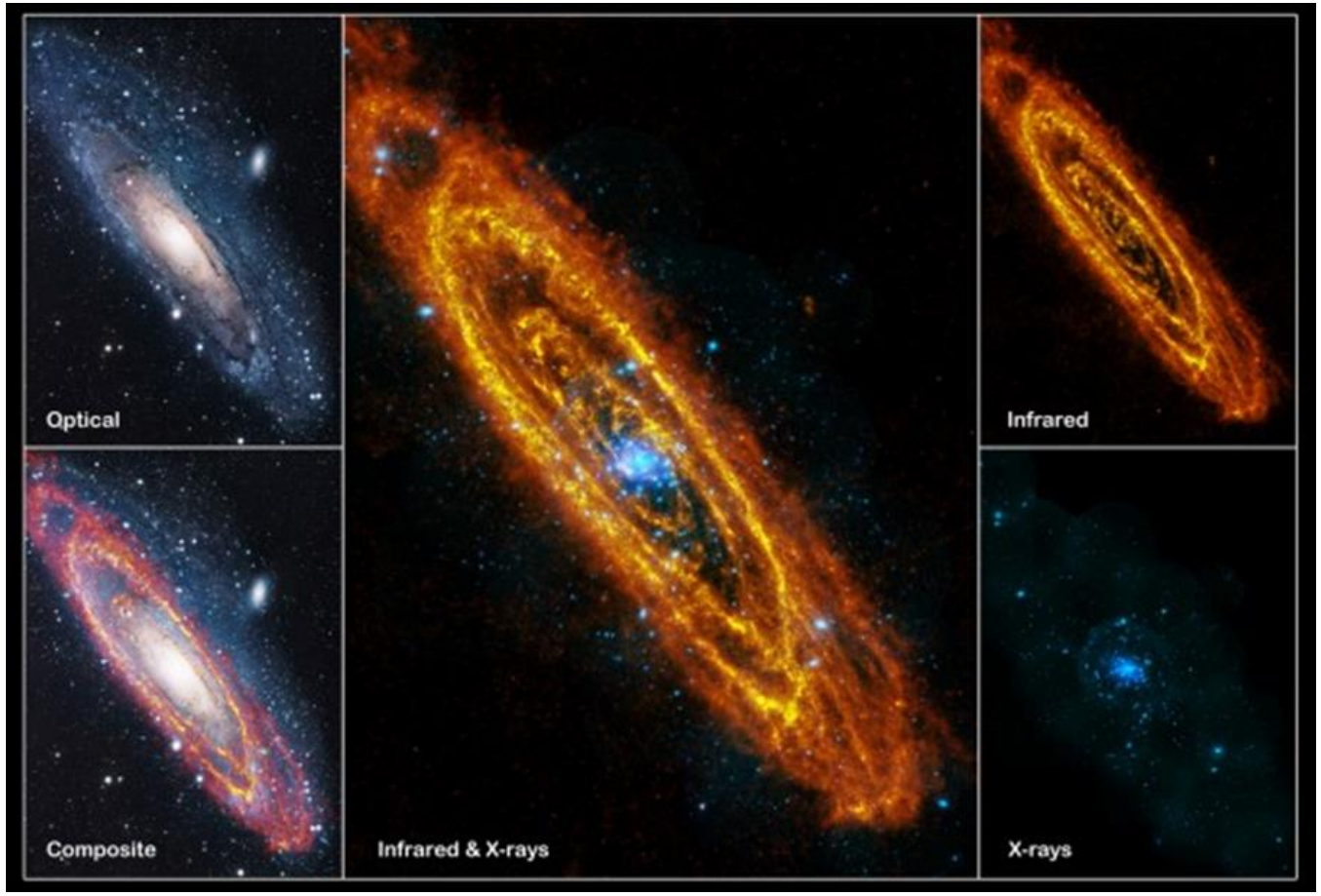
Viewing objects in wavebands other than optical



Galaxy cluster Abell 2029: In the optical waveband one can spot an extremely massive galaxy in the center of the cluster. The gas falling into the gravitational well emits huge amounts of the X-ray radiation due to the accretion .

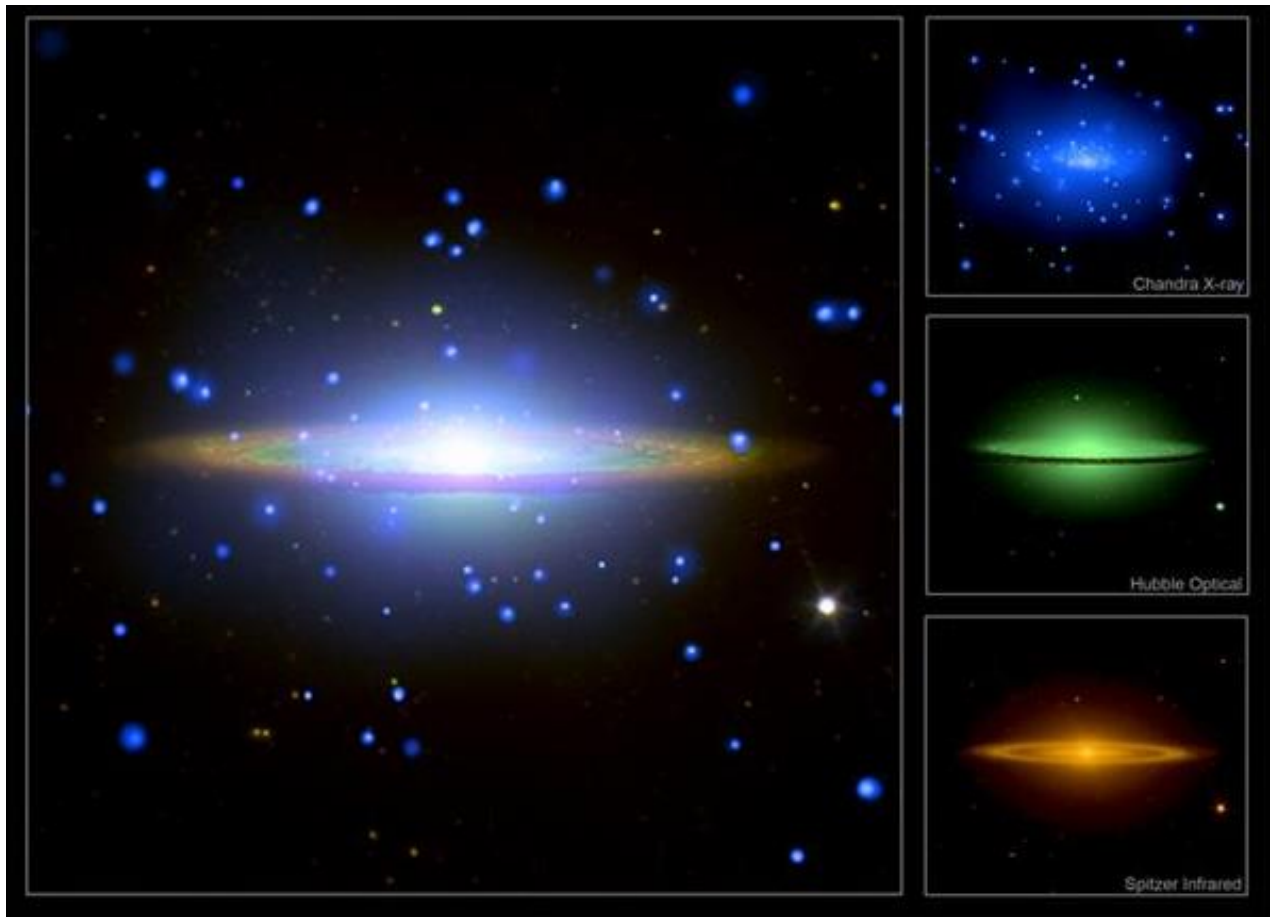
Viewing objects in wavebands other than optical

The Andromeda Galaxy



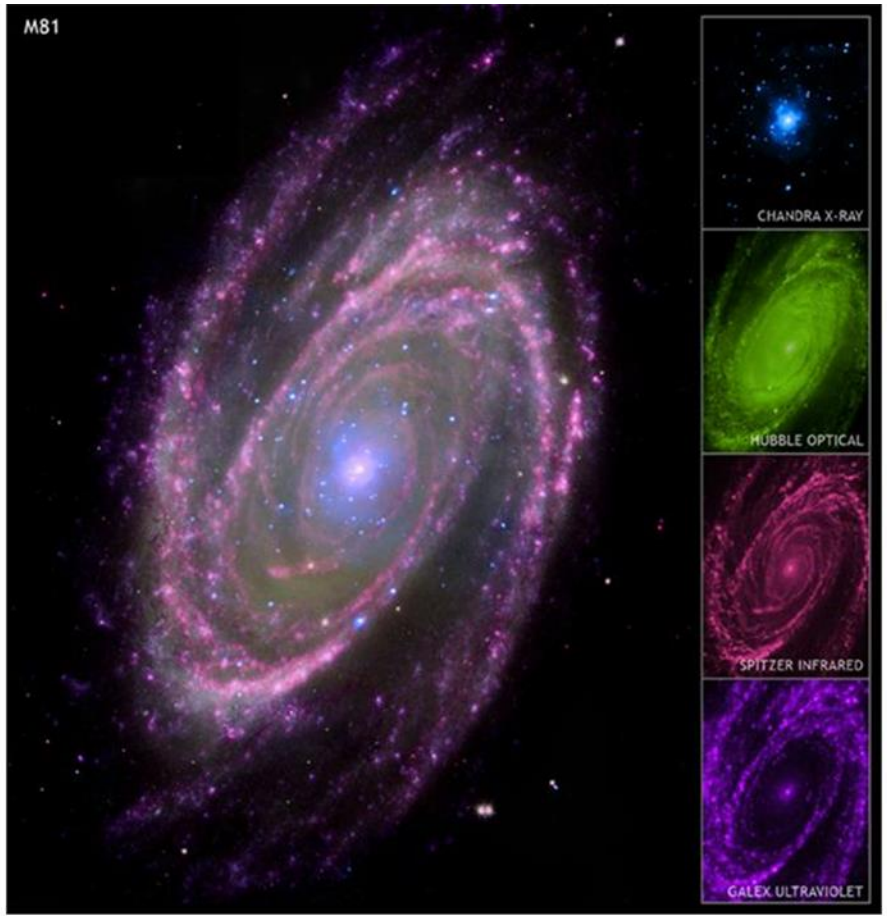
Viewing objects in wavebands other than optical

The Sombrero galaxy (M104)



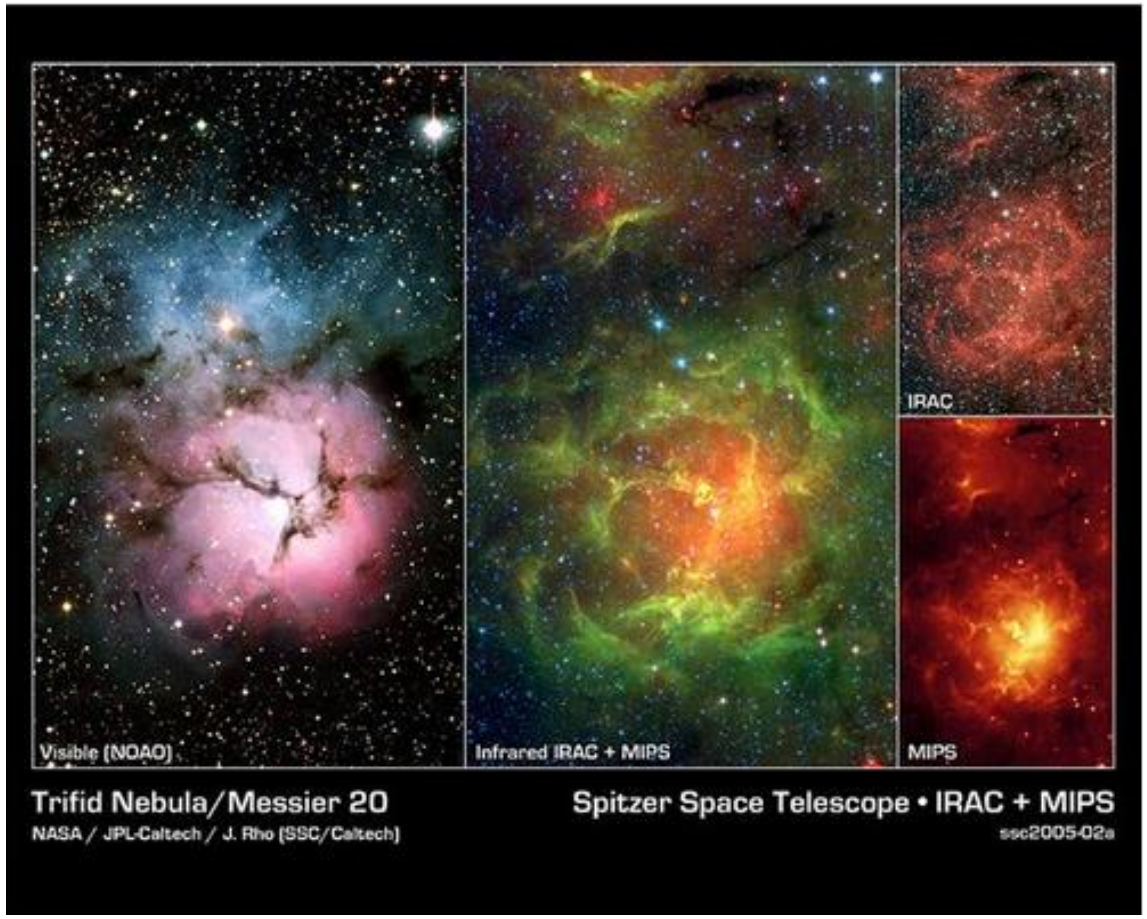
Viewing objects in wavebands other than optical

The spiral galaxy (M81)



Viewing objects in wavebands other than optical

The Trifid Nebula (M20)



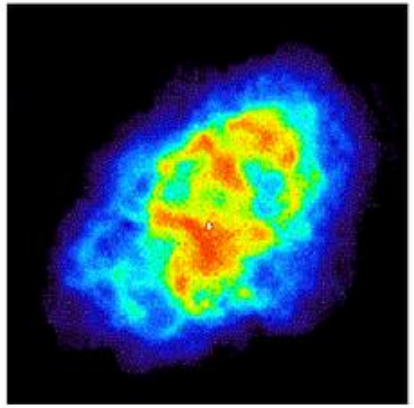
Viewing objects in wavebands other than optical

The Crab Nebula, remnant of SN 1054

X-ray is synchrotron radiation

Optical and Infrared are mostly atomic line radiation

Radio is both synchrotron and thermal bremsstrahlung radiation



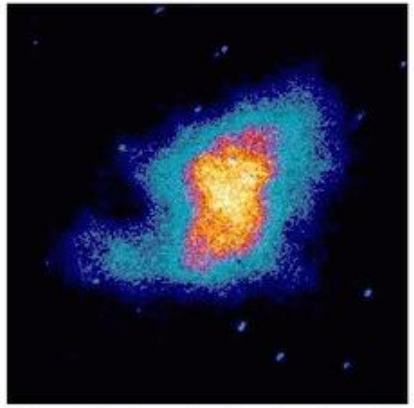
Radio wave (VLA)



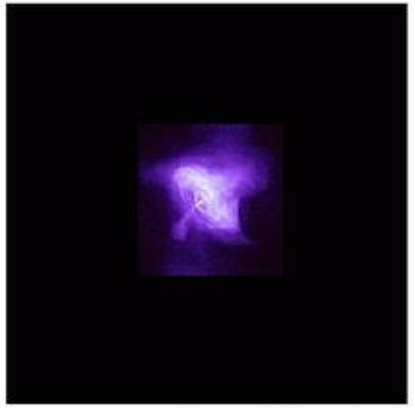
Infrared radiation (Spitzer)



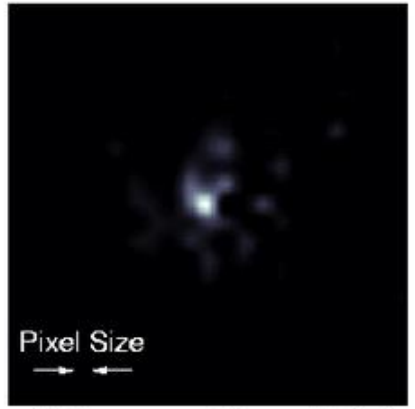
Visible light (Hubble)



Ultraviolet radiation (Astro–1)

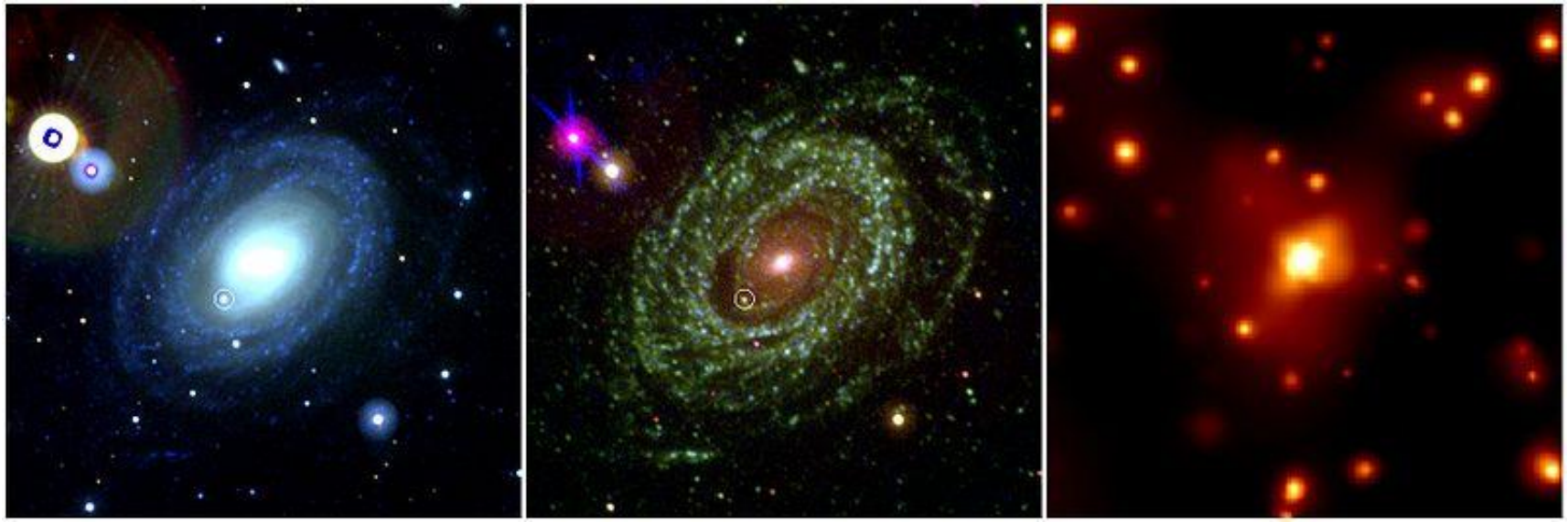


Low–energy X–ray (Chandra)



High–energy X–ray (HEFT)
*** 15 min exposure ***

Viewing objects in wavebands other than optical



Supernova 2005ke is a Type Ia supernova. Shown here is the explosion in optical, ultraviolet and X-ray wavelengths.

What observations in different wavebands are especially good for?

Continuum Radiation Processes

Continuum Radiation Processes

- A photon is a quantum of light, or the smallest possible packet of light at a given wavelength. It is emitted by an atom during a transition from one energy state to another;
- When we think of light as a wave, its wavelength is $\lambda = c_0 / \nu$, where ν is the frequency, and the speed of light is $c_0 \approx 3 \times 10^{10}$ cm/sec;
- In UV-optical astronomy, we **usually** use the Ångström as a unit if wavelength: 1Å = 10^{-8} cm = 0.1 nm. Be familiar with Ångströms!
- When we think of light as particles, the energy of an individual photon is $E = h\nu$, where Planck's constant is $h = 6.626 \times 10^{-27}$ erg s.

Black Body Radiation

- Light can be produced in many ways. The most fundamental source of radiation is a so-called “black body”.
- A black body is an idealized physical body that absorbs all incident electromagnetic radiation. Because of this perfect absorptivity at all wavelengths, a black body is also the best possible emitter of thermal radiation.
- Black body radiation is thus radiation which is in thermal equilibrium and which continuous spectrum depends only on the body's temperature.

Black Body Radiation

Depending on its temperature T, a black body emits radiation according to Planck's law:

$$L_\lambda(T) = \frac{2\,h\,c_0^2}{\lambda^5}\left[\exp\left(\frac{h\,c_0}{\lambda\,k_B T}\right) - 1\right]^{-1}$$

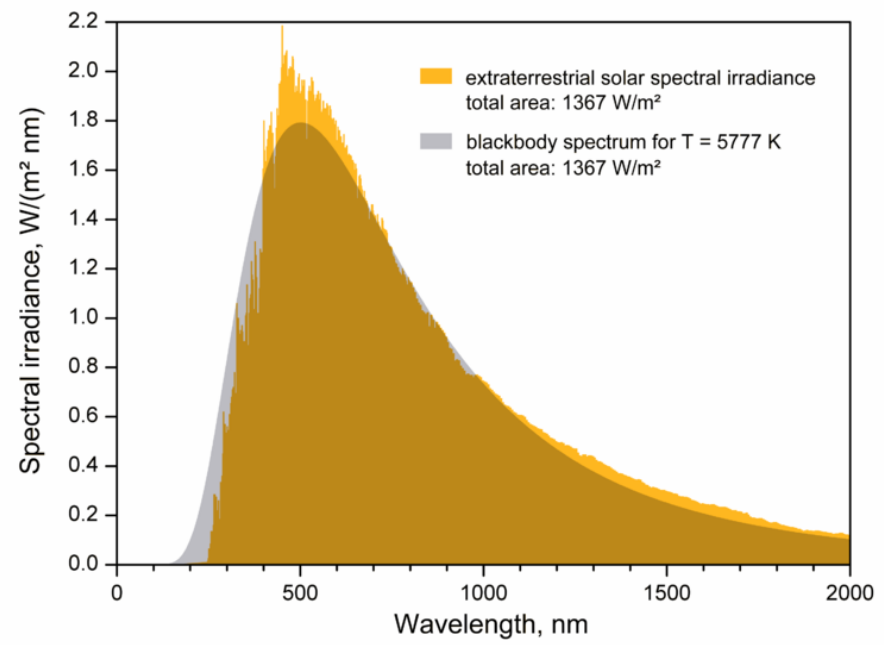
where $k_B = 1.38 \times 10^{-16}$ erg/K is the Boltzmann constant and $L_\lambda(T)$ the spectral radiance at the wavelength λ.

The corresponding formula for $L_\nu(T)$ is

$$L_\nu(T) = \frac{2\,h\,\nu^3}{c_0^2}\left[\exp\left(\frac{h\,\nu}{k_B T}\right) - 1\right]^{-1}$$

Black Body Radiation

- A black body is the extreme case of an optically thick medium. In the Universe, many plasma systems emit radiation approximately according to Planck's law, at least, in certain wavelength ranges.

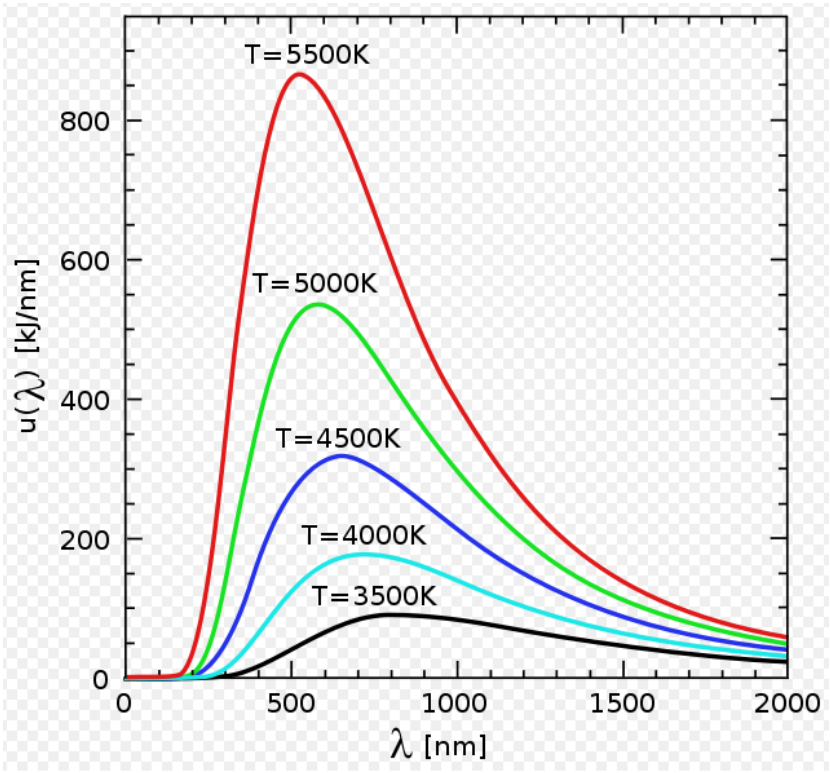


Properties of the blackbody spectrum

- As the temperature increases, the peak of the blackbody radiation curve moves to higher intensities and shorter wavelengths.
- **Wien's displacement law:**
  $\lambda_{max} \approx 0.29/T$ [cm]$= 2.9 \times 10^7/T$ [Å]

  λ_{max} is a function only of the temperature

  For $T=10^7$ K we have $\lambda_{max} \approx 2.9$ Å
  This is a typical X-ray wavelength range



$L(T+\Delta T) > L(T)$ at all λ for positive ΔT

Properties of the blackbody spectrum

- The Rayleigh-Jeans approximation is used at low frequency, particularly in Radio Astronomy.

Raylegh-Jeans regime $h\nu \ll kT$

$$B_\nu \simeq \frac{2\nu^2}{c^2} kT$$

- Wien's law is useful at optical, UV and X-ray wavelengths.

Wien regime $h\nu \gg kT$

$$B_\nu \simeq \frac{2h\nu^3}{c^2} e^{-h\nu/kT}$$

Properties of the blackbody spectrum

The Stefan–Boltzmann law gives the total energy being emitted at all wavelengths by the blackbody:

$$E=\sigma T^4$$

where the Stefan–Boltzmann constant

$$\sigma=5.67\times 10^{-5}\ \text{erg cm}^{-2}\ \text{s}^{-1}\ \text{K}^{-4}$$

Other radiation processes in astrophysics

- **Thermal Bremsstrahlung** (braking) radiation relates to the radiation emitted by free electrons in an ionised plasma due to encounter with ions. The braking radiation results when particles moving in an ionized plasma are accelerated by the Coulomb forces exerted by other charged particles.
- This is one of the most common continuum radiation processes in astrophysics.

Other radiation processes in astrophysics

- **Synchrotron emission** arises from the motion of relativistic electrons in a magnetic field.
- This radiation is the most important source of continuum radiation at radio wavelengths, and is seen at infra-red, optical, ultra-violet and even X-ray wavelengths.
- The synchrotron jets in radio galaxies and quasars like M87 and 3C273 can be seen at all wavelengths from low frequency radio to X-ray.

Other radiation processes in astrophysics

- **Compton scattering** is a process where energy is transferred from a photon to a free electron during an interaction between the two.
- For low photon energies ($h\nu << m_e c^2$) this process reduces to the classical case of Thomson scattering.
- Of more interest in astrophysics is the process where the electron is very energetic, and energy is transferred to the photon. This is known as **inverse Compton scattering radiation**.

Multi-Wavelength Observations

- Different processes dominate at different wavelengths
- To establish the physics of a source you often need to observe it at a variety of wavelengths

Multi-Wavelength Observations

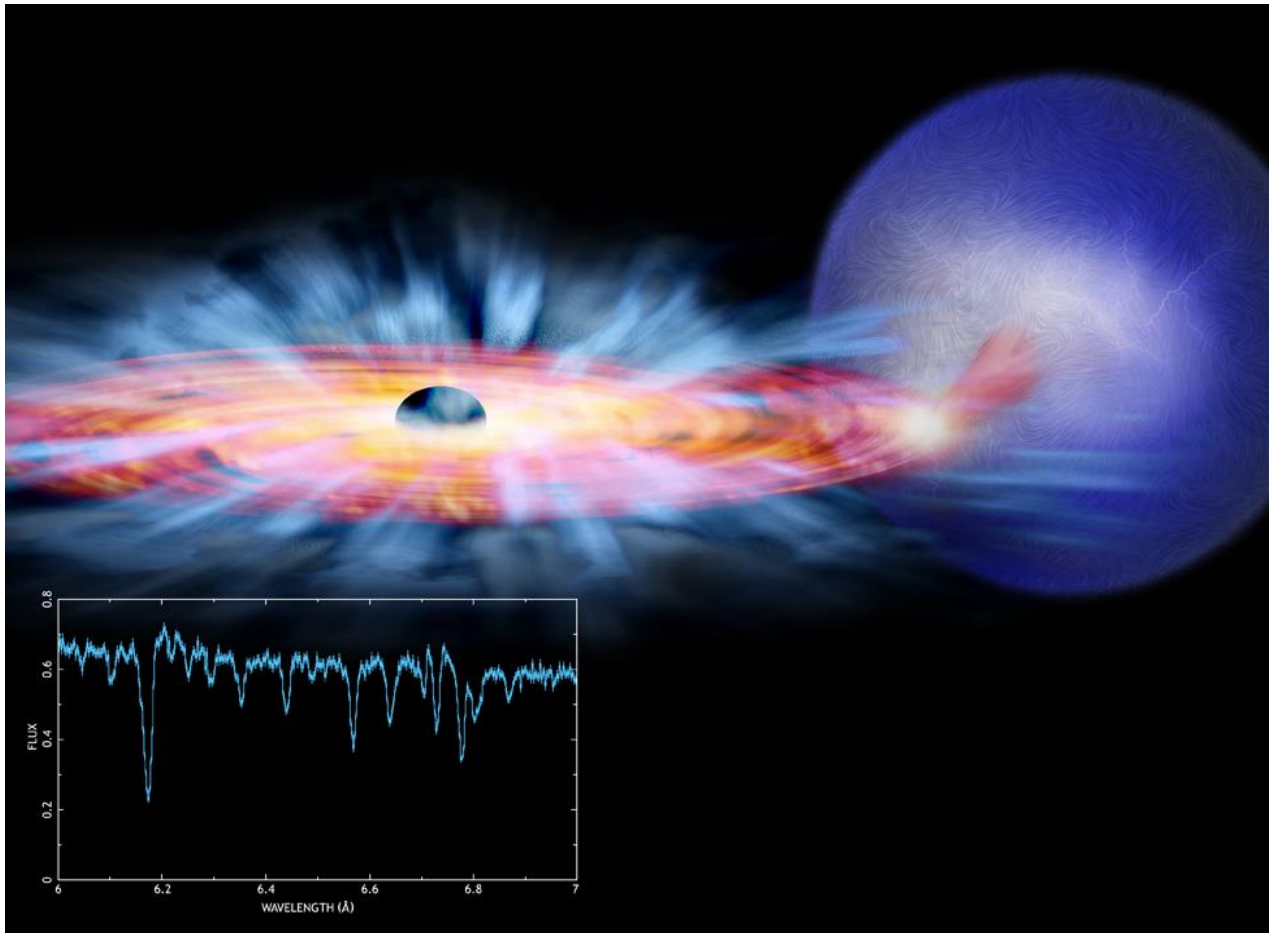
X-ray binary:

X-ray and gamma-ray UV

Optical

Infra-Red

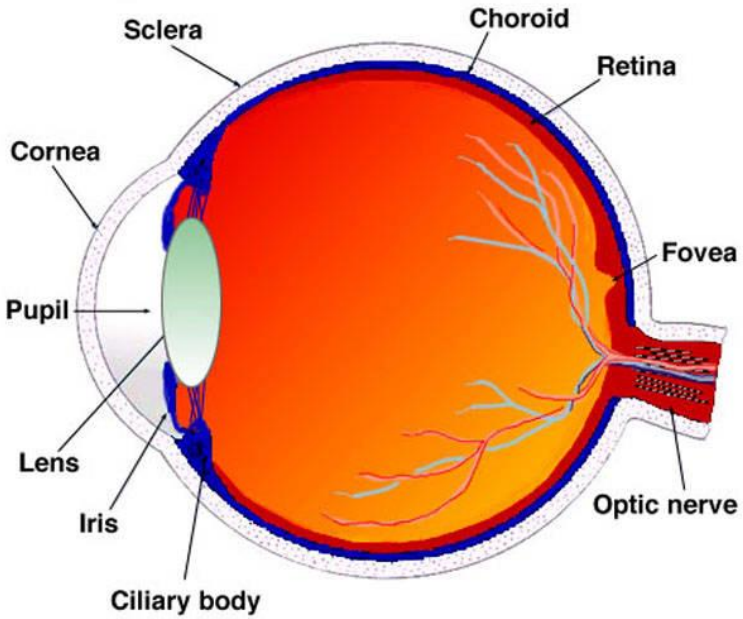
Radio



Detectors

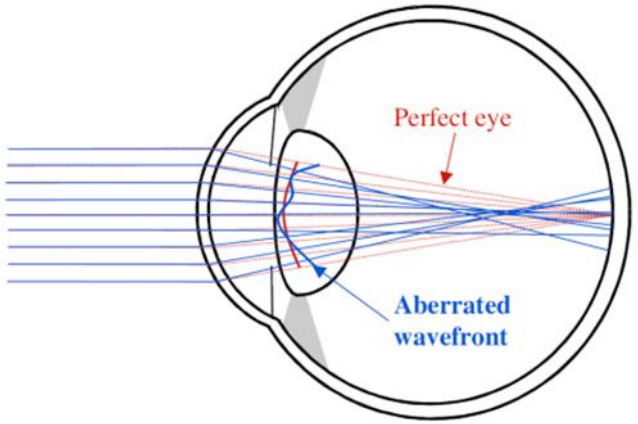
The central theme of this course is the detection and characterization of photons with both ground-based instruments and instruments aboard spacecraft.

The human eye as a natural detector

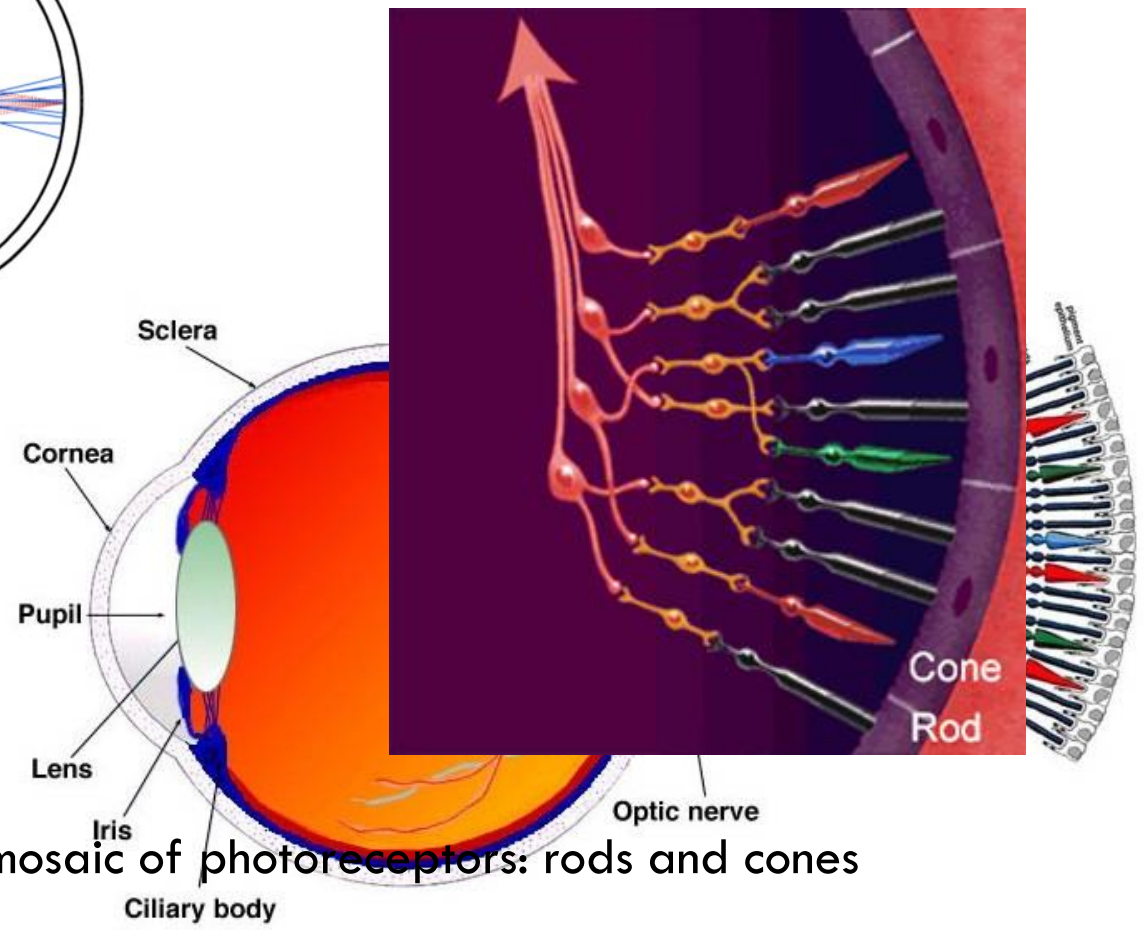


- The most fundamental of the detectors
- Rarely used for primary detection but there are still a few applications in which it performs even better than other detection systems
- The eye and brain act together in the visual process. A simple consideration of the eye on basic optical principles will be misleading

The human eye as a natural detector



A lens (as a photo-camera objective) produces an image on the retina.



The retina is covered with a mosaic of photoreceptors: rods and cones

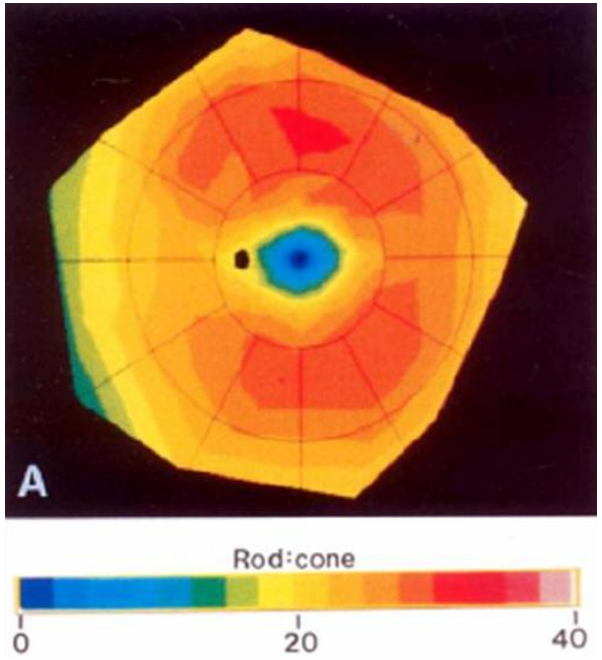
Photoreceptor mosaics

Self reading material

There are two different types of photoreceptors distributed unevenly across the retina:

- **Rods** (approximately 100,000,000) - very sensitive to low levels of light. In bright light conditions the response of the rods is saturated;
- **Cones** (approximately 5,000,000) – faster but less sensitive photoreceptors.

The rod/cone ratio is lowest in the foveal region and higher in the periphery (peripheral vision!).



Photoreceptor mosaics

Self reading material

- Trade off: Sensitivity to light versus spatial resolution.
- Two parallel systems:
  - One that favour sensitivity to light (Rods)
  - One that favour resolution (Cons)

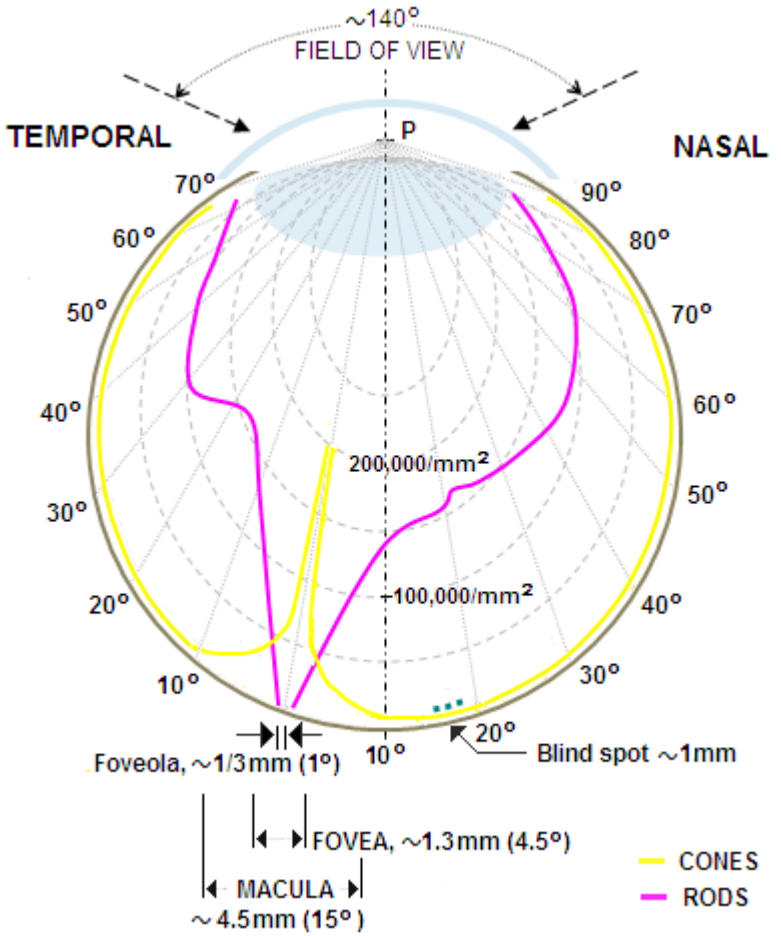
- **Rods:** low resolution image of the world that persists even in low illumination condition.
- **Cones:** high resolution image of the world in good illumination.

Photoreceptor mosaics

Self reading material

Fovea is area of highest concentration of photoreceptors

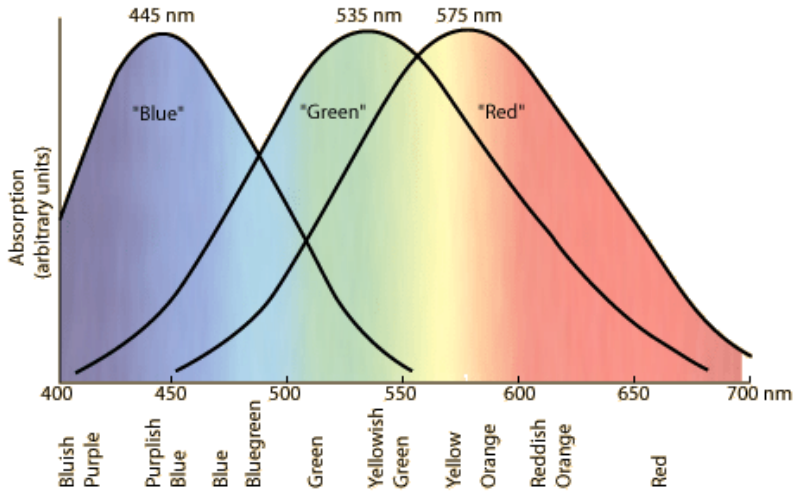
- fovea contains no rods, just cones
- approximately 50,000 cones in the fovea
- cannot see dim light sources (like stars) when we look straight at them!



Cones and Colours

Self reading material

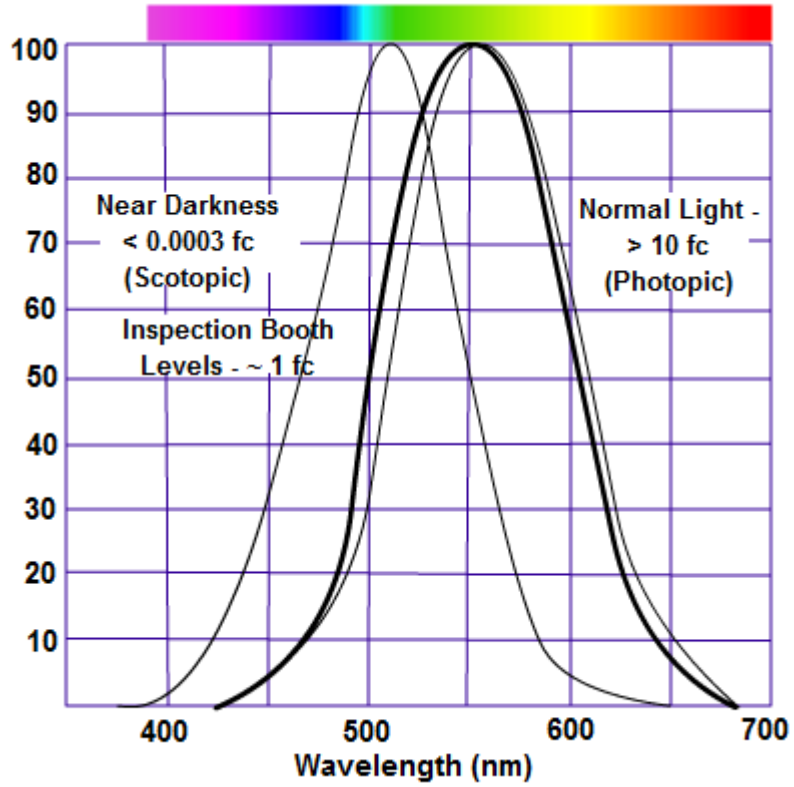
There are three types of cones, each one of them responding best to different wavelengths (short - blue, middle - green, and long - red). Their combined responses generate colour vision.



By population, about 64% of the cones are red-sensitive, about 32% green sensitive, and about 2% are blue sensitive. The "blue" cones have the highest sensitivity (~2 times) and are mostly found outside the fovea.

The Human Eye's Response to Light

Self reading material



The sensitivity of the cones is only about 1% of the maximum for the rods.

Let us now make a brief excursion away from our main topic (Astronomy), but it is very important.

Color blindness

Color blindness or color vision deficiency (CVD) is the decreased ability to see color or differences in color.

The most common cause of color blindness is a variation in the functionality of one or more of the three classes of cone cells in the retina.

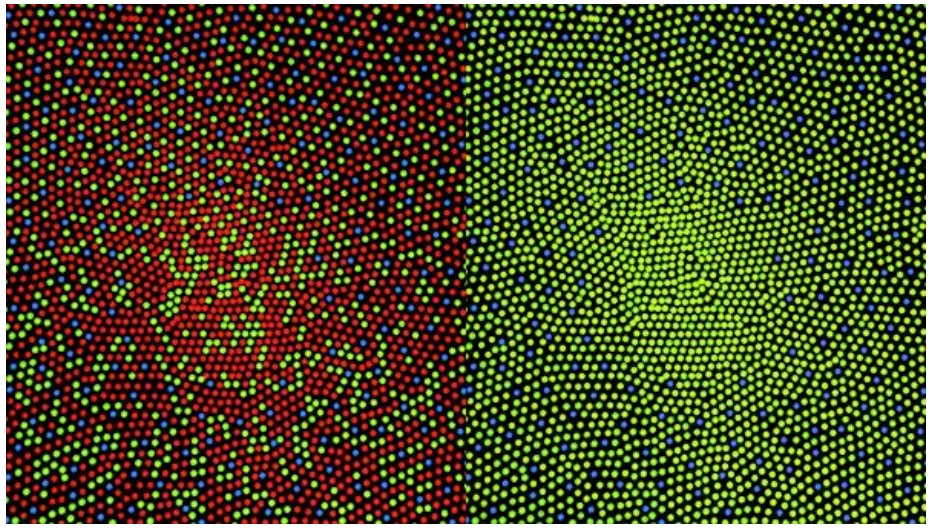
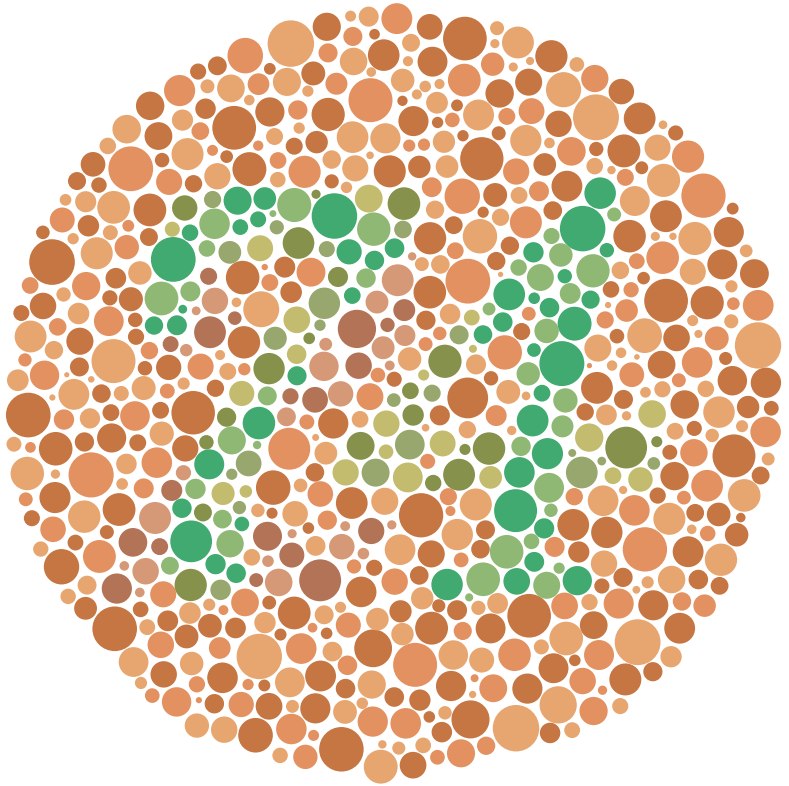


Illustration of the distribution of cone cells in the fovea of an individual with normal color vision (left), and a color blind (protanopic) retina. The center of the fovea holds very few blue-sensitive cones.

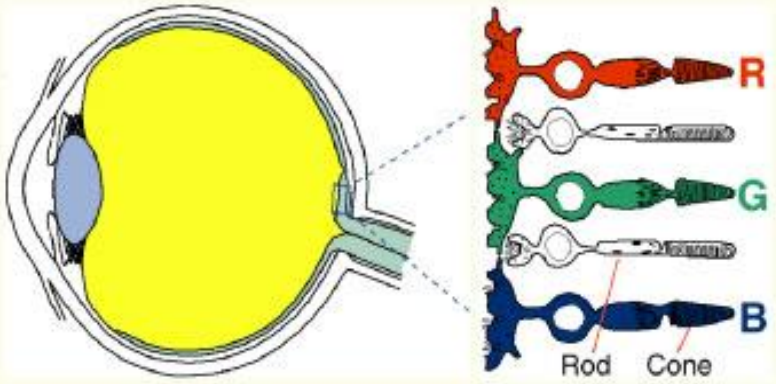


What do you see here?

How common is color blindness?

One in twelve males (8%) and *one in 200* females (0.5%) are red-green color blind.

(Asian: 5%, French and Scandinavian : >10%)



Type 1 (protanopes) and Type 2 (deuteranopes): functional defects in red and green cone cells, respectively.

People with defects in blue cone cells (type 3: tritanopes) are relatively rare (*one in tens of thousands.*)

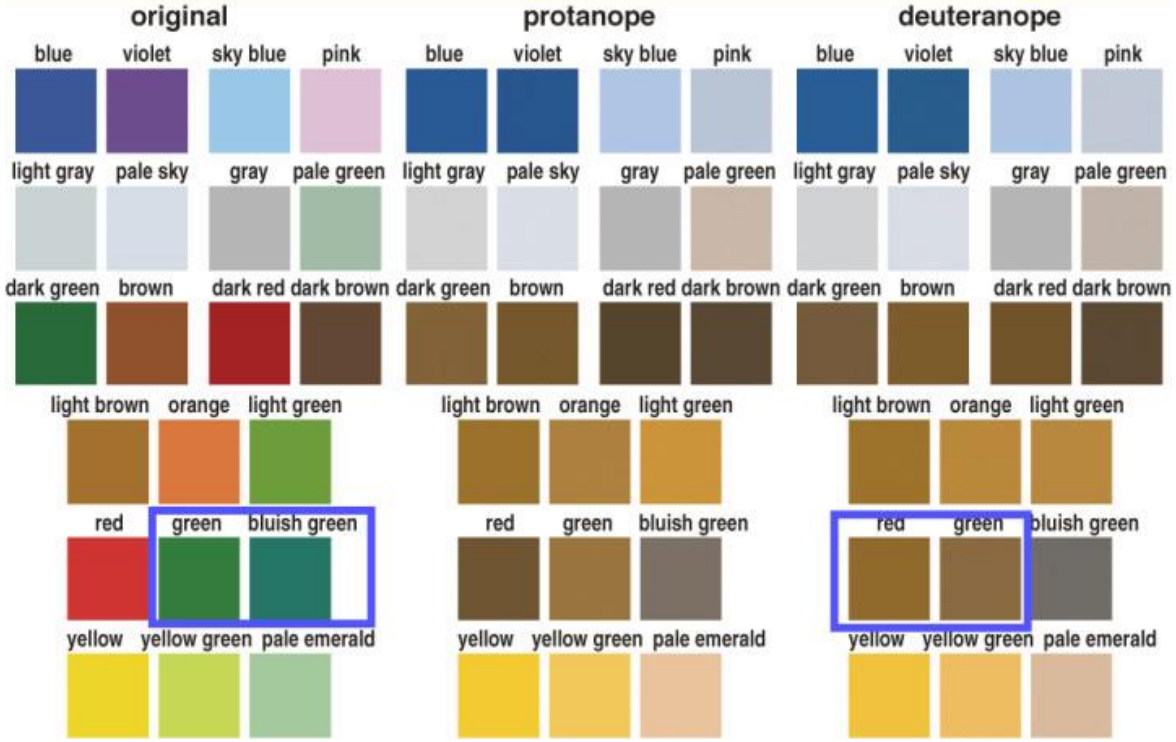
Red-green color blindness is *commoner than AB blood group*. There should be around 10 color blind people in the room with 250 people !

Color blindness

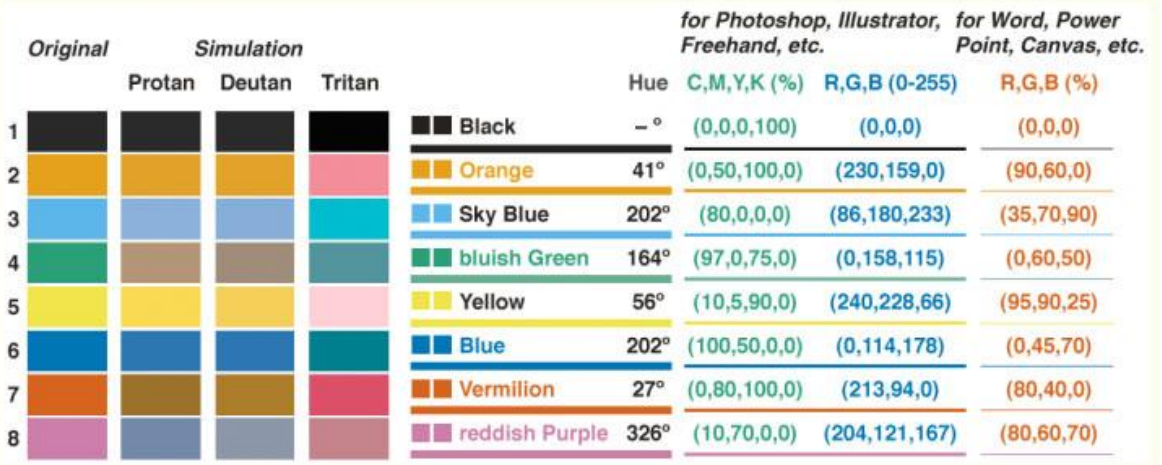
may make some academic activities more **difficult**.

Colors difficult to distinguish and identify

The border of color-name categories is not the same among people.



Example of colors that are easier to identify



1. **Red:** Avoid pure red (RGB=100,0,0%). *Use vermilion* (RGB=80,40,0%) or *change to orange* (RGB=90,60,0%).
2. **Green:** Avoid pure green (RGB=0,50,0%), which is confusing with red or brown. *Use bluish green* (RGB=0,60,50%)
3. **Light green** (RGB=0,100,0%) and **yellow** (RGB=100,100,0%) will appear the same to color blinds.
   Avoid using colors between yellow and green.

There are always color blind people among the audience, readers and referees. Please take this into account when preparing your presentations. **Read a presentation about it on a course web-page!**

The Human Eye (tech. specs)

- **Quantum Efficiency (QE).** A critical performance measure of any photon detector is its quantum efficiency as a function of wavelength, $\eta(\lambda)$:

$$\eta = C/N,$$

where C is the number of counted events and N the number of incoming photons.

$$\eta_{Eye} \approx 1\%$$

The Human Eye (tech. specs)

- **Detector Linearity** is one of the most important specification for a detector that is to be used for quantitative analysis. It is defined as the range over which the detector response is linearly related to the stimulus.
- **The Weber–Fechner law** (the case of vision):
  The eye senses brightness approximately logarithmically over a fairly broad range. Hence stellar magnitude is measured on a logarithmic scale.

The Human Eye (tech. specs)

- **Angular Resolution** (spatial resolution) – describes the ability of any image-forming device to distinguish small details of an object. Angular resolution of the eye – 1-10 arcminutes.
- **The response time** is the time a device takes to react to a given input. The response time of the eye – <0.1 sec under well-lit condition and 1 sec under low light condition.



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 2

Vitaly Neustroev

Telescopes

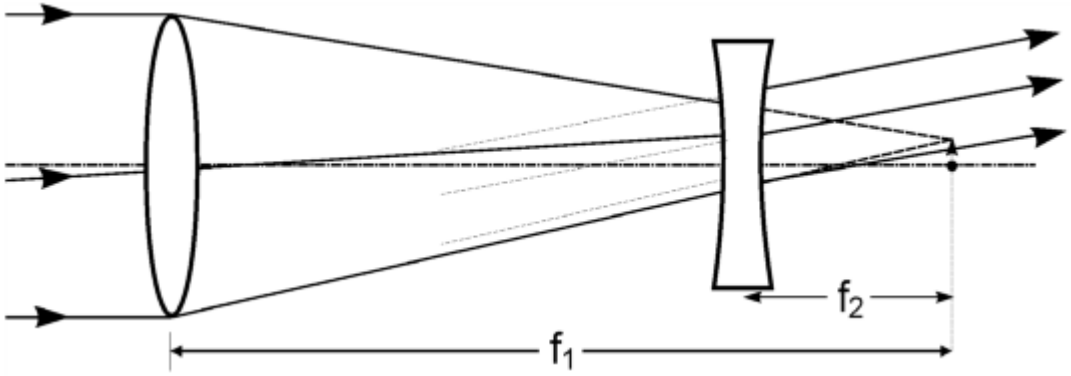
- See how radiation is brought to a focal plane
- Basic technique is reflection off curved surfaces
- Different techniques and materials at different frequency or wavelength.

Telescopes

- Historical introduction
- General principles
- Diffraction and Spatial resolution
- Aberrations in reflecting telescopes
- Optical configuration
- Telescope configurations
- Optical Telescopes
- Segmented mirror telescopes
- Adaptive optics
- New large telescopes
- Problems of Extremely large telescopes
- Radio Telescopes
- Atmospheric transparency
- Infra-Red Telescopes
- UV Telescopes
- X-ray telescopes
- Gamma-ray telescopes

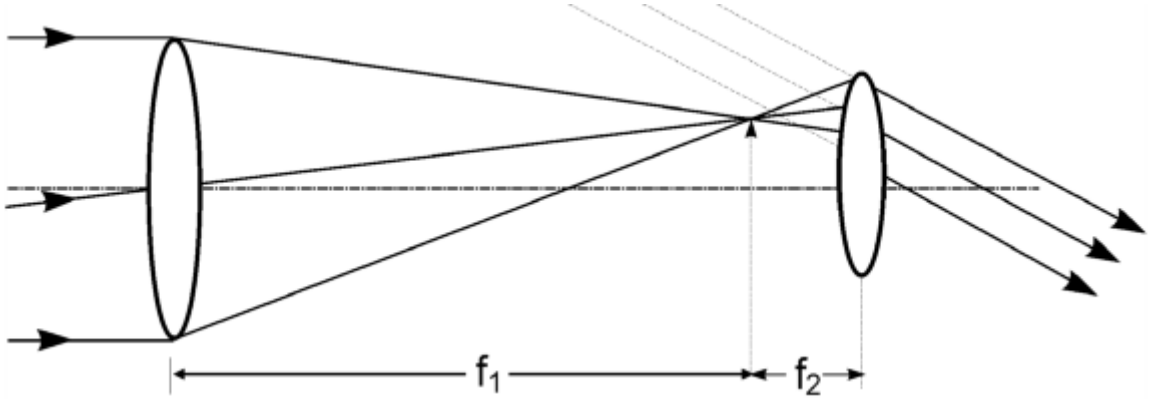
Historical introduction (1)

In July of 1609, Galileo Galilei developed the first astronomical optical telescope. His telescope is formed by the combination of a concave and a convex lens, today known as the Galileo telescope system. The front convex lens is called an objective because it is close to the object being viewed and the rear concave lens is called an eyepiece because it is closest to the observer's eye.



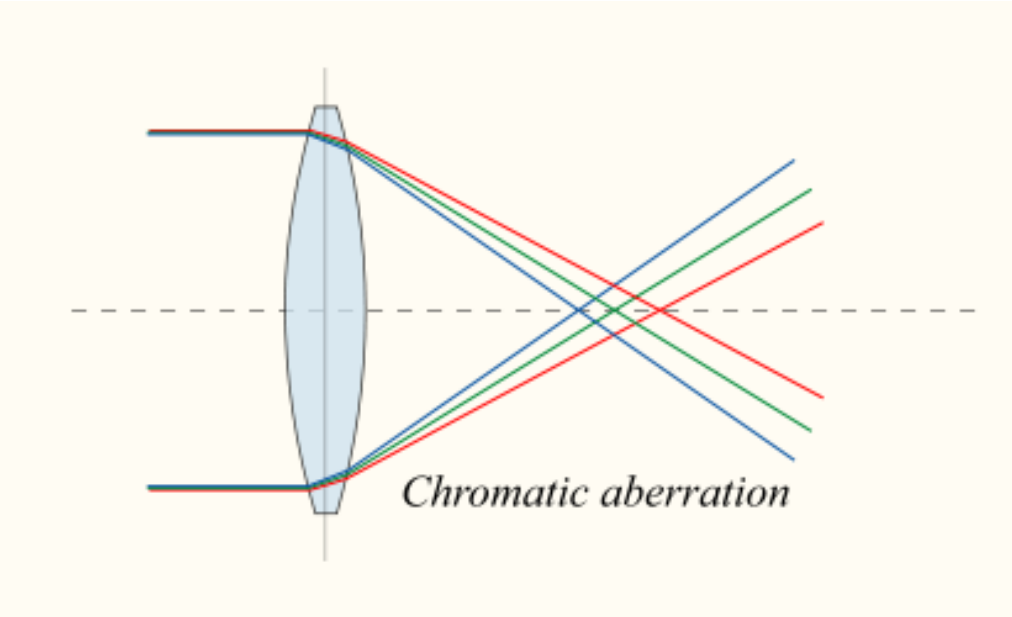
Historical introduction (2)

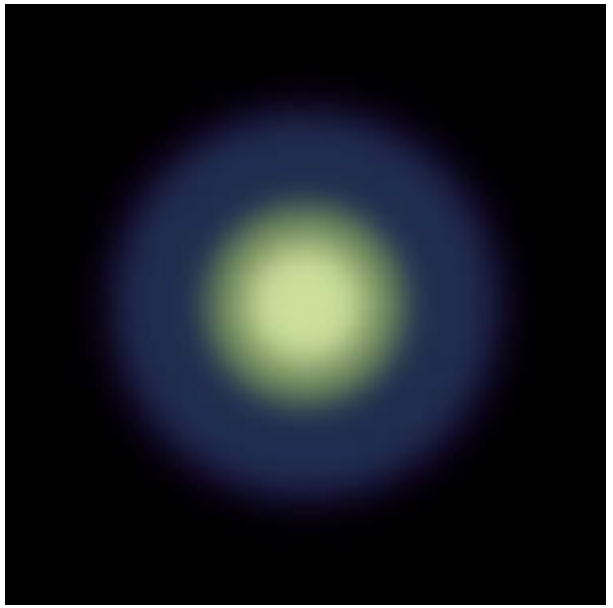
In 1611, Johannes Kepler invented another type of longer telescope comprising two convex lenses, known as the Kepler telescope system. The Kepler telescope forms an upside down image, but with a slightly larger field of view.



Kepler's telescope and Chromatic Aberration (1)

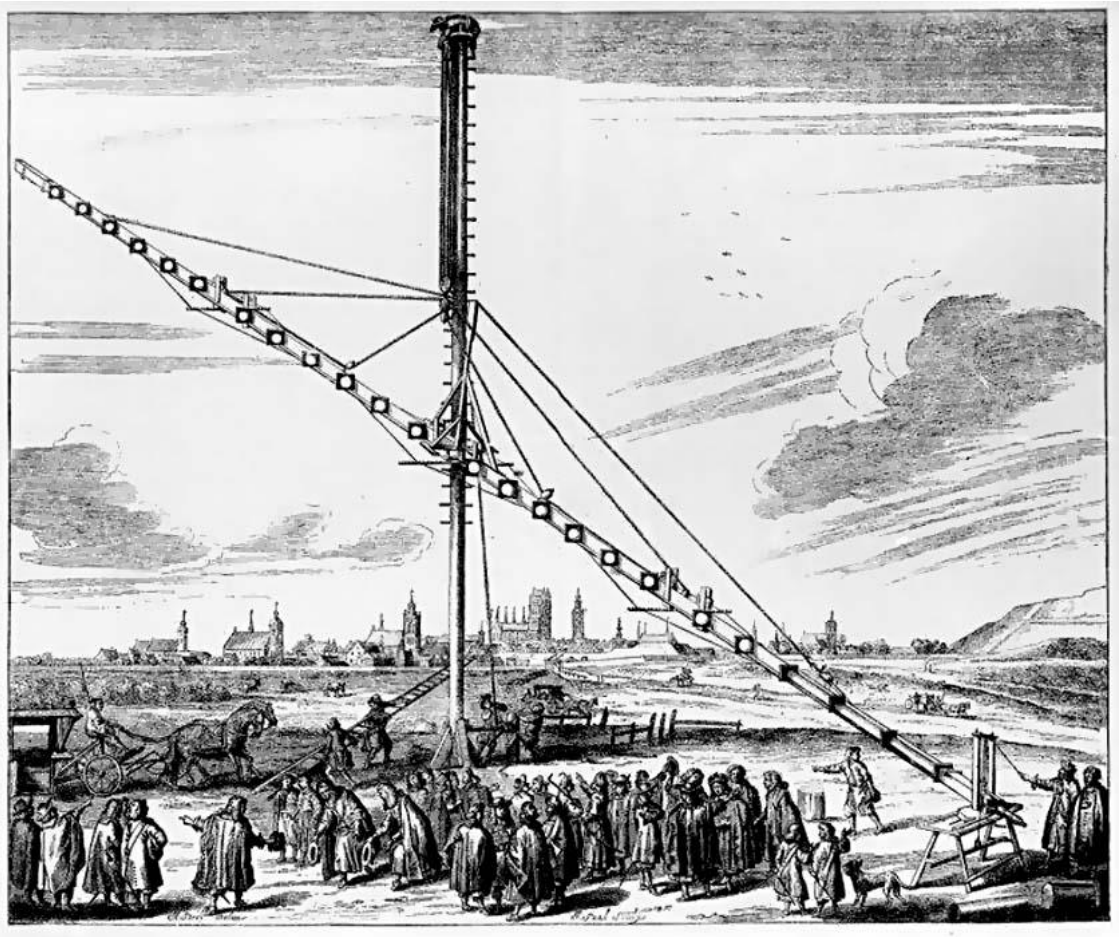
The sharpness of the image in Kepler's telescope is limited by the chromatic aberration introduced by the non-uniform refractive properties of the objective lens



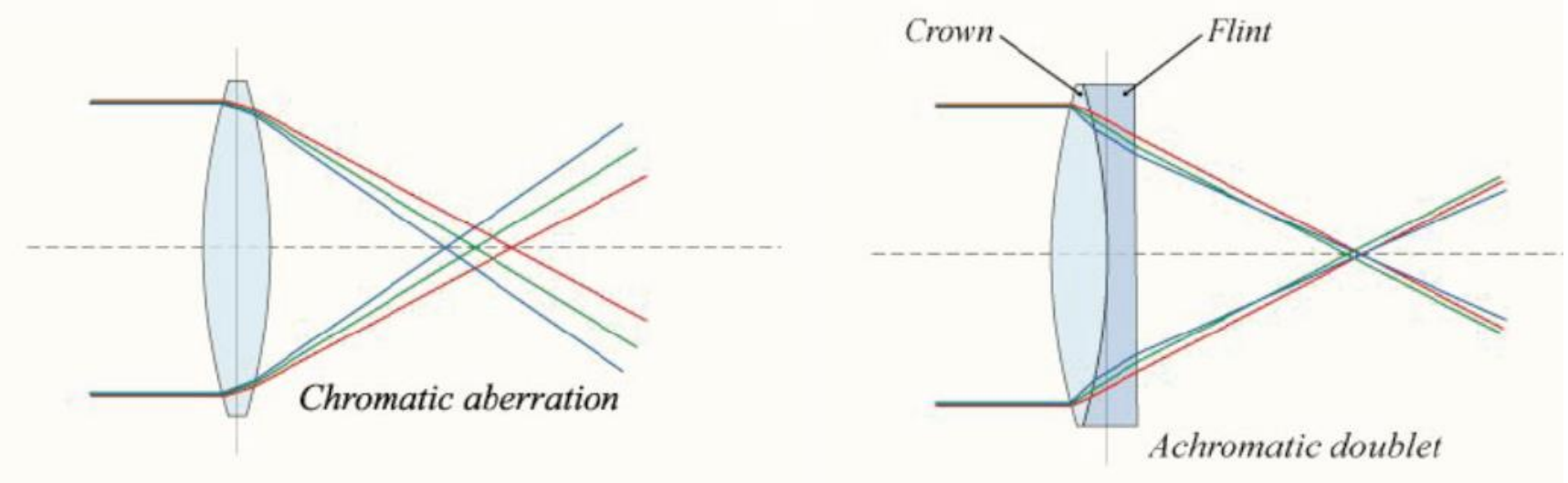


Kepler's telescope and Chromatic Aberration (2)

The only way to overcome this limitation was to create objectives with very long focal lengths.

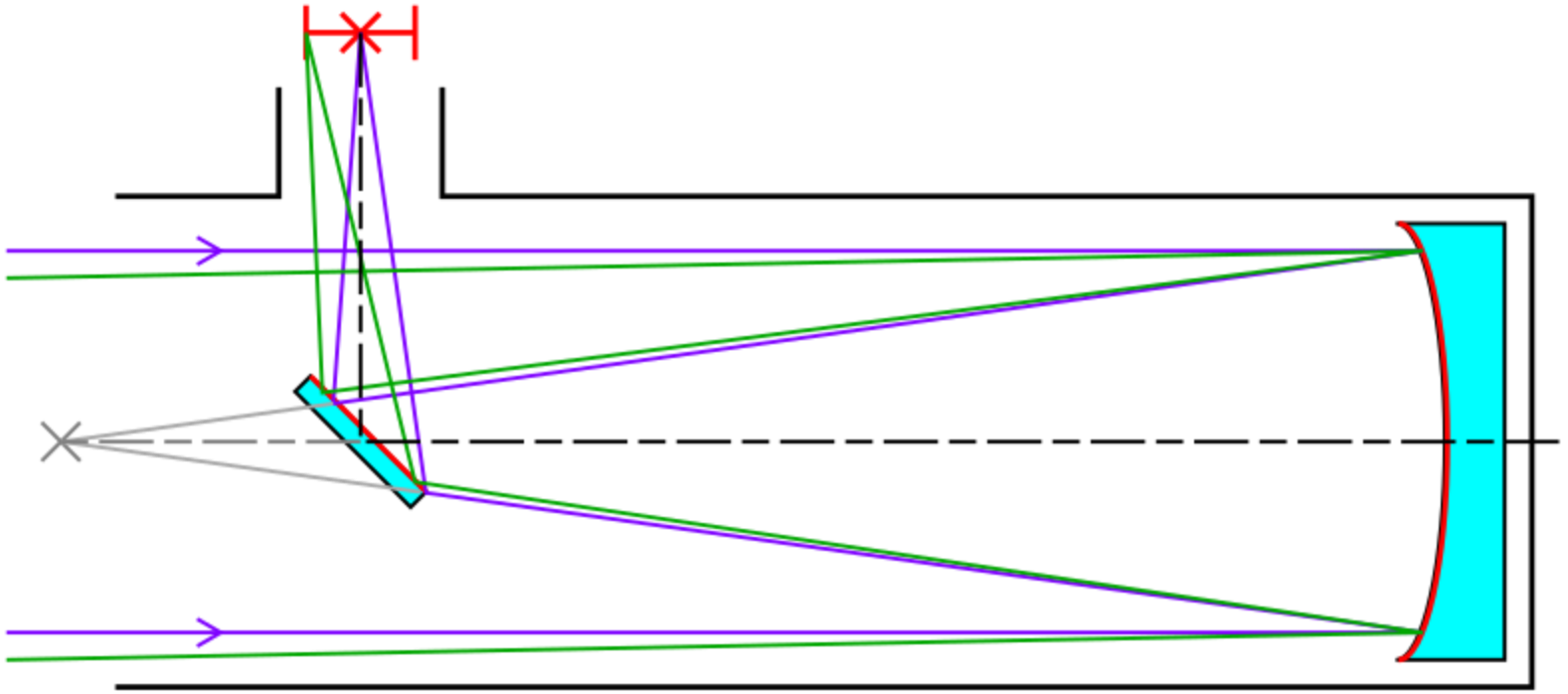


Chromatic Aberration: possible solution



Historical introduction (3)

In 1668, Sir Isaac Newton invented and built the first reflecting telescope. The telescope uses a concave primary mirror and a flat diagonal secondary mirror.



Advantages of the Newtonian design

- They are free of chromatic aberration found in refracting telescopes.
- Newtonian telescopes are usually less expensive for any given aperture than comparable quality telescopes of other types.
- Since there is only one surface that needs to be ground and polished into a complex shape, overall fabrication is far simpler than other telescope designs.
- A short focal ratio can be more easily obtained, leading to wider field of view.
- The eyepiece is located at the top end of the telescope. Combined with short f-ratios this can allow for a much more compact mounting system, reducing cost and adding to portability.

Simple reflecting telescopes

At wavelengths from Radio to UltraViolet, most telescopes are **simple dishes** made of

wire mesh (low frequency radio);

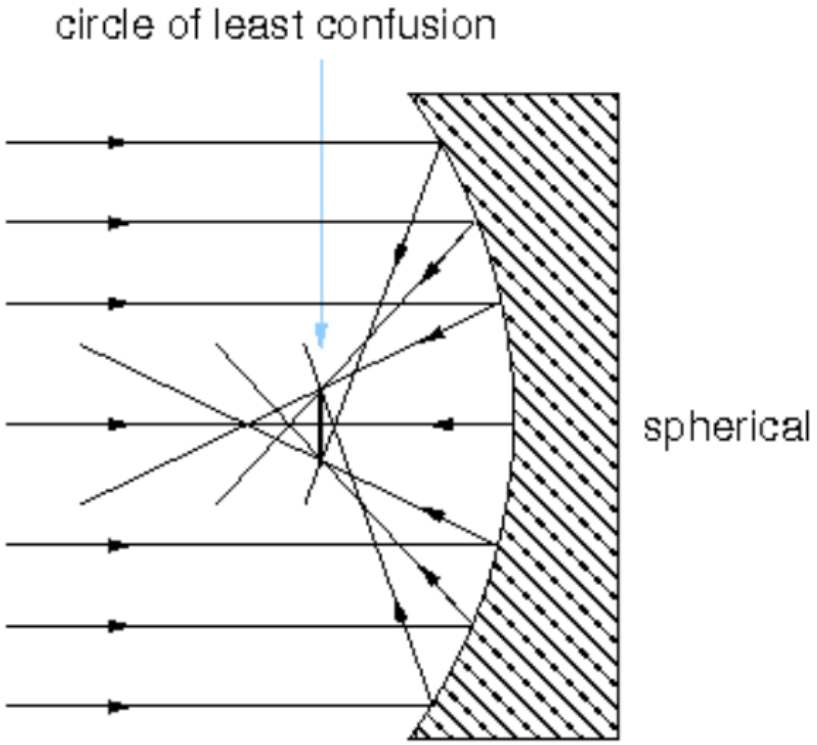
metal plates (high frequency radio);

glass covered with a deposit of

- gold (infra-red);
- silver (optical and near-infrared) or
- aluminium (optical and UV).

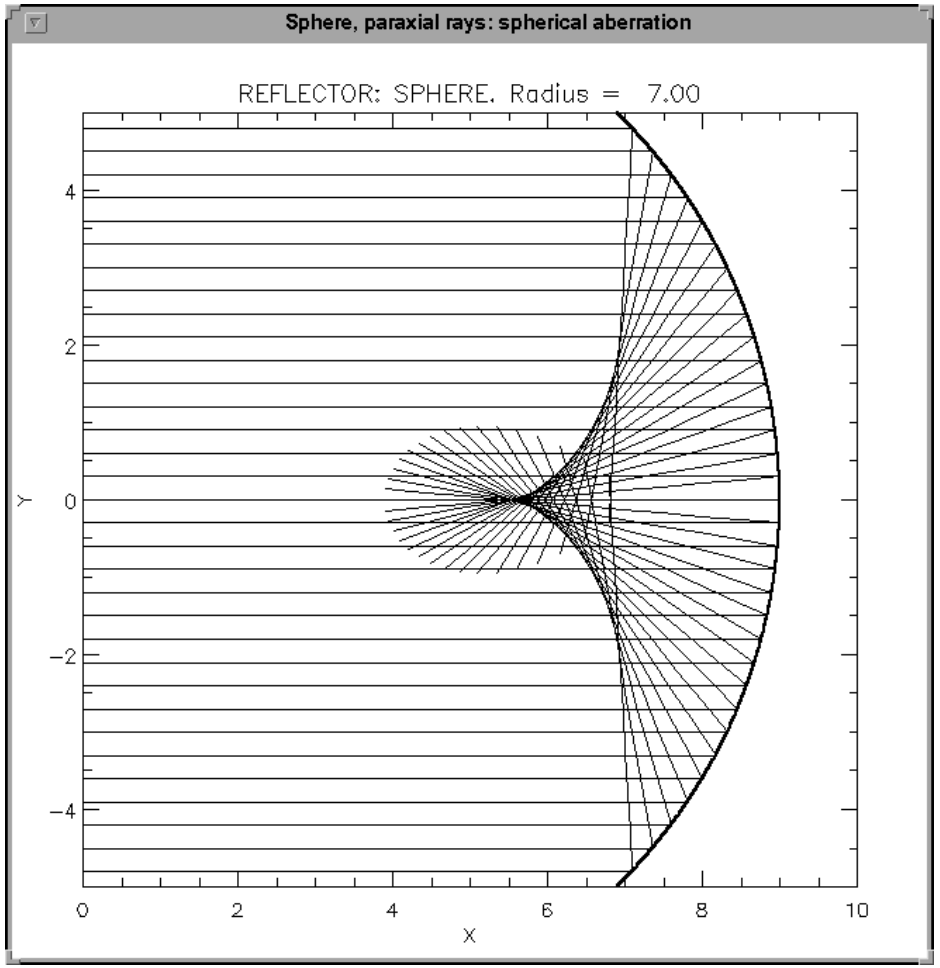
Often there are overcoatings to enhance reflectivity and protect the reflecting surface.

Spherical Aberration (1)



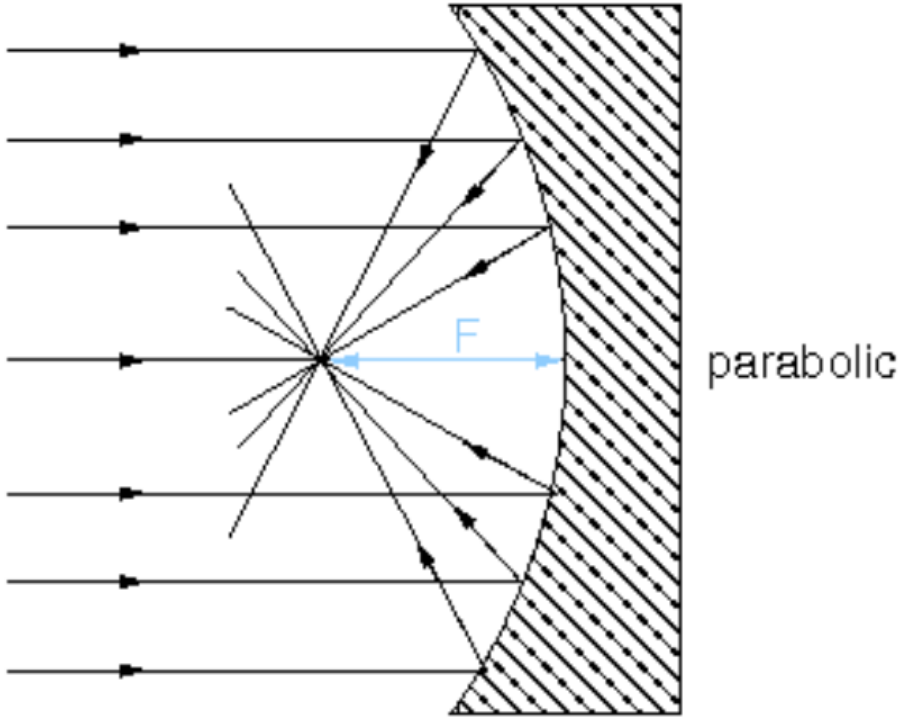
Spherical surface suffers from spherical aberration – rays hitting the outside of the dish come to focus at a different point on the optical axis from those hitting the centre.

Spherical Aberration (2)



Spherical surface suffers from spherical aberration – rays hitting the outside of the dish come to focus at a different point on the optical axis from those hitting the centre.

Paraboloidal reflector



Paraboloidal reflector brings all rays to focus at the same point on the optical axis, and eliminates spherical aberration.

Most telescopes at radio, IR, optical and UV wavelengths are based upon paraboloidal reflectors

A circular aperture of telescopes

- All astronomical observations at wavelengths from mid-frequency radio to X-ray use telescopes, which are a means of bringing a parallel beam of light from a distant source to a focus.
  The majority of telescopes have a circular aperture.



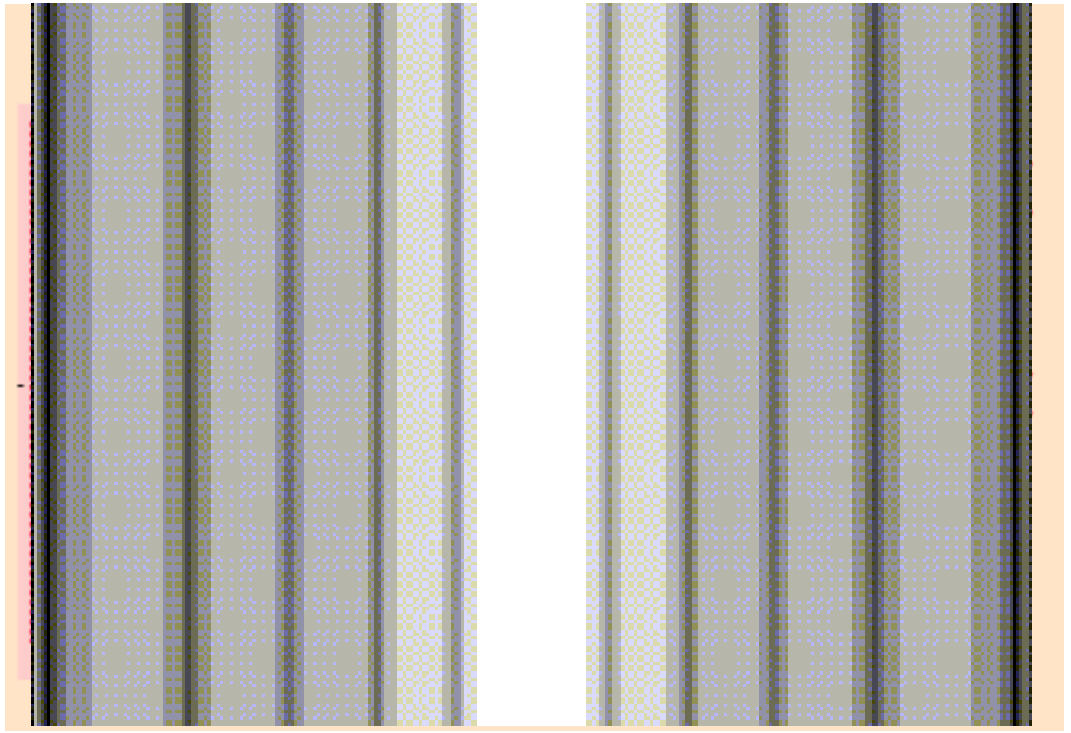


Diffraction through a single slit

- Diffraction places fundamental limit on angular resolution of telescopes.

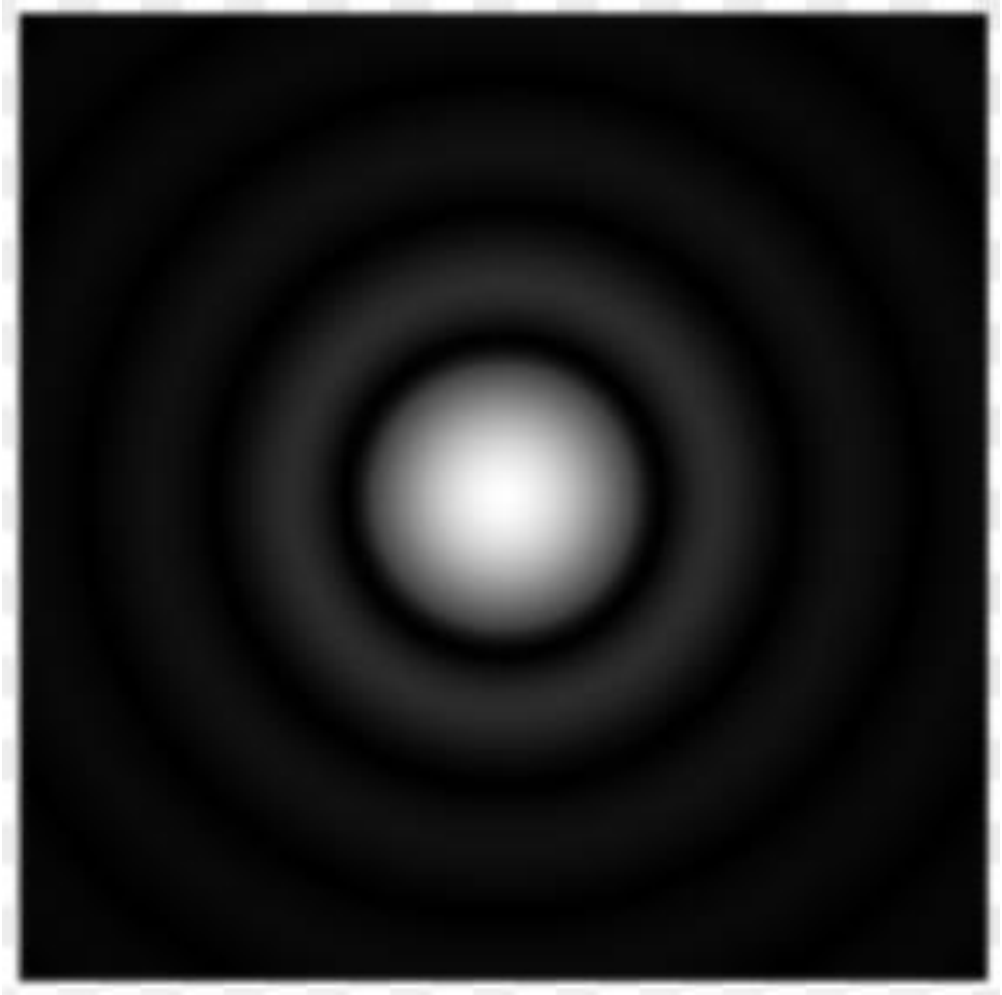
- Diffraction through a single slit (due to the wave nature light): the waves from secondary sources within the slit interfere.

Image of a narrow slit:



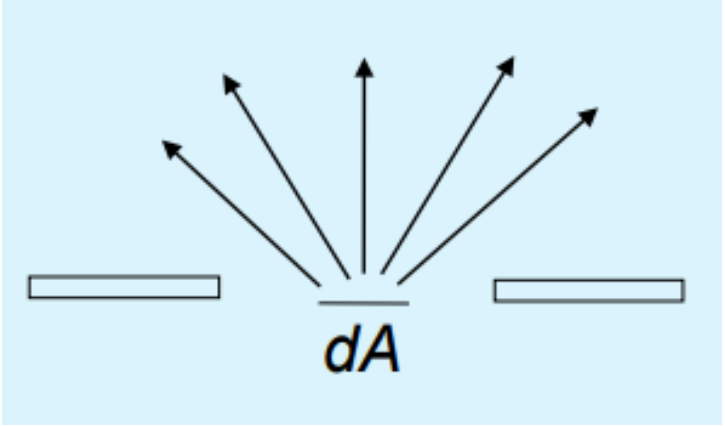
Diffraction problem for a circular aperture

- The interference pattern on the screen behind the circular aperture of a finite size is created similarly to the slit: the waves from different secondary sources in the apperture interfere.
- The resulting pattern is a set of co-centric circular bright and dark bands with the highest peak in the center.



Diffraction and Spatial resolution (1)

- An approach for diffraction by a circular aperture:
  - Take an element of area dA which radiates in all directions. We then need to sum all contributions, over all such elements, at some angle θ.



- Analytical expression for diffraction pattern obtained, for example, in the Kitchin book "Astrophysical techniques"

http://cnx.org/content/m13097/latest/

Does not work anymore

Circular Aperture

The circular aperture is particularly important because it is used a lot in optics. A telescope typically has a circular aperture for example.

We can use the same expression for the E field that we had for the rectangular aperture for any possible aperture, as long as the limits of integration are appropriate. So we can write

$$\vec{E} = \frac{\vec{\varepsilon_A}}{R} e^{i(kR-\omega t)} \int\int_{aperture} e^{-iK(Yy+Zz)/R} \, dy \, dz$$

For a circular aperture this integration is most easily done with cylindrical coordinates. Look at the figure

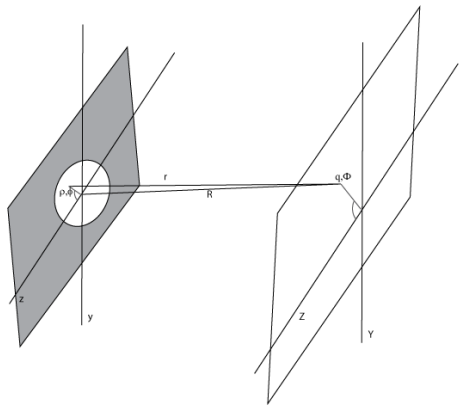


Figure 1.

Then we have

$$z = \rho \cos\varphi$$
$$y = \rho \sin\varphi$$
$$Z = q \cos\Phi$$
$$Y = q \sin\Phi$$

Then

$$Yy + Zz = \rho q \cos\varphi \cos\Phi + \rho q \sin\varphi \sin\Phi$$

or

$$Yy + Zz = \rho q \cos(\varphi - \Phi)$$

and the integral becomes

$$\vec{E} = \frac{\vec{\varepsilon_A}}{R} e^{i(kR-\omega t)} \int_0^a \int_0^{2\pi} e^{-iK\rho q \cos(\varphi-\Phi)/R} \rho \, d\rho \, d\varphi$$

In order to do this integral we need to learn a little about Bessel functions.

$$J_0(u) = \frac{1}{2\pi} \int_0^{2\pi} e^{iu\cos v} \, dv$$

Is the definition of a Bessel function of the first kind order 0.

$$J_m(u) = \frac{1}{2\pi} \int_0^{2\pi} e^{i(mv + u\cos v)} \, dv$$

Is the definition of a Bessel function of the first kind order m.

They have a number of interesting properties such as the recurrence relations

$$\frac{d}{du}[u^m J_m(u)] = u^m J_{m-1}(u)$$

so that for example when $m = 1$

$$\int_0^u u' J_0(u') \, du' = u J_1(u).$$

In order to numerically calculate the value of a Bessel function one uses the expansion

$$J_m(x) = \sum_{s=0}^{\infty} \frac{(-1)^s}{s!(m+s)!} \left(\frac{x}{2}\right)^{m+2s}.$$

Now we want to evaluate the integral

$$\vec{E} = \frac{\vec{\varepsilon_A}}{R} e^{i(kR-\omega t)} \int_0^a \int_0^{2\pi} e^{-iK\rho q \cos(\varphi-\Phi)/R} \rho \, d\rho \, d\varphi$$

which we can do at any value of Φ since the problem is symmetric about Φ. So we can simplify things greatly if we do the integral at $\Phi = 0$

$$\vec{E} = \frac{\vec{\varepsilon_A}}{R} e^{i(kR-\omega t)} \int_0^a \int_0^{2\pi} e^{-iK\rho q \cos(\varphi)/R} \rho \, d\rho \, d\varphi$$

which becomes

$$\vec{E} = \frac{\vec{\varepsilon_A}}{R} e^{i(kR-\omega t)} 2\pi \int_0^a J_0(-K\rho q/R) \rho \, d\rho$$

Now J_0 is an even function so we can drop the minus sign and rewrite the expression as

$$\vec{E} = \frac{\vec{\varepsilon_A}}{R} e^{i(kR-\omega t)} 2\pi \int_0^a J_0(K\rho q/R) \rho \, d\rho$$

To do this integral we change variables

$$w = k\rho q/R$$
$$\rho = \frac{wR}{kq}$$
$$d\rho = \frac{R}{kq} \, dw$$

so that

$$\int_0^a J_0(K\rho q/R) \rho \, d\rho = \int_0^{kaq/R} \left(\frac{R}{kq}\right)^2 J_0(w) w \, dw$$
$$= \left(\frac{R}{kq}\right)^2 \left(\frac{kaq}{R}\right) J_1(kaq/R)$$
$$= a^2 \left(\frac{R}{kaq}\right) J_1(kaq/R)$$
$$= a^2 \frac{J_1(kaq/R)}{kaq/R}$$

So finally we have the result

$$\vec{E} = \vec{\varepsilon_A} \frac{e^{i(kR-\omega t)}}{R} 2\pi a^2 \frac{J_1(kaq/R)}{kaq/R}$$

Or recognizing that πa^2 is the area of the aperture A and squaring to get the intensity we write

$$I = I_0 \left[\frac{2J_1(kaq/R)}{kaq/R}\right]^2$$

If you want to write this in terms of the angle θ then one uses the fact that $q/R = \sin\theta$

$$I(\theta) = I(0) \left[\frac{2J_1(ka\sin\theta)}{ka\sin\theta}\right]^2$$

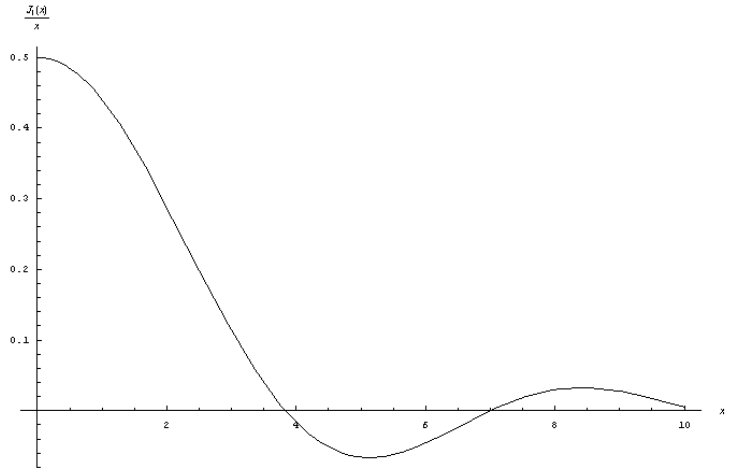


Figure 2.

Above is a plot of the function $J_1(x)/x$. Notice how it peaks at $1/2$ which is why there is the factor of two in the expression for the irradiance. Below is a 3D plot of the same thing (ie. $J_1(r)/r$). Notice the rings.

Diffraction and Spatial resolution (2)

Answer is (first derived in 1835 by Airy):

$$I_\theta \propto [\, J_1(2m)/m \,]^2$$

where $m = (\pi\, r \sin\theta \,/\, \lambda)$,

λ is wavelength

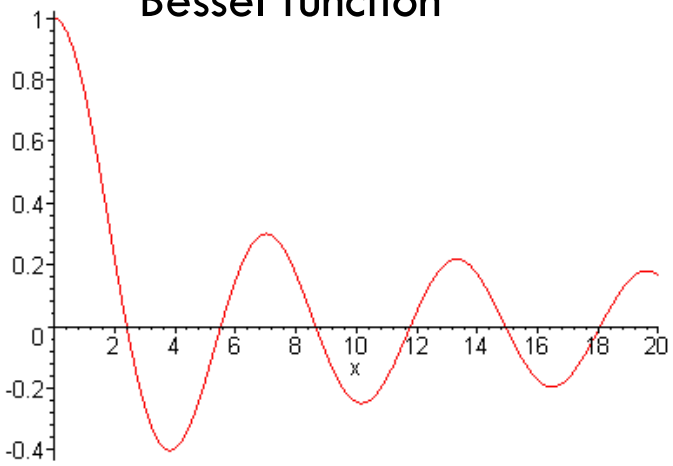
θ is the angle to the normal from the aperture,

r is the radius of the aperture,

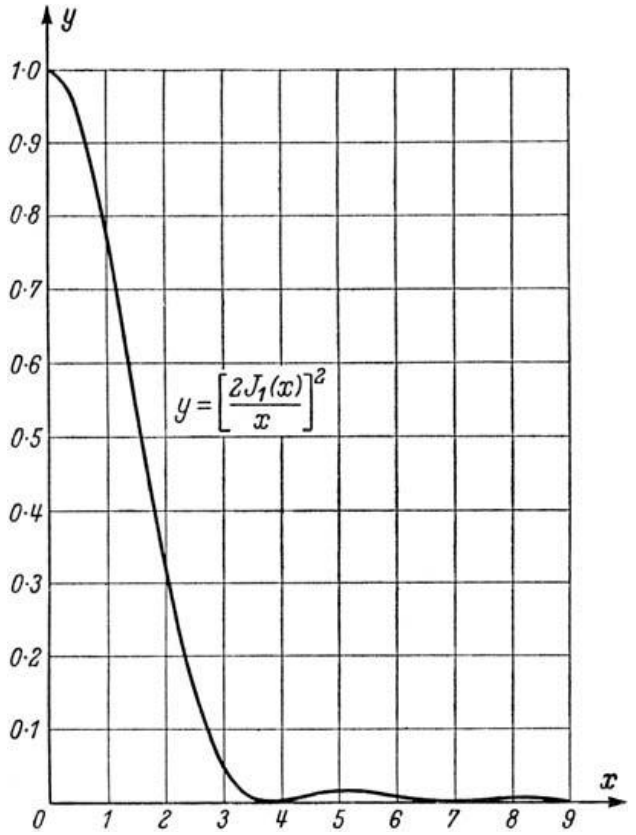
J_1 is a Bessel function of the 1st kind and 1st order.

Diffraction and Spatial resolution (3)

Bessel function



Diffraction of a circular aperture with normalized intensity



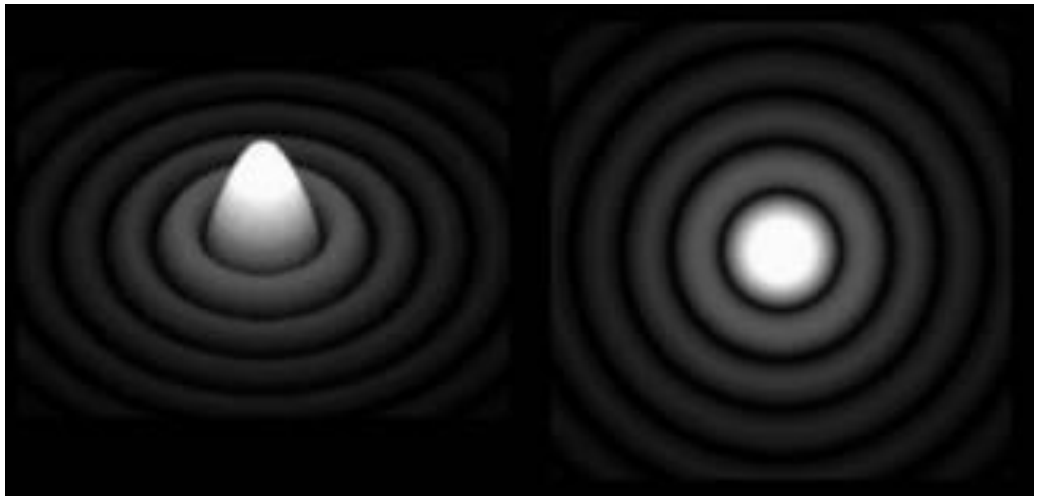
The first zero of the Intensity function occurs at $m = 1.916$, or for small θ:

$$\theta \cong 1.22\,\lambda / D \text{ (radians)}$$

$D = 2r$ is the diameter of the aperture

Airy pattern

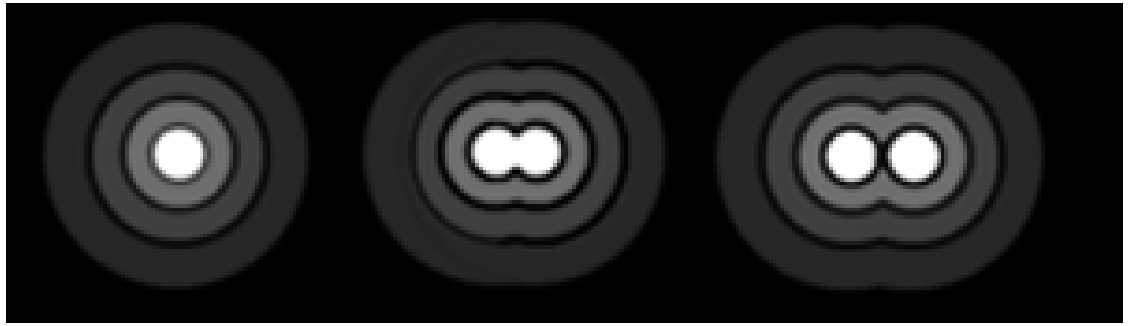
- Airy pattern:



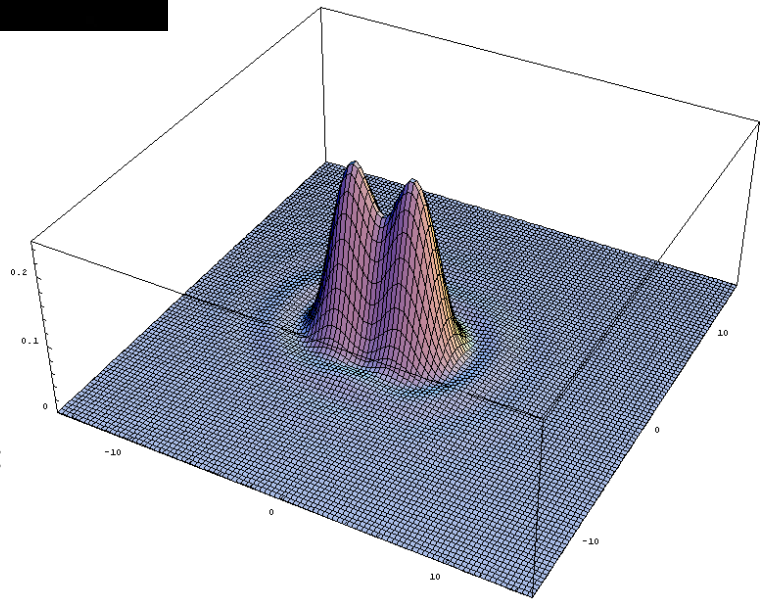
- The 1st maximum is roughly 1.75% of the central intensity, the 2nd – 0.42%.
- 84% of the light arrives within the central peak called the **Airy disk**.

Effect of diffraction on image resolution

- Two distant point sources through a circular aperture:



- **The Rayleigh criterion:** two sources are just resolved if the centre of the 1st Airy pattern is superimposed on the 1st dark ring of the 2nd pattern.



- According to Rayleigh criterion, the minimum resolvable angular separation or angular limit of resolution is given by:

$$\theta \cong 1.22\,\lambda / D \text{ (radians)}$$

λ is wavelength and D is the diameter of the aperture

Diffraction and Spatial resolution (4)

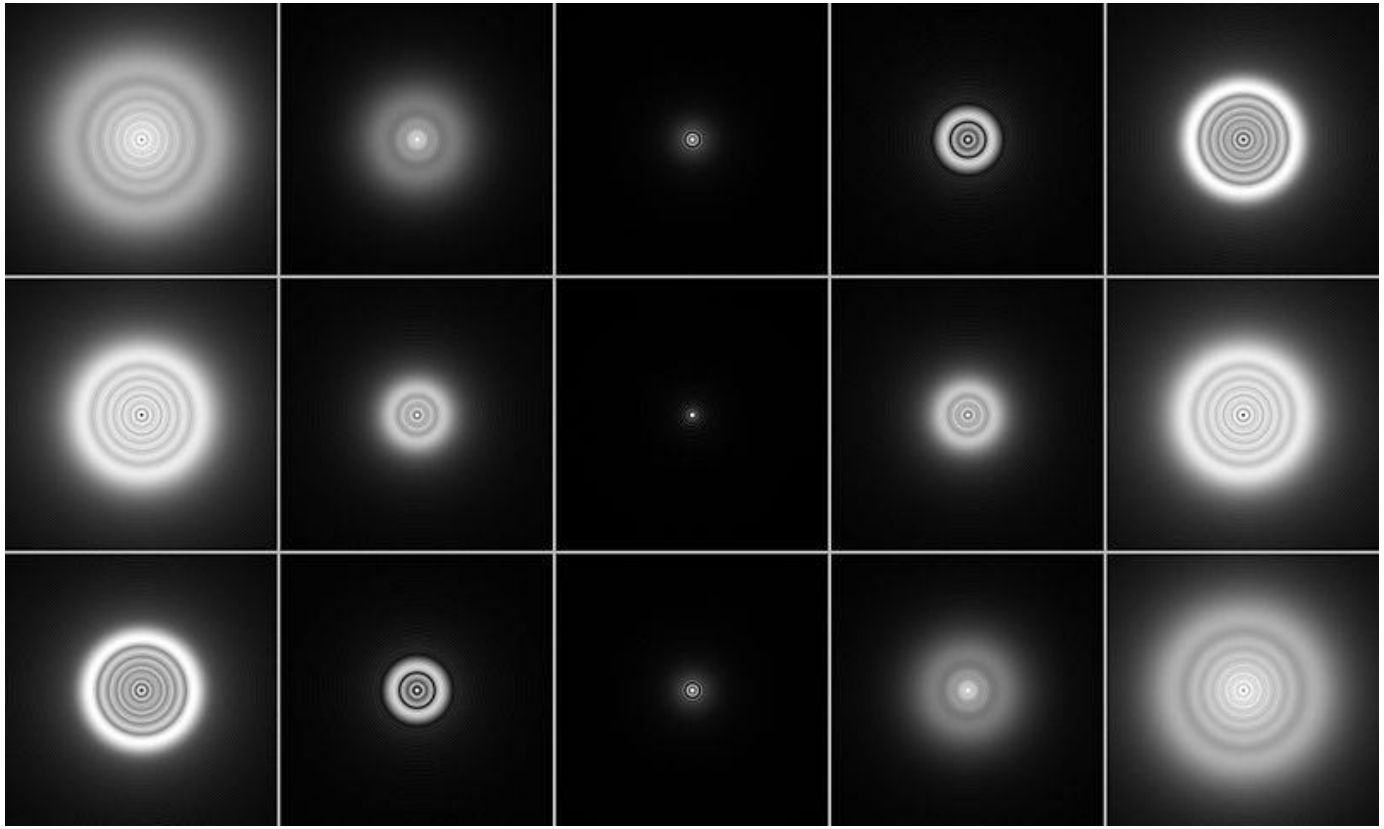
- The **fundamental** limit to the spatial resolution of a circular aperture telescope is set by the size of the Airy diffraction pattern, but **in practice** the resolution is often worse than that, particularly at wavelengths shorter than 10 μm (100000 Å).
- Causes are:
  - Distortion of the wavefront by the atmosphere (seeing);
  - Imperfections in manufacture or support of the reflecting surface;
  - Aberrations.

Aberrations in reflecting telescopes

- **Spherical aberration** – caused by different parts of the mirror bringing rays to focus at different points on the optical axis.
- **Coma** – teardrop or comet shaped aberration proportional to the square of the distance of the image off axis.
- **Astigmatism** – cylindrical wavefront distortion resulting in an ellipticity of the image which is orthogonal on different sides of focus. Can occur on-axis if the mirror is stressed, off axis it results from magnification at the secondary (i.e. if final focal length >> primary focal length) and is proportional to the square of the distance of the image off axis.
- **Curvature of the field of the image**

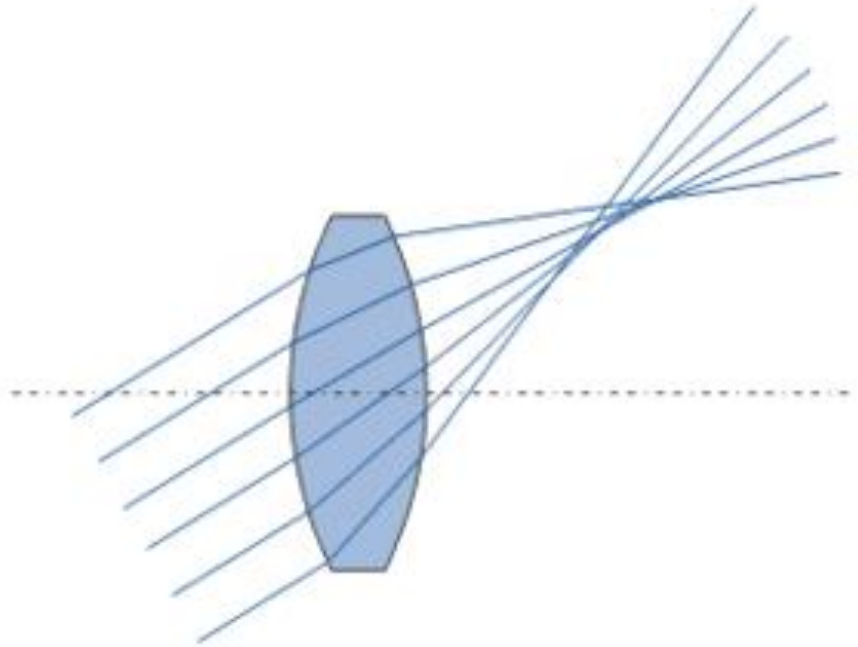
Spherical aberration

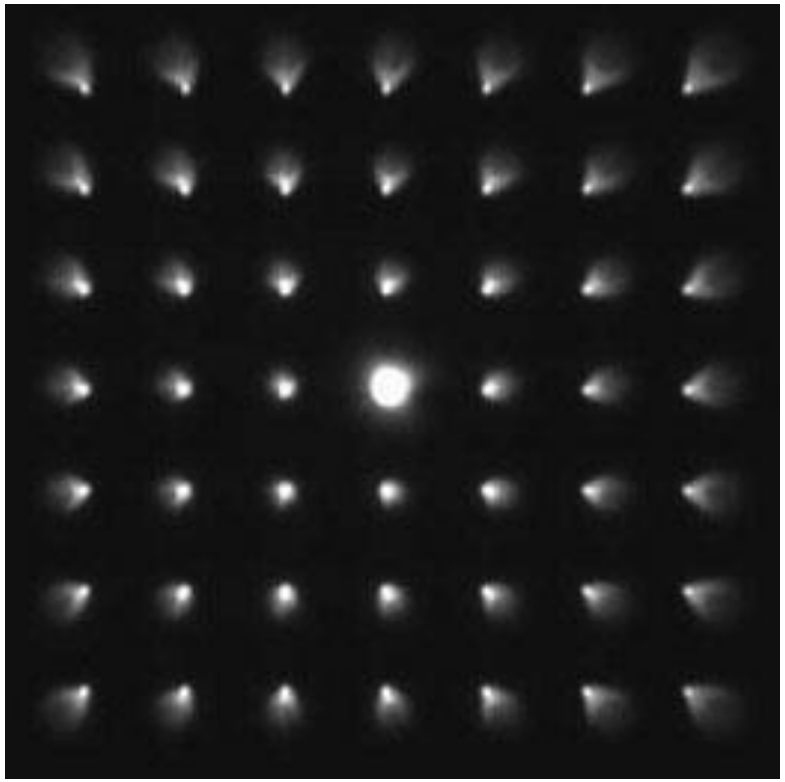
- **Spherical aberration** – caused by different parts of the mirror bringing rays to focus at different points on the optical axis.



Coma

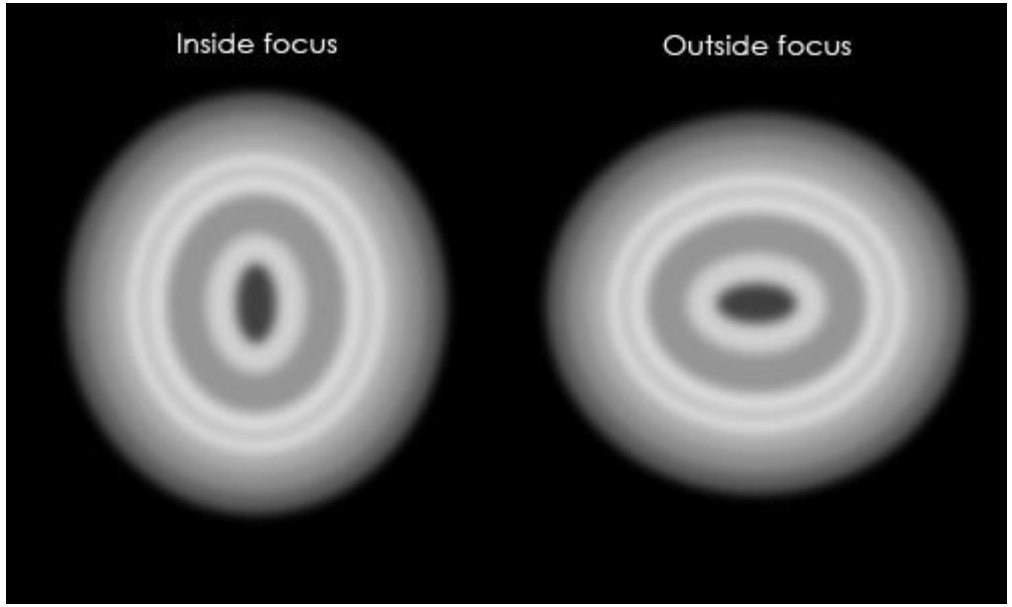
- **Coma** – teardrop or comet shaped aberration proportional to the square of the distance of the image off axis.





Astigmatism

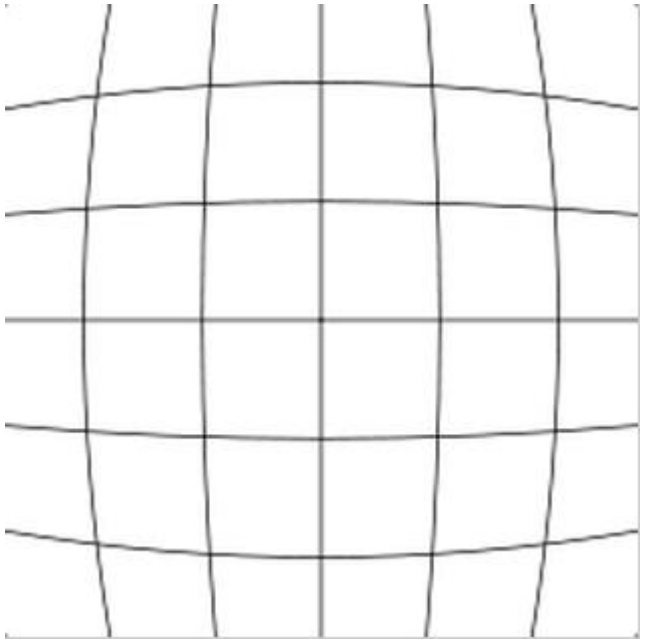
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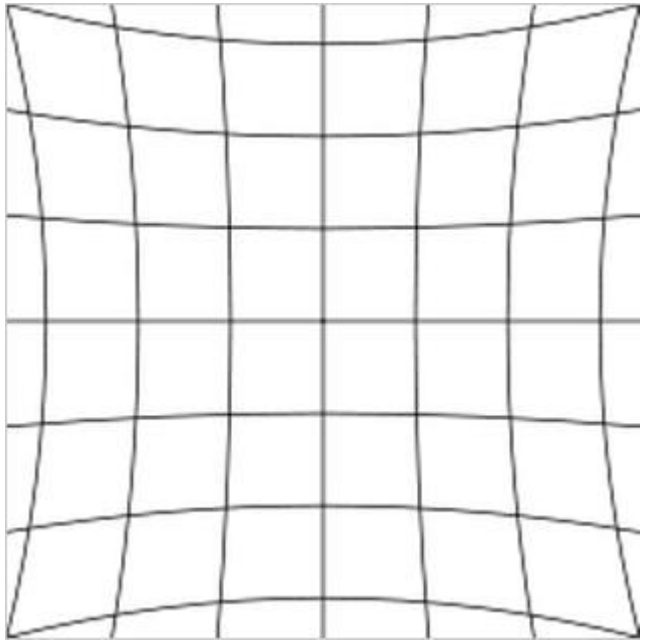


Curvature of the image field

- **Curvature of the field of the image –**

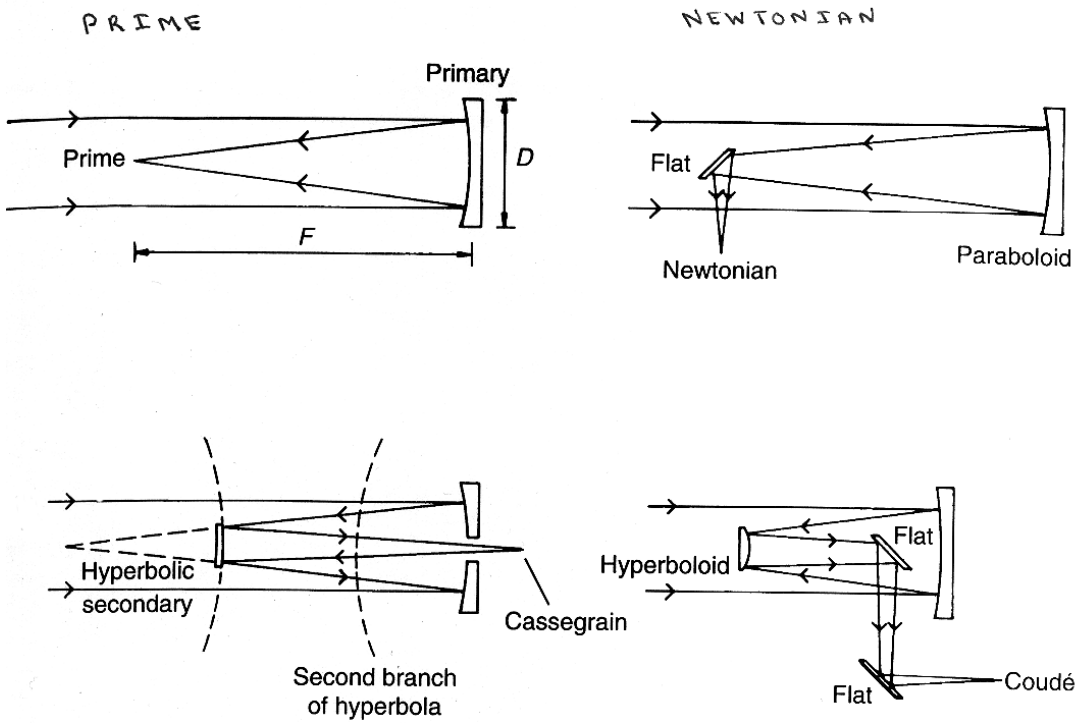
  Barrel and Pincushion distortions of the image





Optical configurations of telescopes

Here "optical" does not mean wavelength, but the way the system is configured to accept rays of radiation.

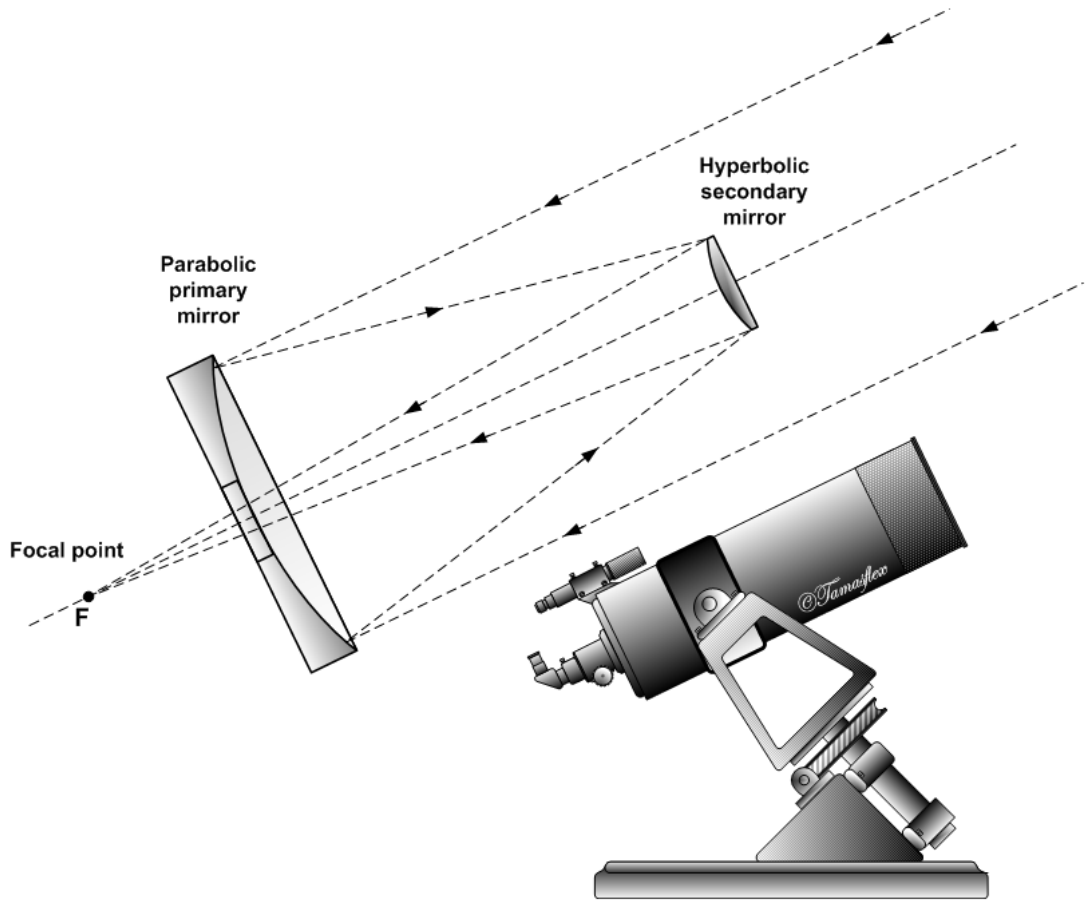


Optical configuration

- **Prime Focus –** detector or receiver is at the focus of the paraboloidal reflector
- **Cassegrain –** a convex hyperboloidal secondary mirror or reflector is used to bring the rays to a focus below a hole in the primary.
- **Gregorian –** the secondary is now a concave elliptical mirror or reflector above the primary focus. **Disadvantage** is less compact than Cassegrain.
- **Nasmyth –** like Cassegrain except that a folding flat is added above or below the primary

Cassegrain optical telescope

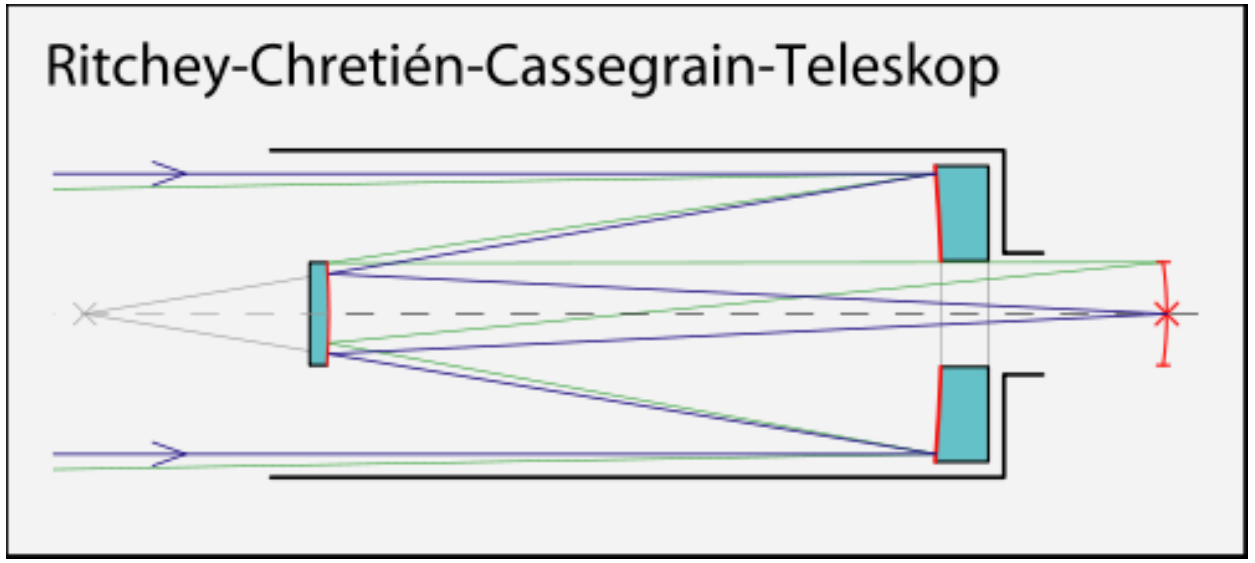
Schematic of Cassegrain optical telescope



Ritchey–Chrétien telescope

Schematic of Ritchey–Chrétien telescope

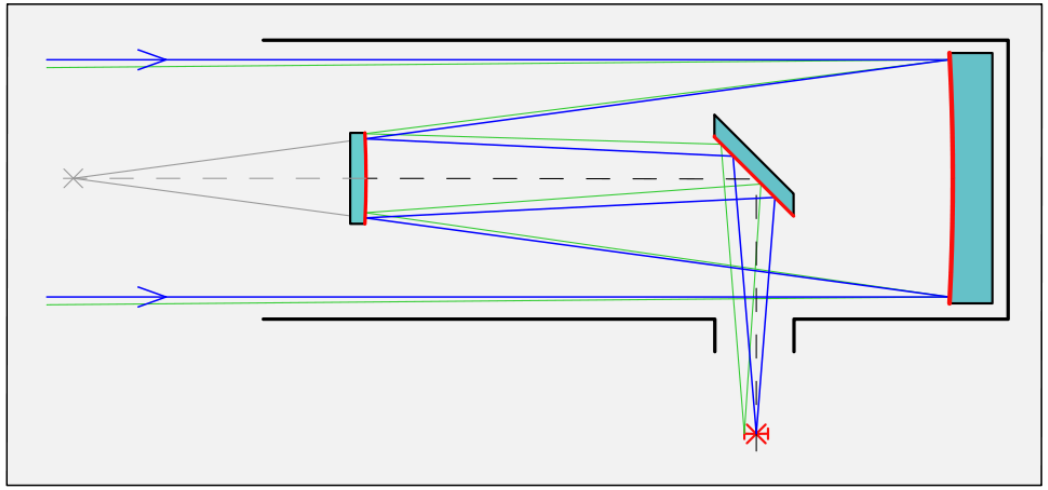
- Both mirrors, a primary and a secondary, are hyperbolic.
- Designed to eliminate coma, thus providing a large field of view compared to a more conventional configuration.
- But more expensive.



Nasmyth telescope

Schematic of Nasmyth telescope

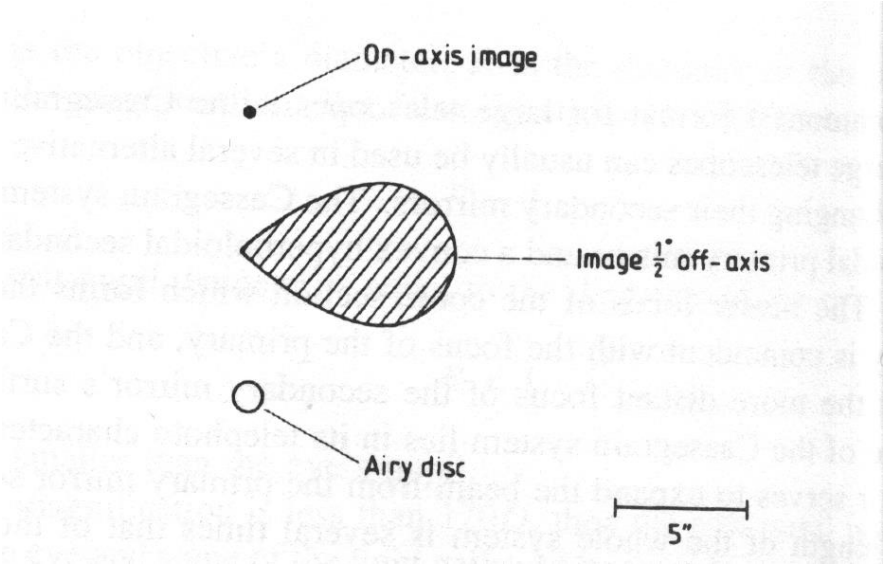
- No hole in the primary mirror is needed.
- The eyepiece or instrument does not need to move with the telescope. This has significant advantages for heavy instruments typically used at research observatories.
- Most modern research telescopes can be configured into a Nasmyth telescope.



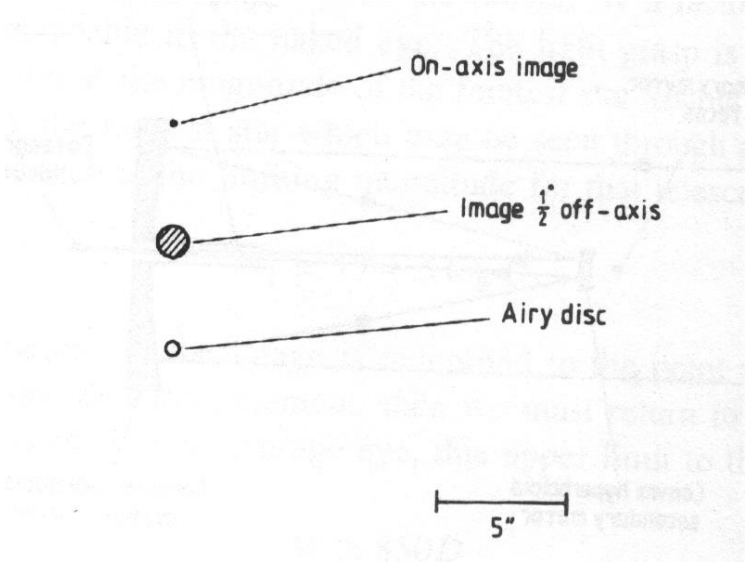
Aberrations in reflecting telescopes

- **Cassegrain** system with paraboloidal primary provides perfect images on-axis, but off-axis coma restricts the useful field of view.
- **Ritchey-Chrétien** optical system
  [where the primary is deepened into a hyperboloid, and a stronger hyperboloid is used as the secondary] allows the coma and astigmatism to be balanced over a wider field.

Wide Field telescopes



Cassegrain optics with paraboloidal primary



Ritchey-Chrétien optics with hyperboloidal primary

Image quality on- and off-axis

Optical configuration

- **Radio telescopes** are most commonly prime focus, occasionally Cassegrain, Gregorian or Nasmyth.
- **NIR, Optical, and UV telescopes** are most commonly Cassegrain of different types, mostly Ritchey-Chrétien, often configurable with interchangeable top end.

Examples of Cassegrain radio antennas

The 14 m Radio Telescope at Metsähovi



Examples of Cassegrain radio antennas

The 32m Radio Telescope at Cambridge



Examples of Cassegrain radio antennas

The 70 meter dish at Jet Propulsion Laboratory (JPL) Goldstone antenna complex.



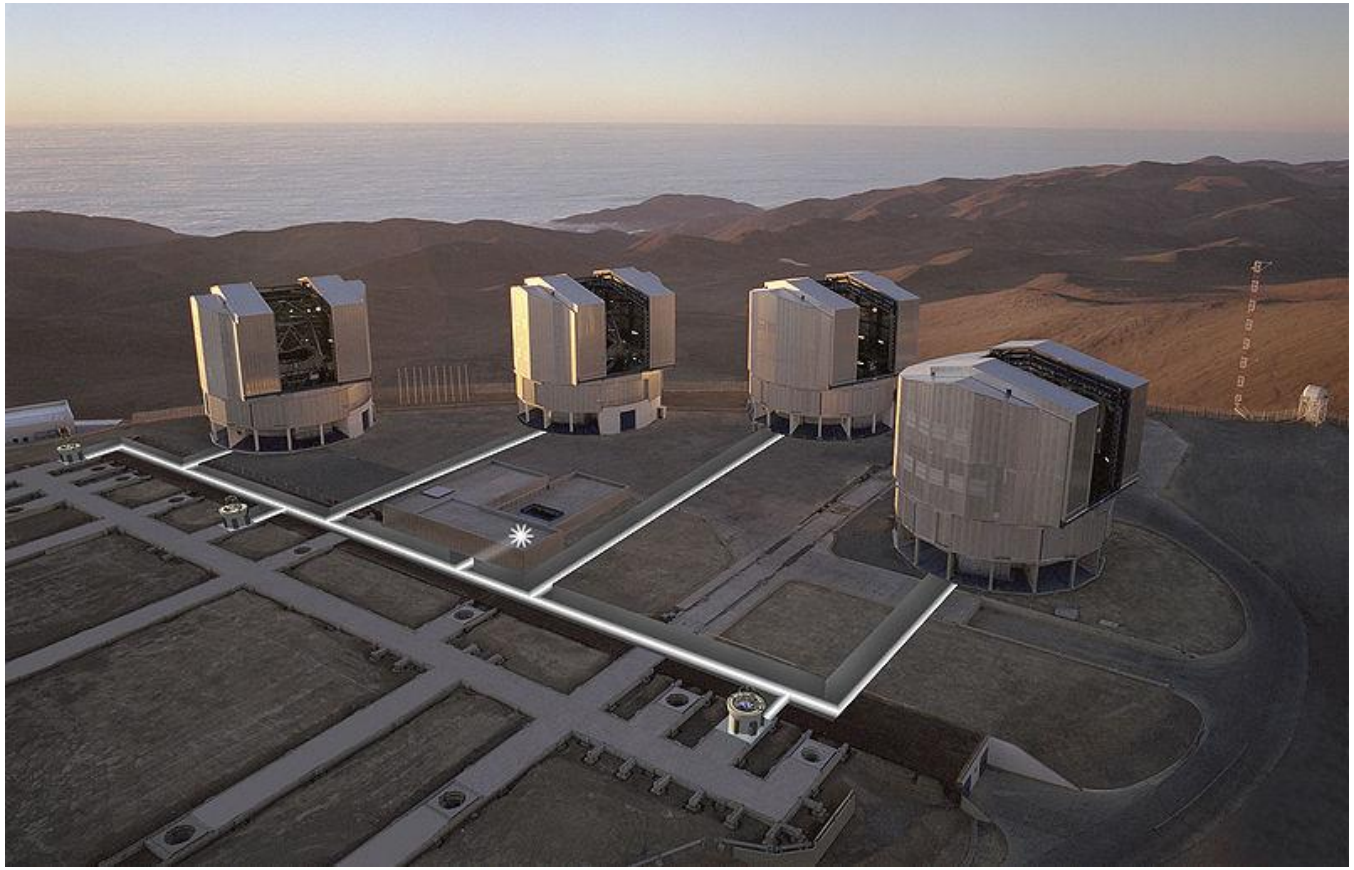
Examples of large Ritchey–Chrétien telescopes

Hubble Space Telescope (2.4-m)



Examples of large Ritchey–Chrétien telescopes

The four 8.2-m telescopes comprising the Very Large Telescope in Chile



Examples of large Ritchey–Chrétien telescopes

The 10.4 m Gran Telescopio Canarias at Roque de los Muchachos Observatory



Examples of large Ritchey–Chrétien telescopes

The Nordic Optical Telescope (2.56-m)



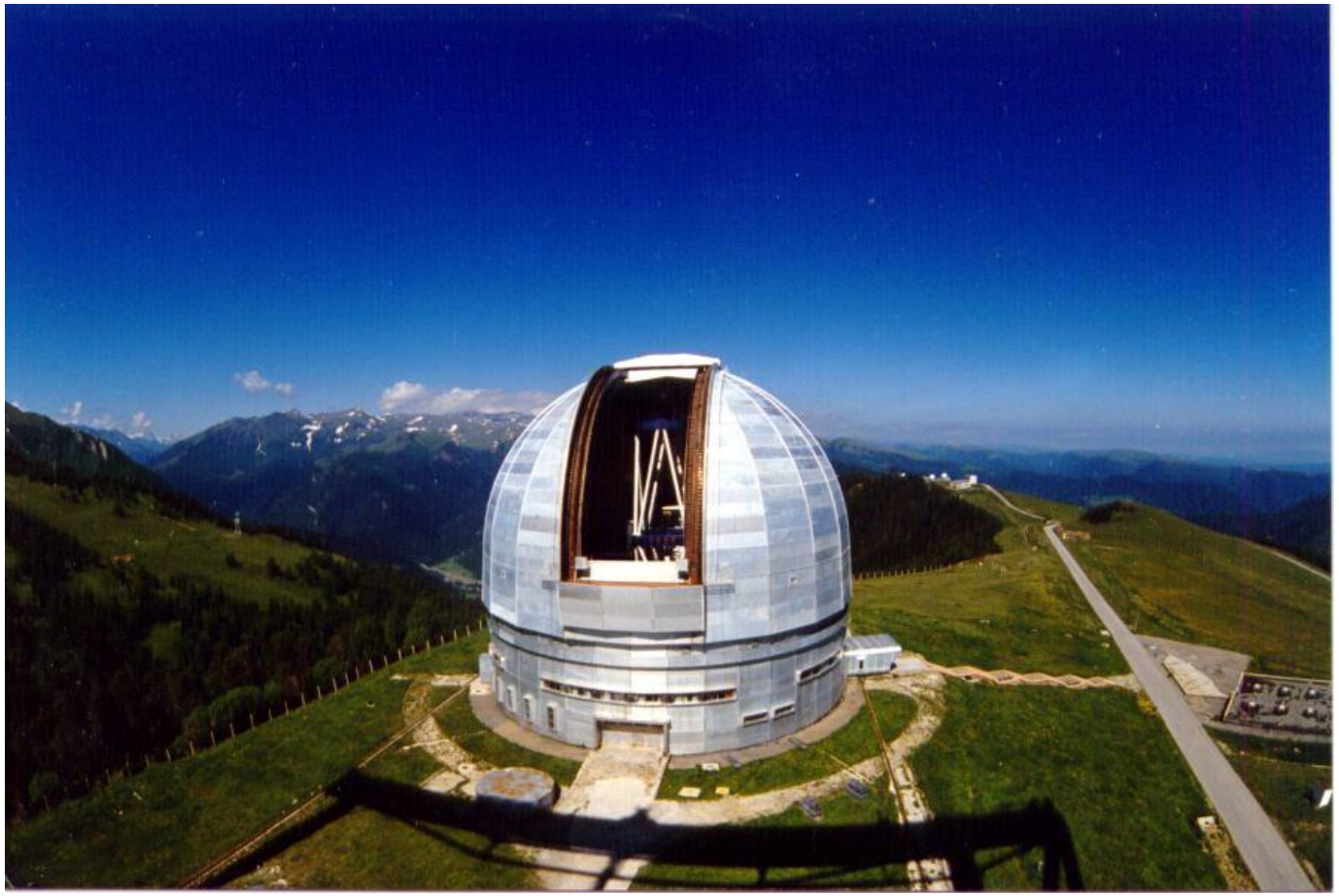
Examples of large Ritchey–Chrétien telescopes

The two 10-m telescopes of the Keck Observatory



Examples of Nasmyth telescopes

The 6-m alt/az mounted telescope (Russia)



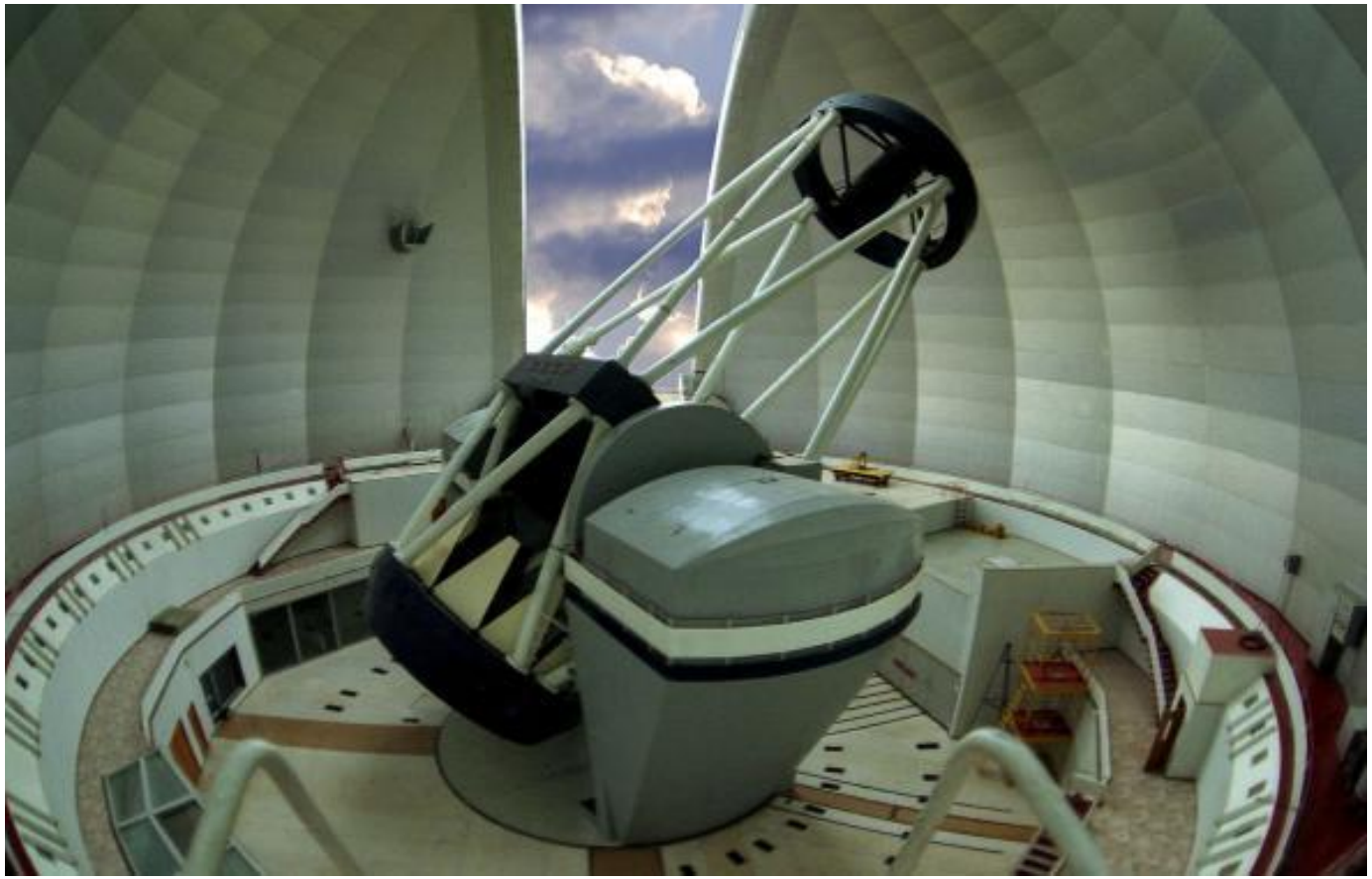
Examples of Nasmyth telescopes

The 6-m alt/az mounted telescope (Russia)



Examples of Nasmyth telescopes

The 6-m alt/az mounted telescope (Russia)



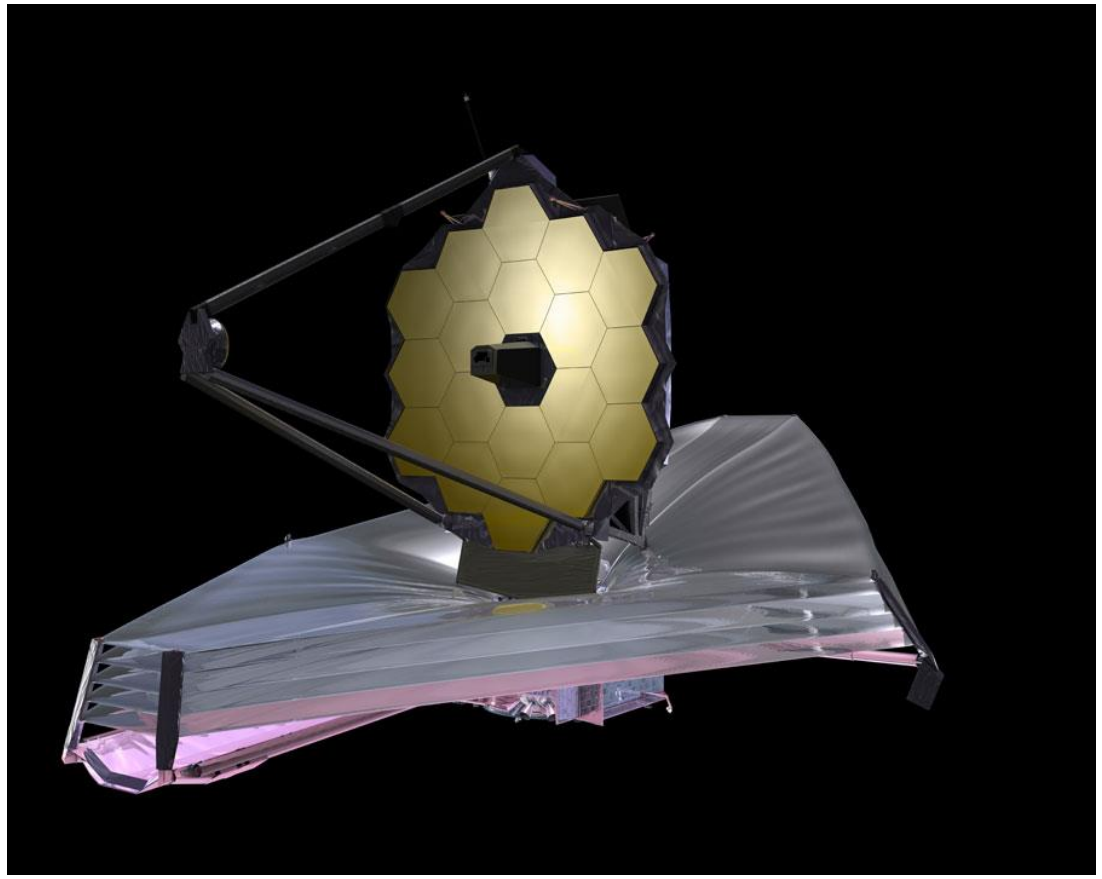
Ultra-violet telescopes

- Ultra-violet radiation at $\lambda < 3100$ Å is absorbed by the atmosphere. Ultra-violet telescopes must be in space, but their design is otherwise similar to optical telescopes.
- The Hubble Space Telescope is the premier ultra-violet telescope.

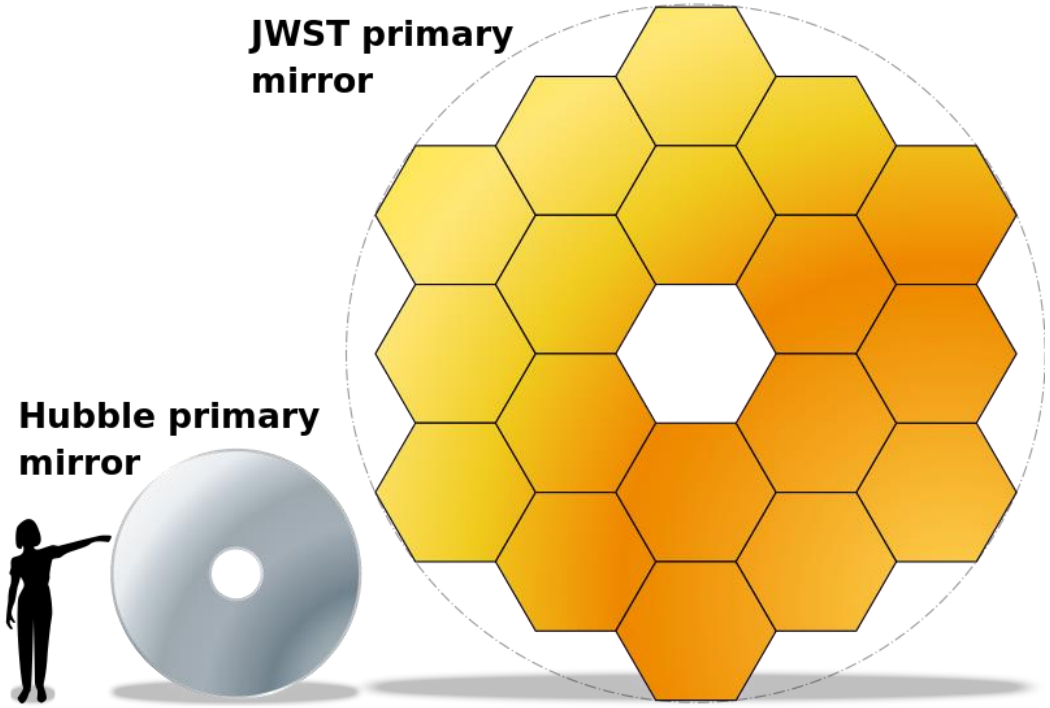
Infra-red telescopes

- Basic principles as for optical telescopes, but the differ in a number of ways. Basic driver for design is that the background comes predominantly from the telescope not from the sky:
  - Choose high, cold sites (e.g. South Pole, Hawaii)
  - Minimize the mechanical structure, telescope structure radiates so that the less are it presents to the detector the better.
  - Low emissivity coatings (gold or silver rather than aluminum) on mirrors.
  - Use rapid beam switching (e.g. with chopping secondary) to subtract of background contribution.
  - Use long focal length, cold baffles and undersized secondary mirror to avoid background radiation reaching the detector.

James Webb Space Telescope

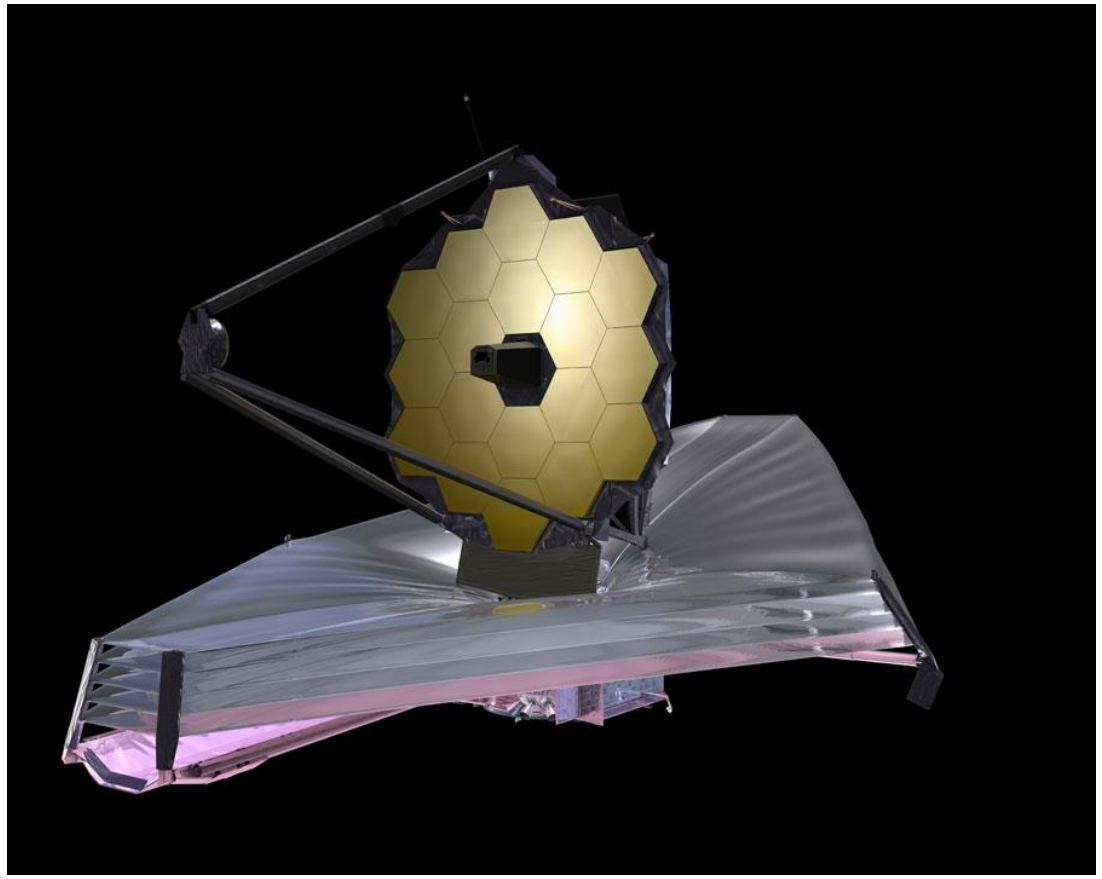


Primary mirror consists of 18 hexagonal mirror segments, which combined create a 6.5-m diameter mirror.



JWST observes from long-wavelength visible light through mid-infrared (0.6–28.3 μm).
The telescope must be kept extremely cold, below 50 K.

James Webb Space Telescope



- JWST's optical design is a three-mirror anastigmat (Korsch telescope).

- Homework: Plot a scheme and explain differences with the classical Cassegrain telescope. What are the advantages?

Telescope mountings

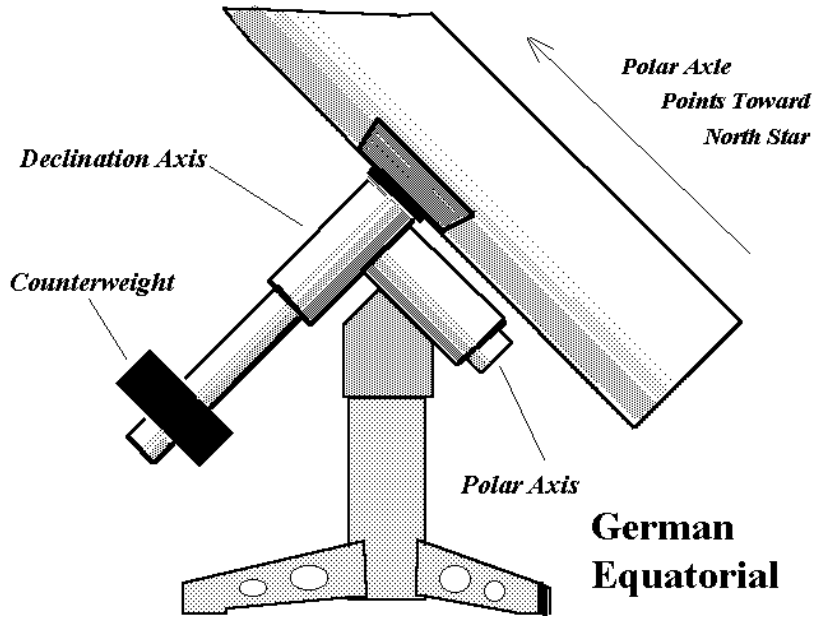
Practical telescope configurations

- Mechanical mounting

  In the past optical telescopes were mounted in an *equatorial* configuration, where one axis points at the north or south pole, and the other axis was horizontal and perpendicular to the polar axis. In this configuration the telescope needs to be driven at a constant angular velocity about the polar axis to compensate for the rotation of the earth and track a source.

Practical telescope configurations

An equatorial mounting



An alt-azimuth mounting

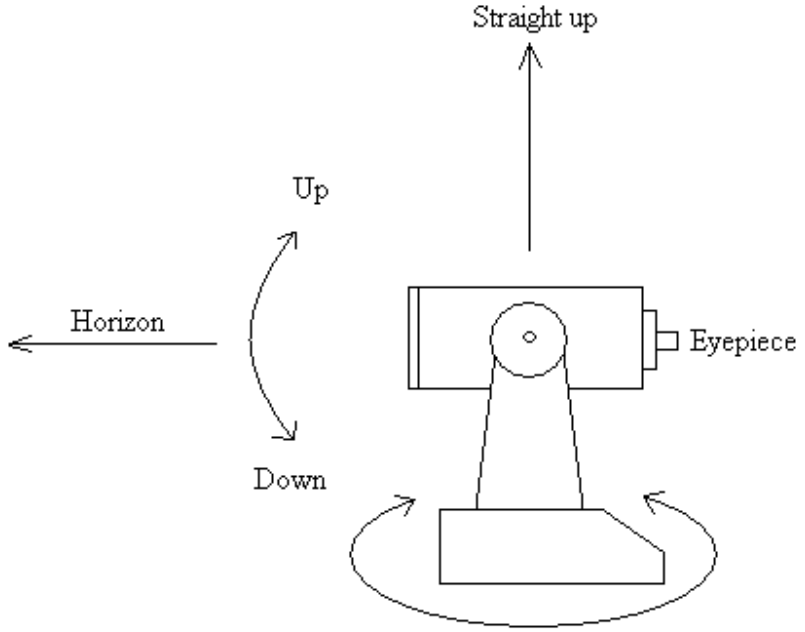


Fig. 1

Practical telescope configurations

All large radio telescopes, and most modern optical and infrared telescopes are mounted in an *altitude-azimuth* configuration. Here the axes are vertical and horizontal, which is mechanically simpler and cheaper to construct. To track a source across the sky, both axes need to be driven at variable angular velocity, but with modern computers this is possible and cheap compared with the difference in the mechanical cost.



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 3

Vitaly Neustroev

Maintaining figure and alignment

- In order to maintain the image quality at the limit set by diffraction (or the atmosphere) it is important to:
  - maintain the paraboloidal figure of the primary reflector
  - maintain the alignment of the optical axes of the primary, secondary and any tertiary reflector with power (i.e. non-flat)
  - maintain the relative orientation of the reflectors, i.e they must not tilt with respect to each other.

If you do not do this – you get aberrations (coma and astigmatism)

Homologous deformation

- Any telescope designed to operate at a wavelength λ should have its surface manufactured to a tolerance of $\lambda/20$ to maintain diffraction limited performance.
- It should maintain this accuracy as its orientation changes, i.e. as it tracks a source across the sky.
- Mechanically this is possible for small steerable dishes (10 metres diameter or less) but difficult for larger structures.

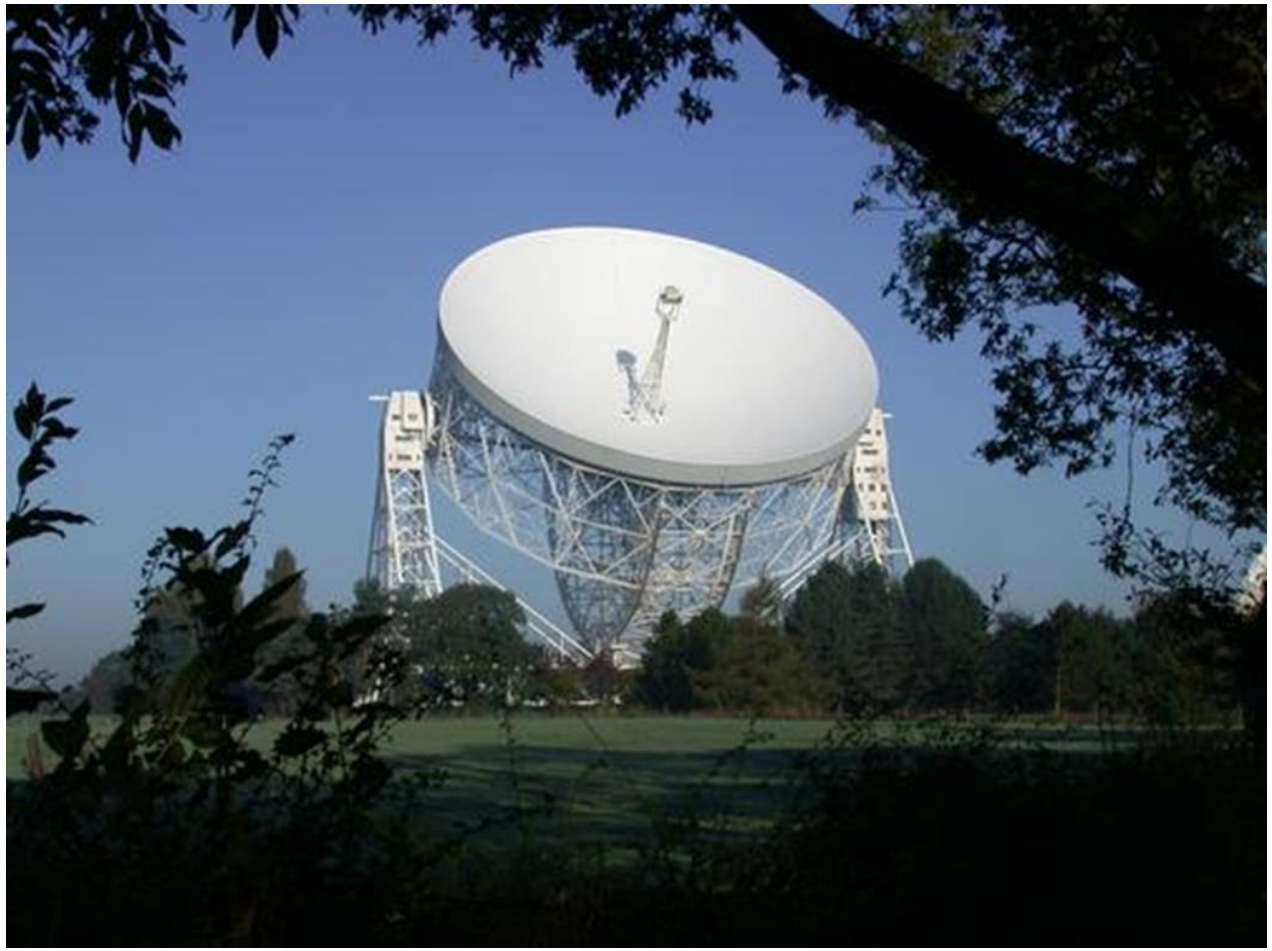
Homologous deformation (radio telescopes)

- Strictly what is required is not that its figure should not change with orientation, but that its figure should remain a paraboloid of rotation.
- Large radio telescopes are designed to deform but to remain paraboloids.
- The Effelsberg 100 metre diameter steerable dish deforms by up to 6cm, but remains a paraboloid to < 0.4 mm, so it performs to specification at wavelengths as short as 1cm.

Radio telescope performance

- Radio telescope performance in terms of maintaining the figure is usually limited by thermal expansion effects and by wind. For this reason, many modern high-frequency radio telescopes are not made from steel or aluminium, but from more exotic materials such as Carbon Fibre or other composite materials.

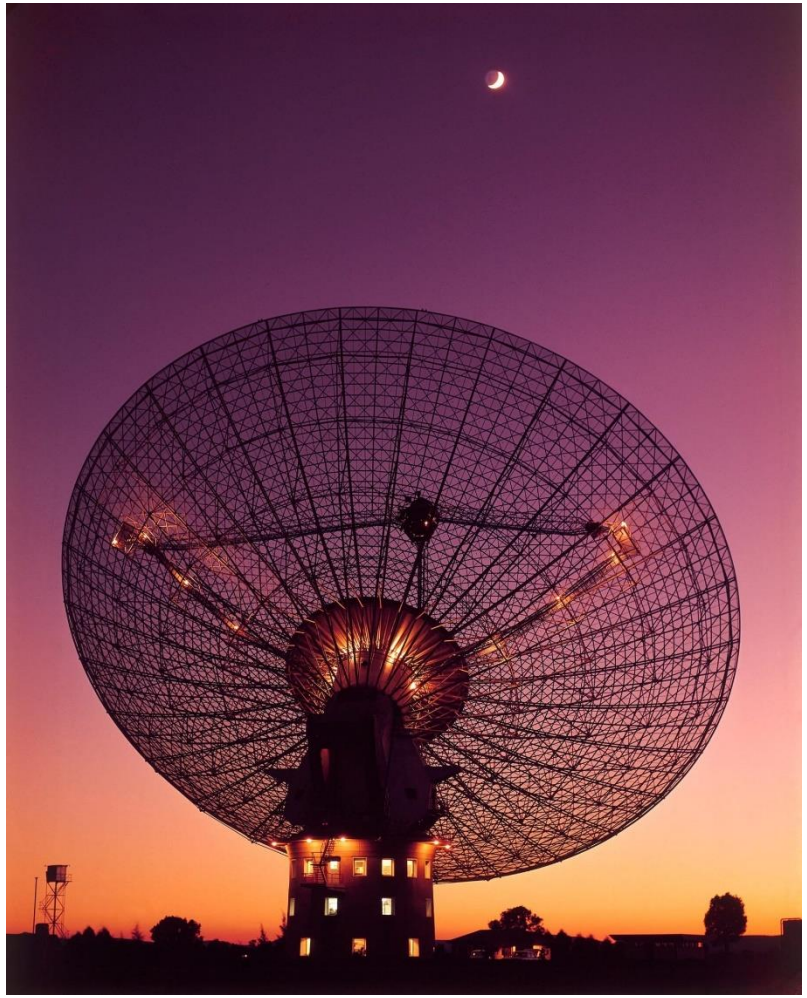
Lovell telescope at Jodrell bank



76.2 metres (250 feet) in diameter

Parkes Radio telescope, Australia

64-m diameter parabolic dish





Above a modern photo, to the left in 1969.
What is the difference? Why?

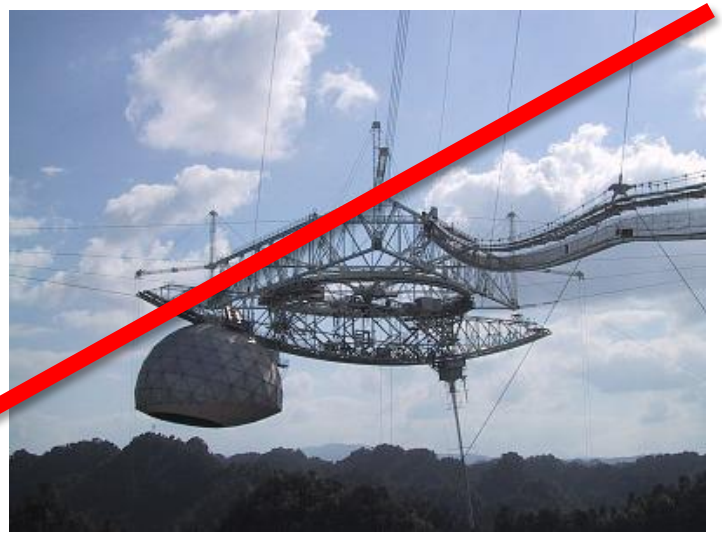
RATAN-600, Russia



A 576 m diameter circle of rectangular radio reflectors and a set of secondary reflectors and receivers

Arecibo radio telescope, Puerto Rico

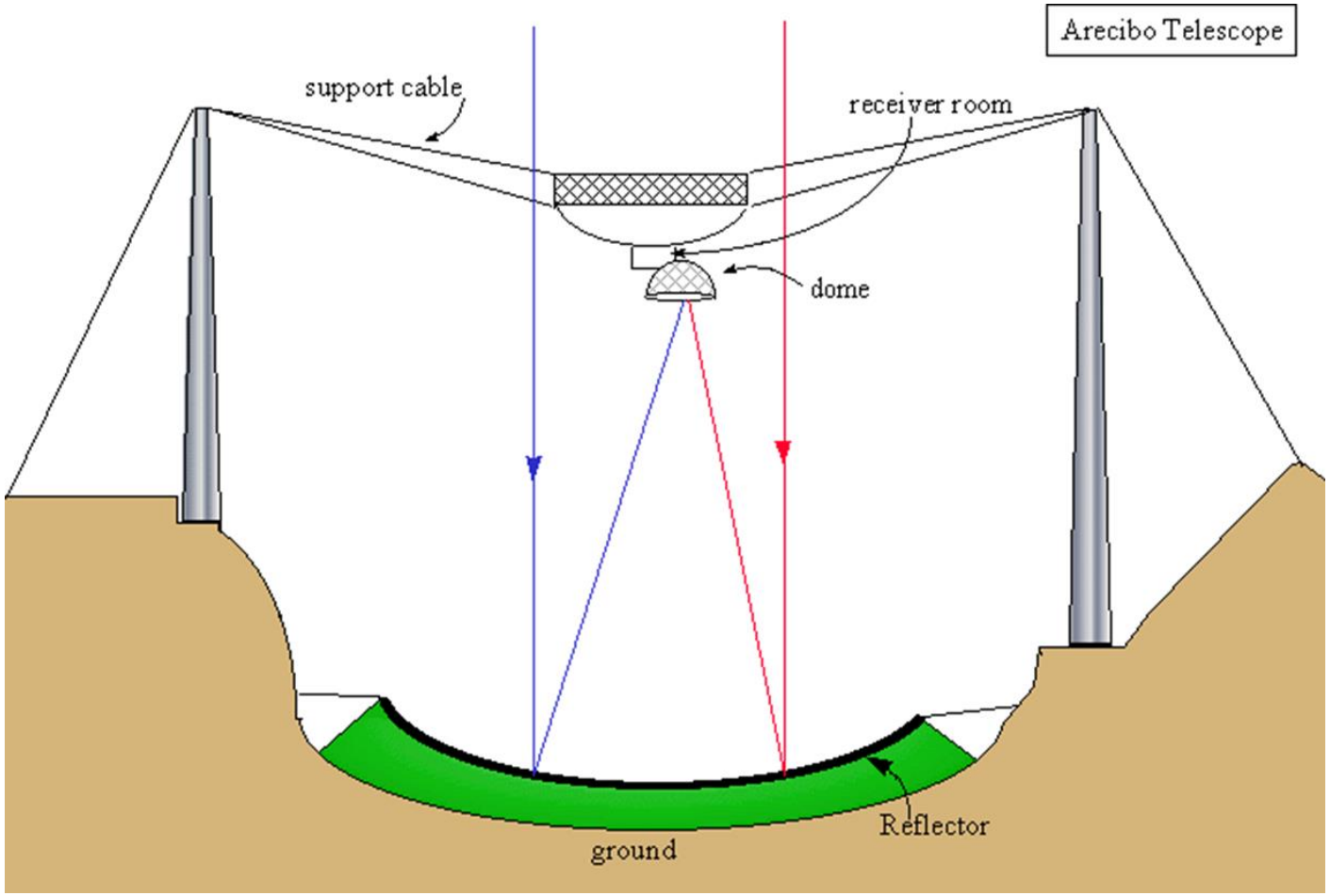






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Arecibo radio telescope, Puerto Rico



A 305 m spherical reflector

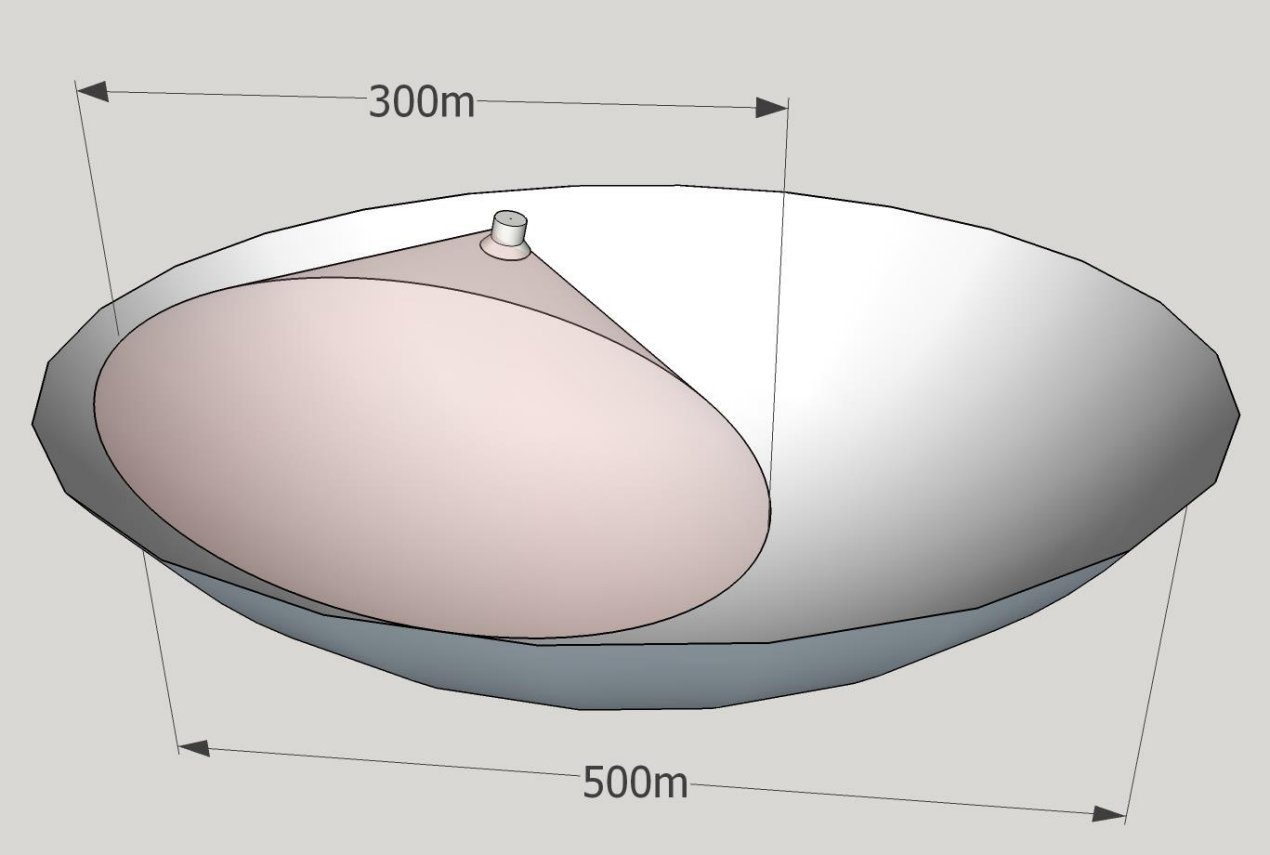
A new king of radio-astronomy – FAST (China)

Five-hundred-meter Aperture Spherical Telescope (FAST)

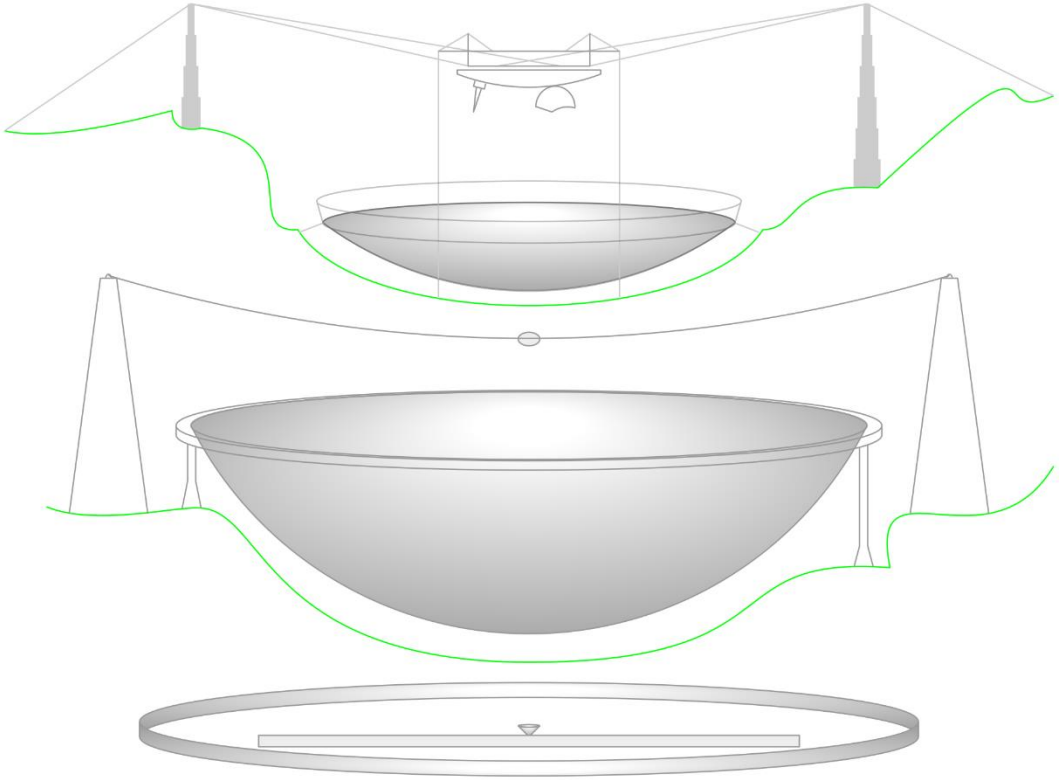


FAST (China)

- Diameter: 500 m
- Illuminated diameter: 300 m



Comparison of the Arecibo (top), FAST (middle) and RATAN-600 (bottom) at the same scale.

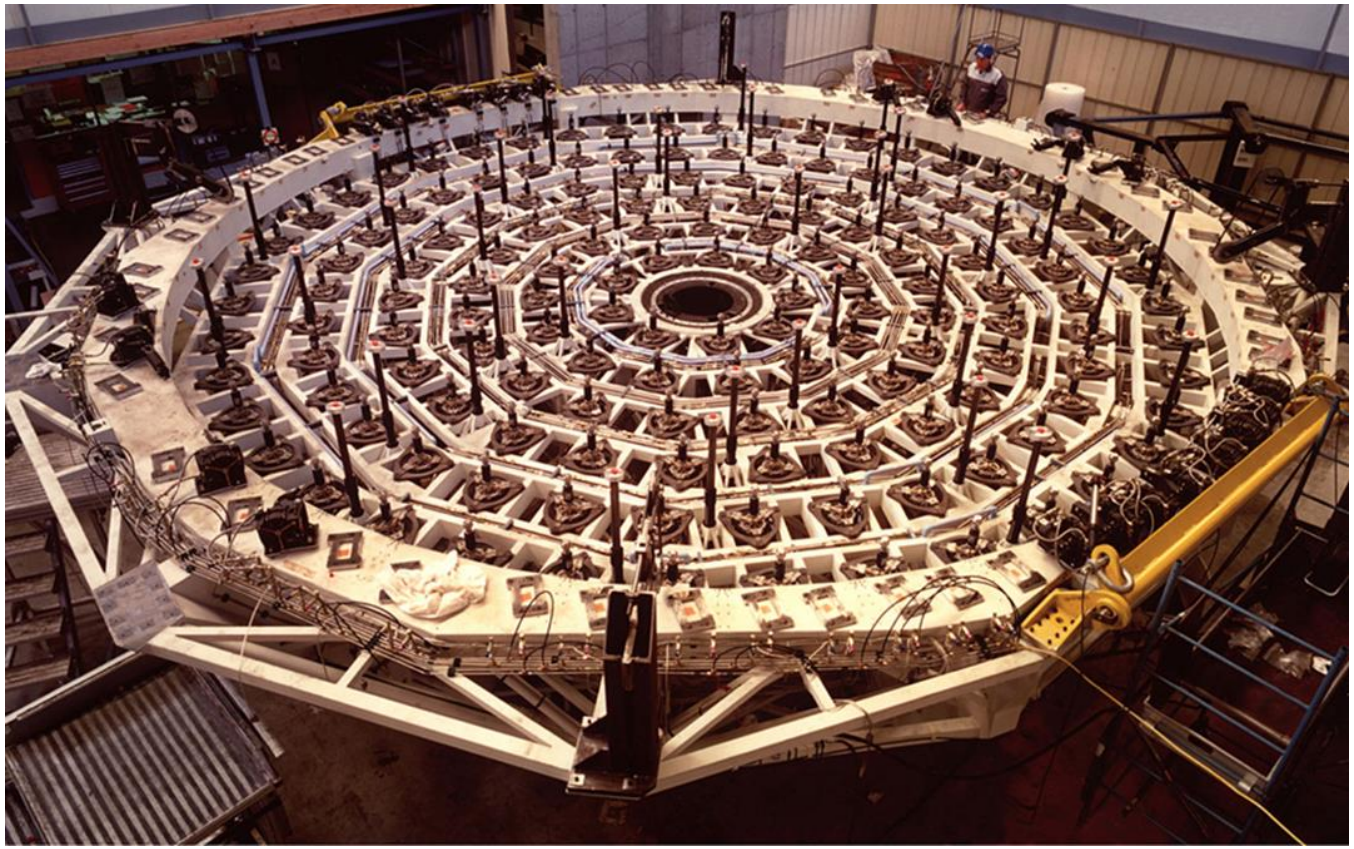


Optical telescopes - active mirror support

- In an optical telescope the primary reflector is usually a zero expansion glass surface coated with a reflecting metal. In small telescopes this can be rigid and maintain its structure.
- In larger telescopes the deformations are modelled as a function of telescope attitude, and a series of pneumatic or hydraulic actuators are used to apply the correct force at each point on the mirror to make sure the shape is maintained.

The VLT primary mirror support

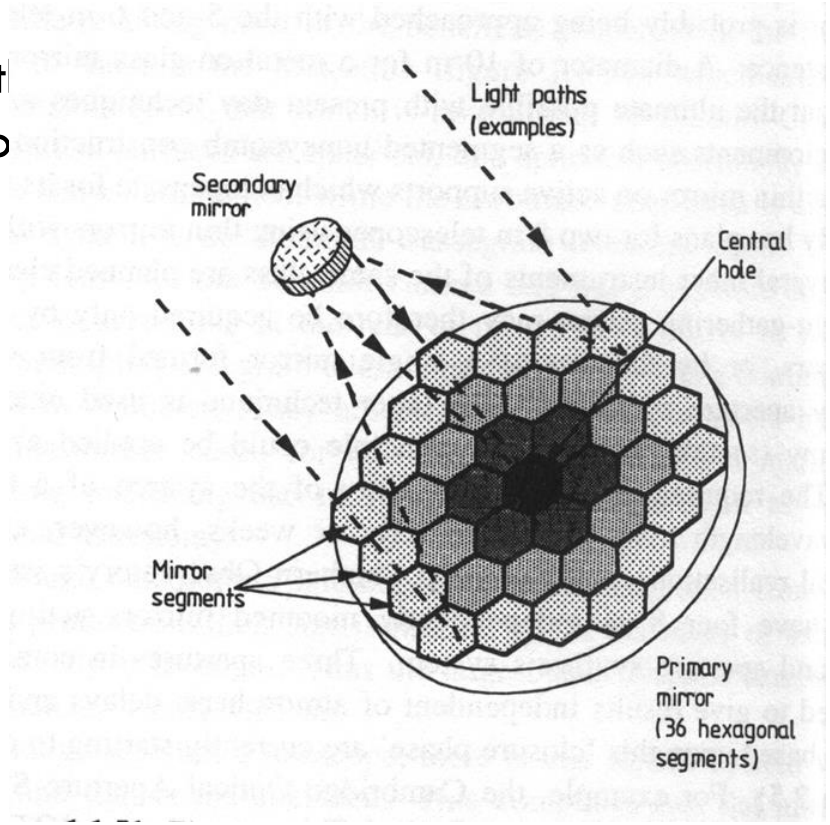
The VLT primary mirror support, showing 150 actuators arranged in six concentric rings.



Segmented mirror telescopes

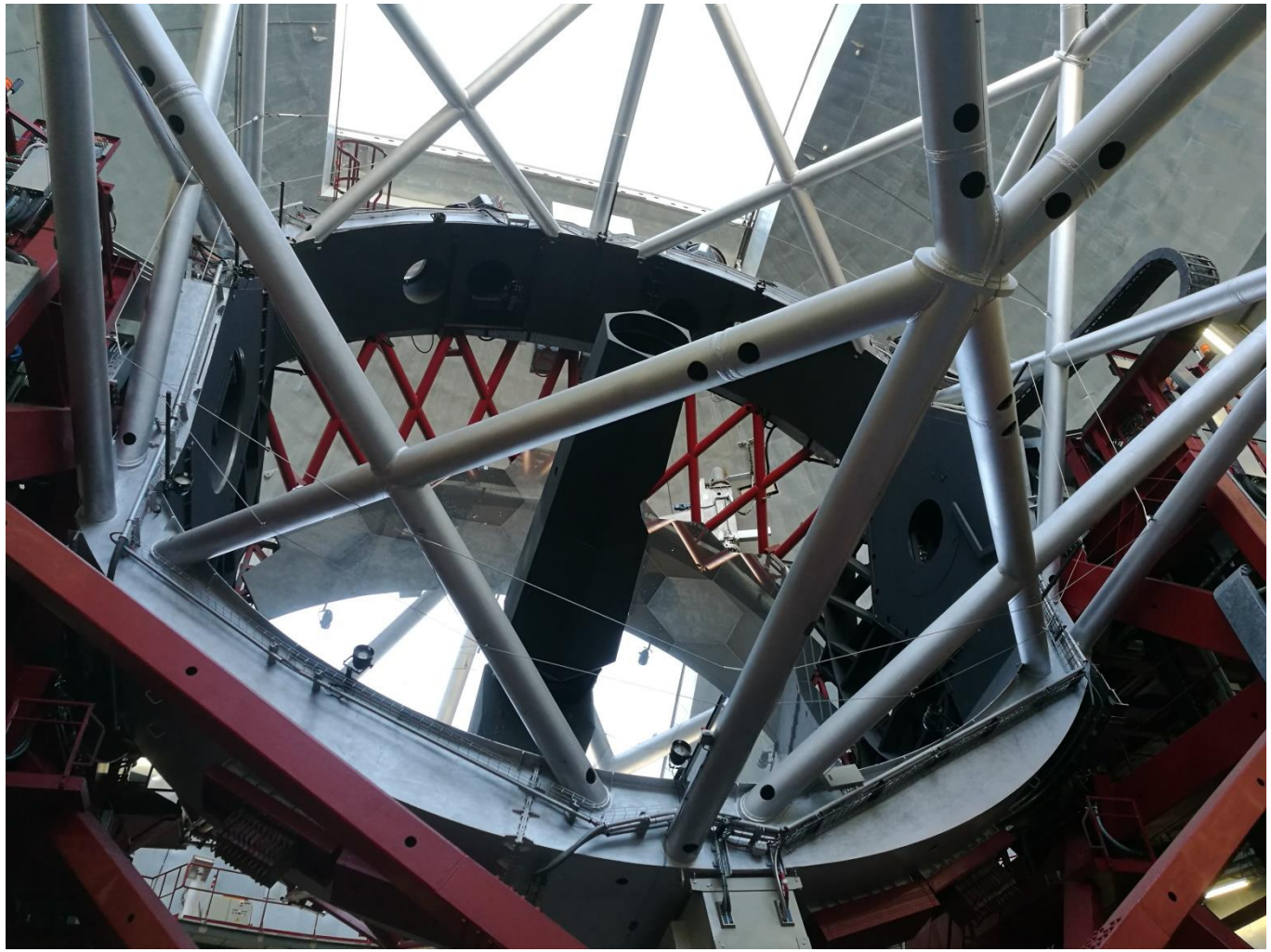
- In the Keck and GranTeCan (GTC) telescopes, instead of a single monolit mirror, the mirror is built up of hexago segments, which can be controlled accurately to bring their light to the same focus *at the same phase*. These telescopes are Ritchey–Chrétien altitude-azimuth designs.





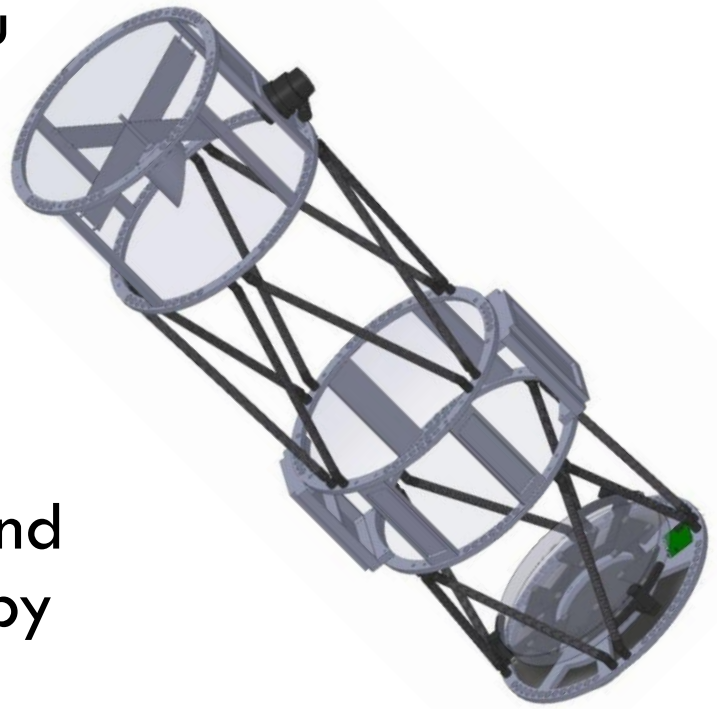
GTC telescope

Segments are kept aligned and in phase by electromechanical support and feedback system

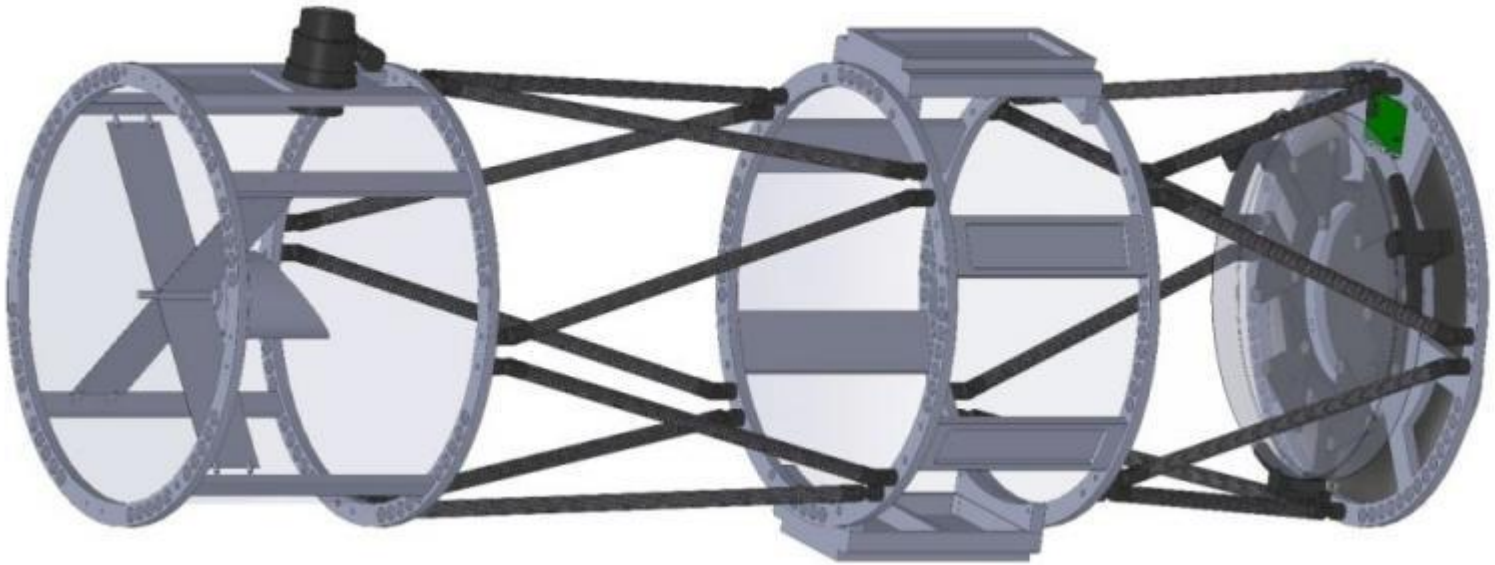


The Serrurier Truss

- It is very important to keep the optical axes of the paraboloidal and/or hyperboloidal primary and secondary mirrors precisely aligned, otherwise you get coma even on-axis.
- **The Serrurier truss** is an open "tube" structure designed so that at any orientation the flexure of primary and secondary mirrors is identical, maintaining the alignment.
- This design allows a relative rotation of primary and secondary mirrors, which can in turn be corrected by pneumatic and/or electromechanical actuators.



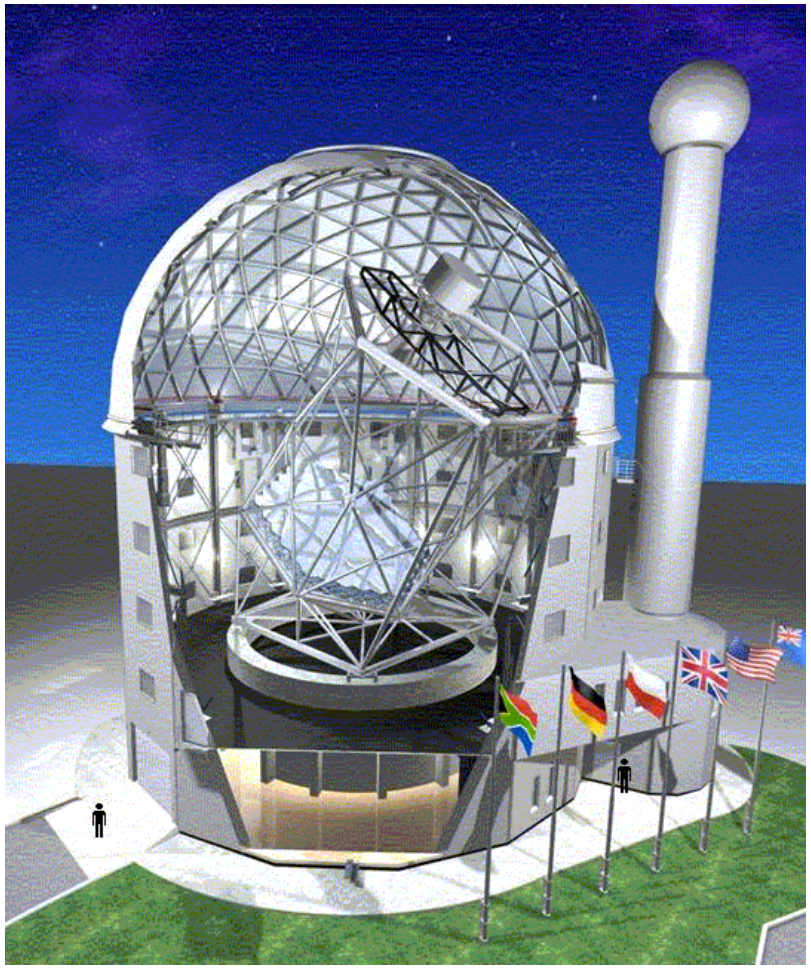
The Serrurier Truss



Southern African Large Telescope

- The segmented 11 by 9.8 meters mirror is spherical.
- SALT has a fixed zenith angle of 37 degrees, optimised for the Magellanic clouds.
- Azimuth only tracking.
- Prime focus top end moves Arecibo style.
- Limited sky area and tracking time.
- Much cheaper than telescopes of similar size but of classical design.





New Large Telescopes

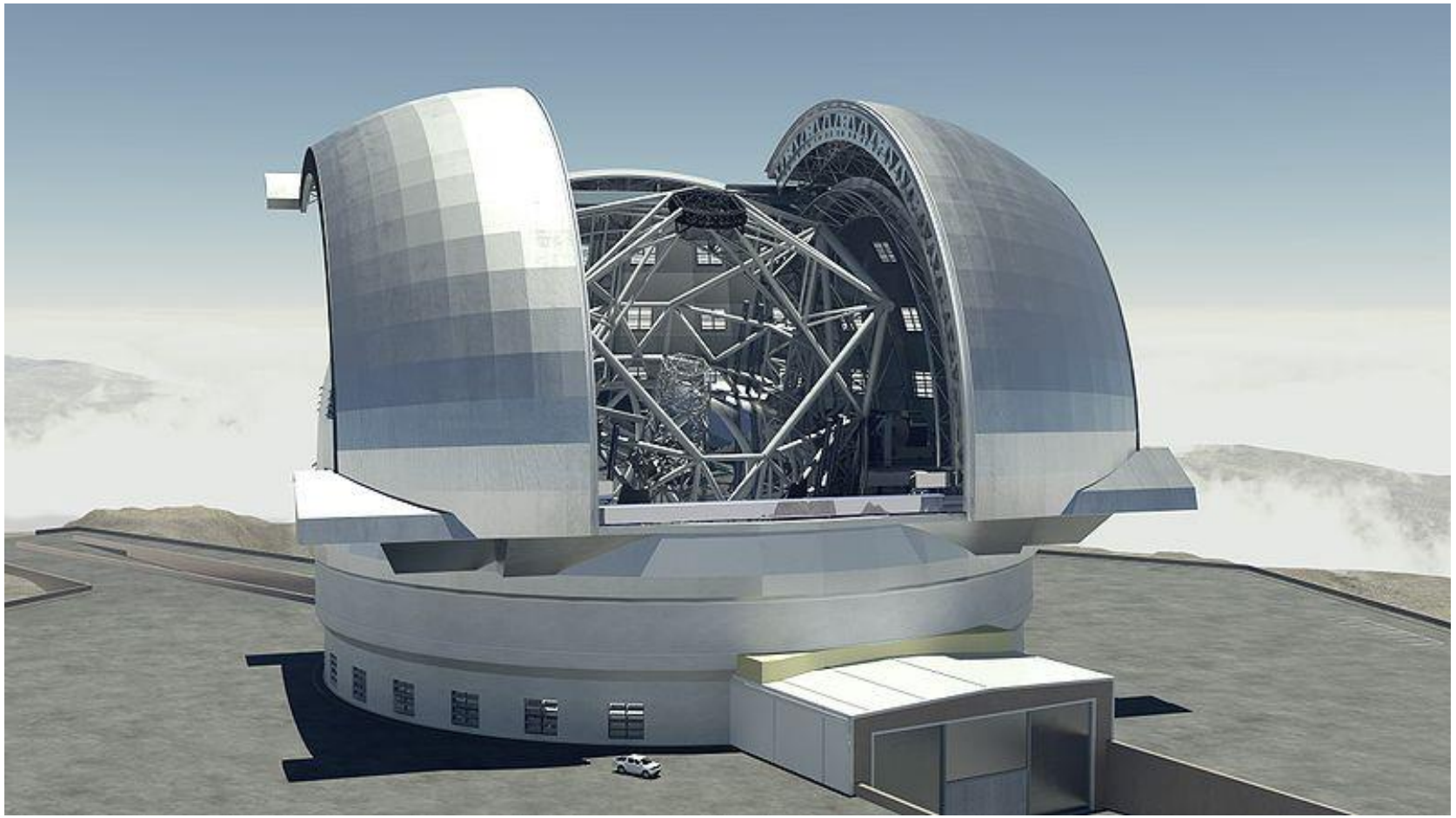
- Current state of the art in optical/near infra-red telescopes is:
  - 8.2 metre monilithic meniscus mirror (VLT)
  - 10 metre equivalent mirror with hexagonal segments (GTC).
- Various proposals for larger 20-100 metre aperture telescopes.
- The European Extremely Large Telescope (ELT) is already under construction.
- International 30 metre telescope (TMT) ~~is preparing to start construction.~~
  - In June 2025 the United States' National Science Foundation dropped support for the TMT in favor of the the Giant Magellan Telescope (GMT).

Problems of Extremely large telescopes

- Need light stiff structures, composite materials, e.g. Silicon Carbide for mirror segments.
- Large downward looking secondary is a huge problem, carbon fibre structures.
- Control loop to keep segments aligned is heavily nested hierarchical control at hundreds of Hz, beyond the scope of current hardware.
- Wind distortion a severe problem.
- Adaptive mirrors up to a metre across may be required.
- Need Multi-Conjugate Adaptive optics with laser guide stars (see next lectures).
- Huge data rates.

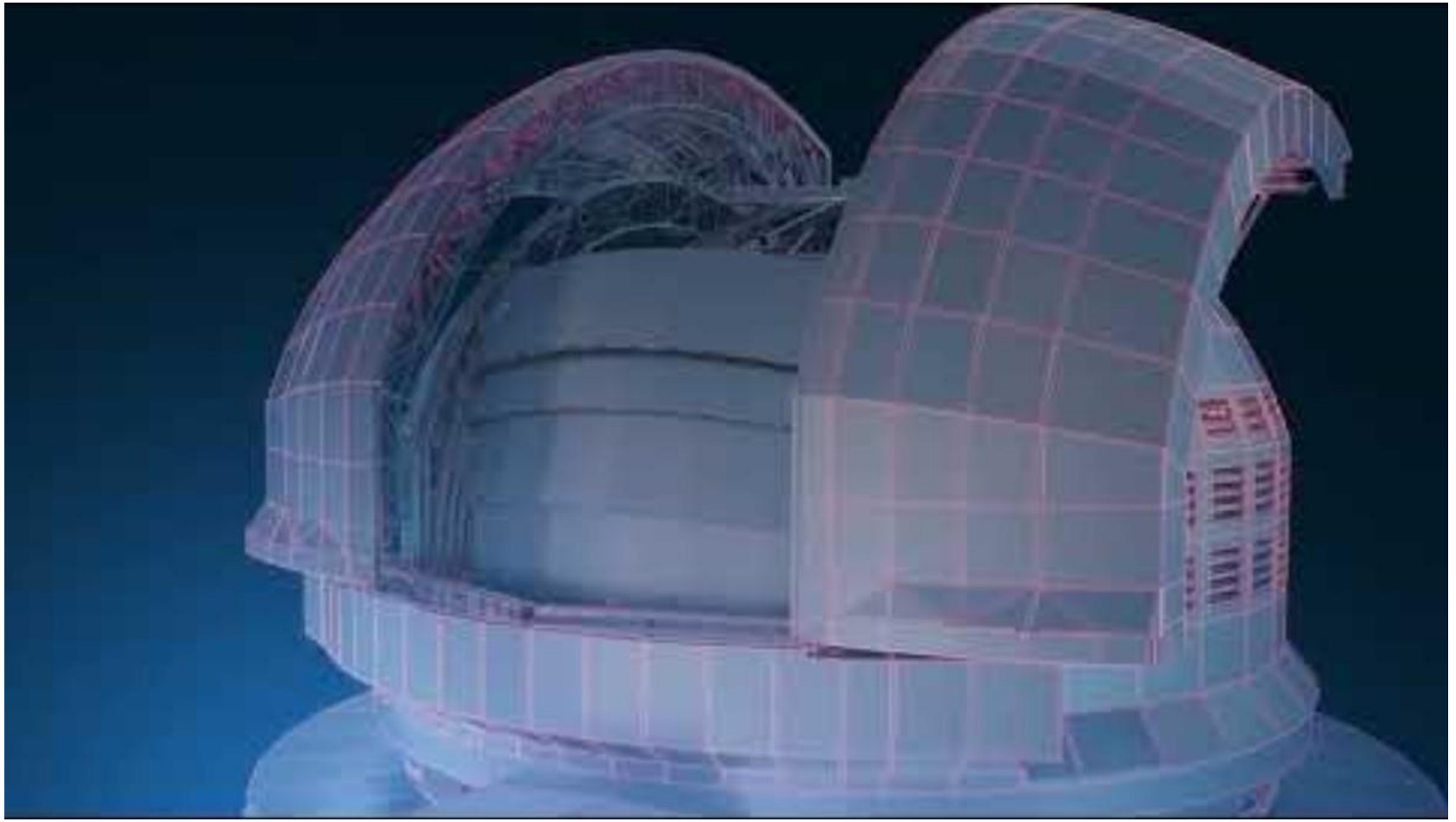
The European Extremely Large Telescope

- Diameter: 39 m (798 hexagonal 1.4 m mirror segments).



The European Extremely Large Telescope

- Diameter: 39 m (798 hexagonal 1.4 m mirror segments).



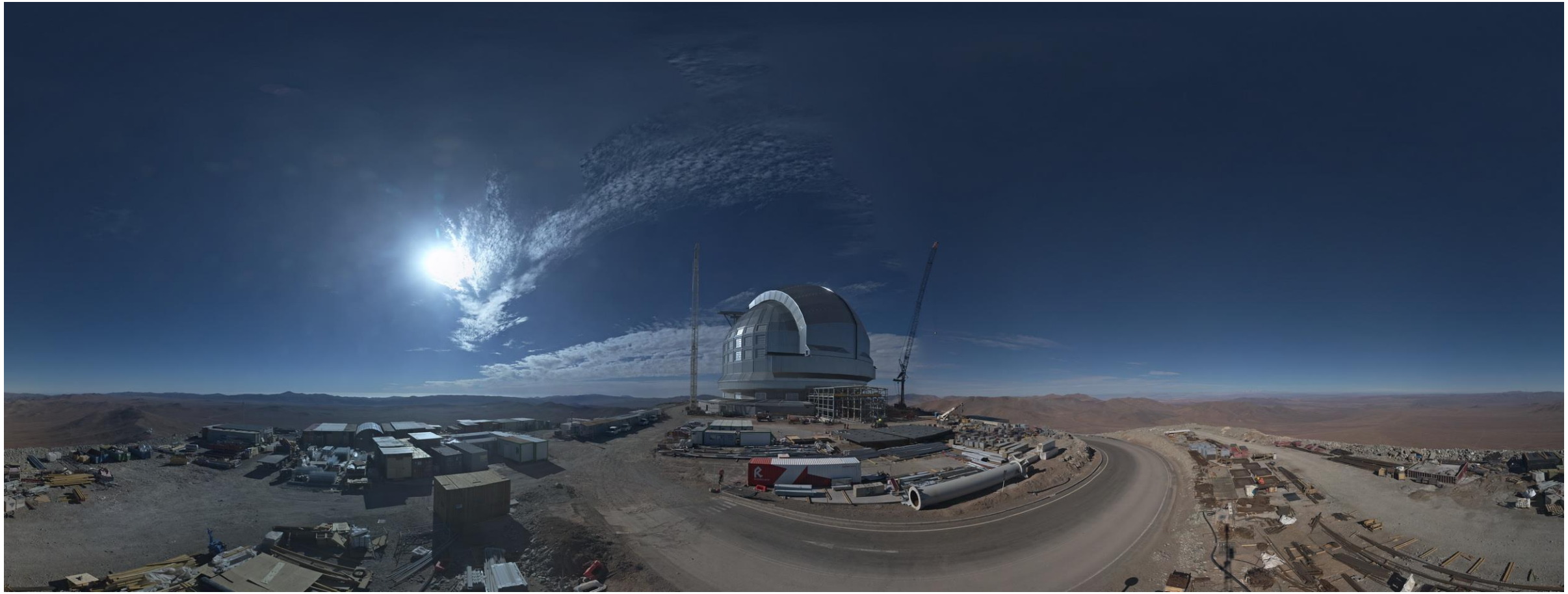
The European Extremely Large Telescope

- Previous year's lecture: 6 September 2025, 21:00 CEST
- 2028 (planned): Technical first light of the ELT



The European Extremely Large Telescope

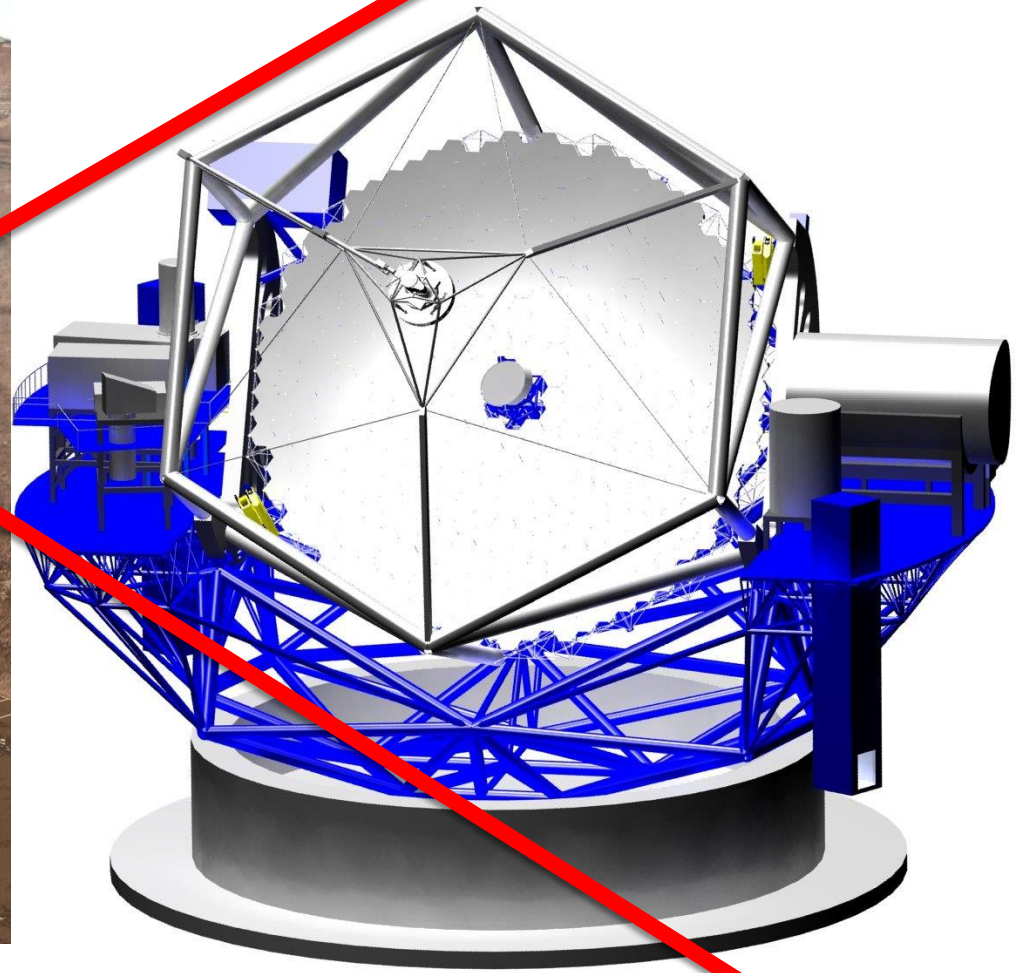
- Current state (4 September 2026, 22:00 CEST)
- 2028 (planned): Technical first light of the ELT





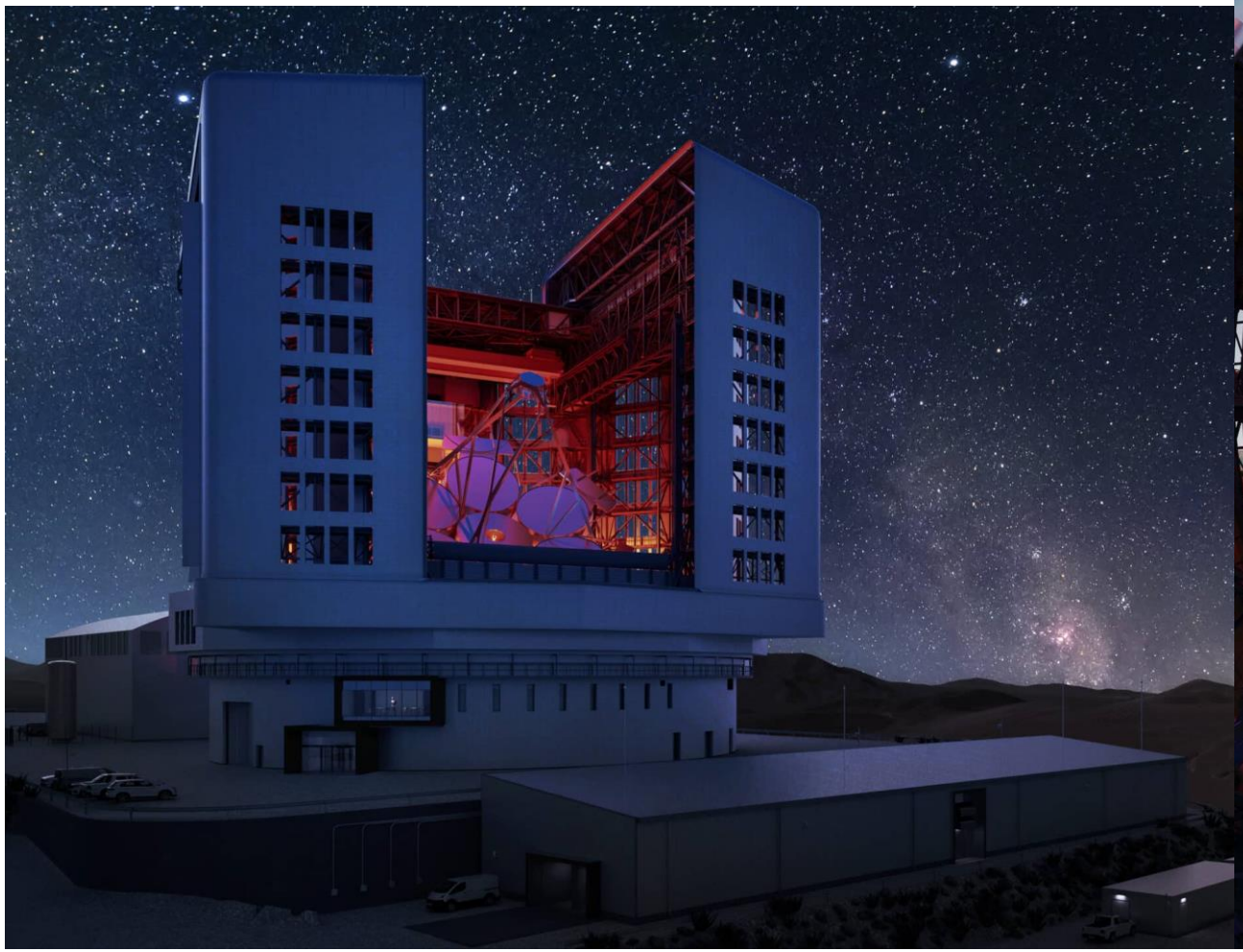
International 30 metre telescope (TMT)

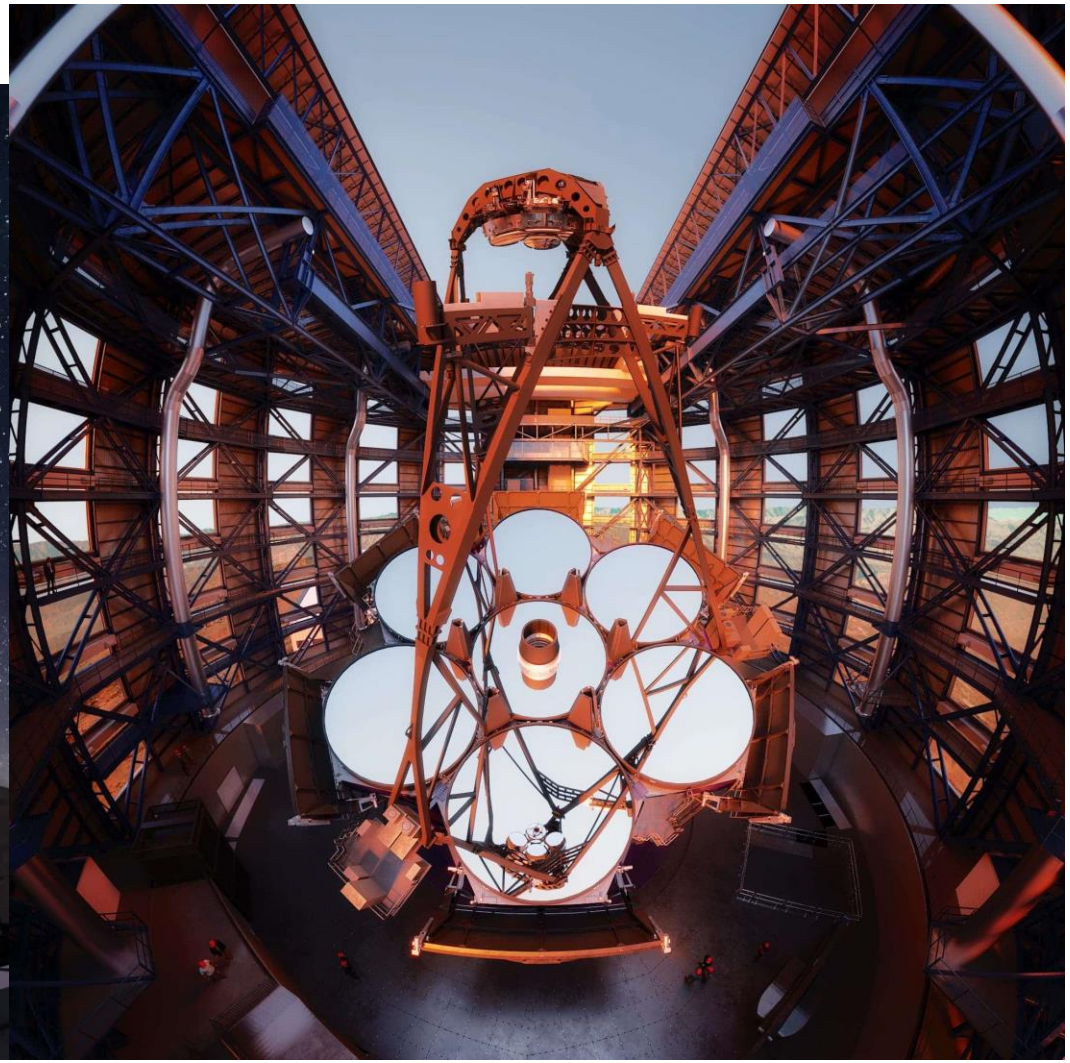




Giant Magellan Telescope (GMT)

7 x 8.4-m mirrors → a 25.4-m surface





Telescopes for high-energy astrophysics

The high-energy spectrum

- Extreme ultraviolet (EUV): 100 – 1000 Å (12-120 eV)
- Soft X-rays: 10 – 100 Å (120 – 1200 eV)
- X-rays: 0.1 – 10 Å (1.2 – 120 keV)
- Soft γ-rays: 0.01 – 0.1 Å (120 – 1200 keV)
- γ-rays: <0.01 Å (>1.2 MeV)

$$\lambda\,[\text{Å}] = 12.40/E\,[\text{keV}]$$

The high-energy spectrum

The main production mechanisms:

- Synchrotron radiation
- The inverse Compton effect
- Free-Free (Bremsstrahlung) radiation

The flux of the radiation varies enormously with wavelength:

- The solar optical flux – 10^{21} photons m^{-2} s^{-1}
- The solar flux at 10 Å – 5×10^{9} photons m^{-2} s^{-1}
- The total flux from all sources for energies >1GeV – few photons m^{-2} **day**$^{-1}$

UV, X-ray, and γ-ray Telescopes

- All radiation at $\lambda < 3100$ Å is absorbed by the atmosphere.
- UV (2000 – 3000 Å) and hard X-rays can penetrate to an altitude of 30 – 40 km while soft X-rays only reach higher altitude:
  - therefore, all UV and X-ray observations have to be carried out from rockets or from space.
- Ultraviolet telescopes are similar to optical in design, but must be in space. The Hubble Space Telescope is the premier ultra-violet telescope.

X-ray telescopes

- Imaging of high-energy photons is a much more difficult task because of their extremely penetrating nature.
- Normal designs of telescope are impossible:
  - in a refractor the photon would be absorbed, scattered or unaffected rather than refracted by the lens;
  - in a reflector the photons would just pass through any material they impinge on at normal incidence.

X-ray telescopes

- X-rays at normal incidence are not reflected by anything, however, at energies <~100 keV, photons may be reflected with up to 50% efficiency off metal surfaces, when their angle of incidence approaches 90° – grazing incidence.
- Approximate empirical formula for the critical angle is $\theta_C = 69 \times \sqrt{\rho/E}$ arcminutes
  where ρ is density of material in g/cm^3 and E is energy of X-rays in keV.

X-ray telescopes

- For e.g. nickel or gold, critical angle is of order few degrees for 1 keV X-rays, $<0.6^{o}$ for 10 keV X-rays, and $<0.1^{o}$ for 100 keV X-rays.

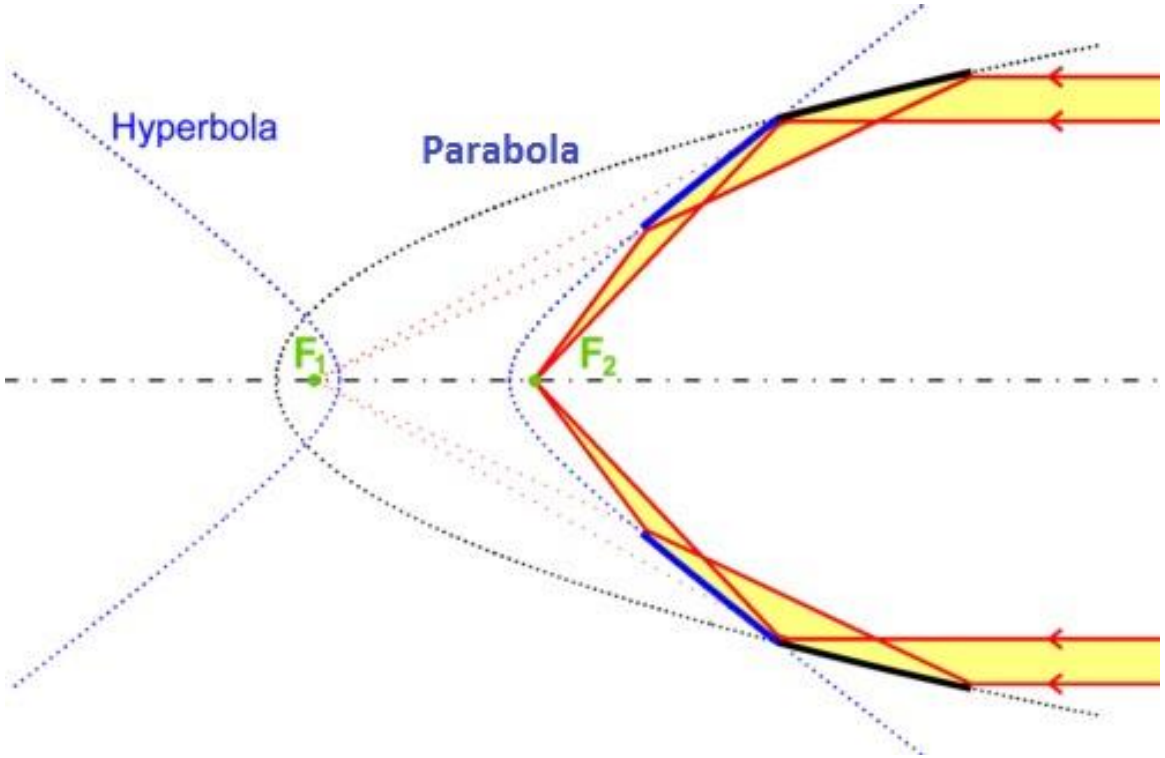
- However, at energies under a few keV, forms of reflecting telescope can be made. Imaging X-ray optical systems were introduced by **Hans Wolter**.
- Its not an easy technology though, there are some difficult problems.

Problems with grazing incidence telescopes

- A telescope with a single grazing incidence paraboloid suffers from severe astigmatism away from the optical axis.
- The aberrations can be corrected by using a pair of confocal surfaces of rotation (i.e. surfaces with a common focus), one a paraboloid and the other a hyperboloid or ellipsoid.
- Surfaces must be extremely accurate for grazing incidence mirrors to work efficiently (remember $\lambda/20$).

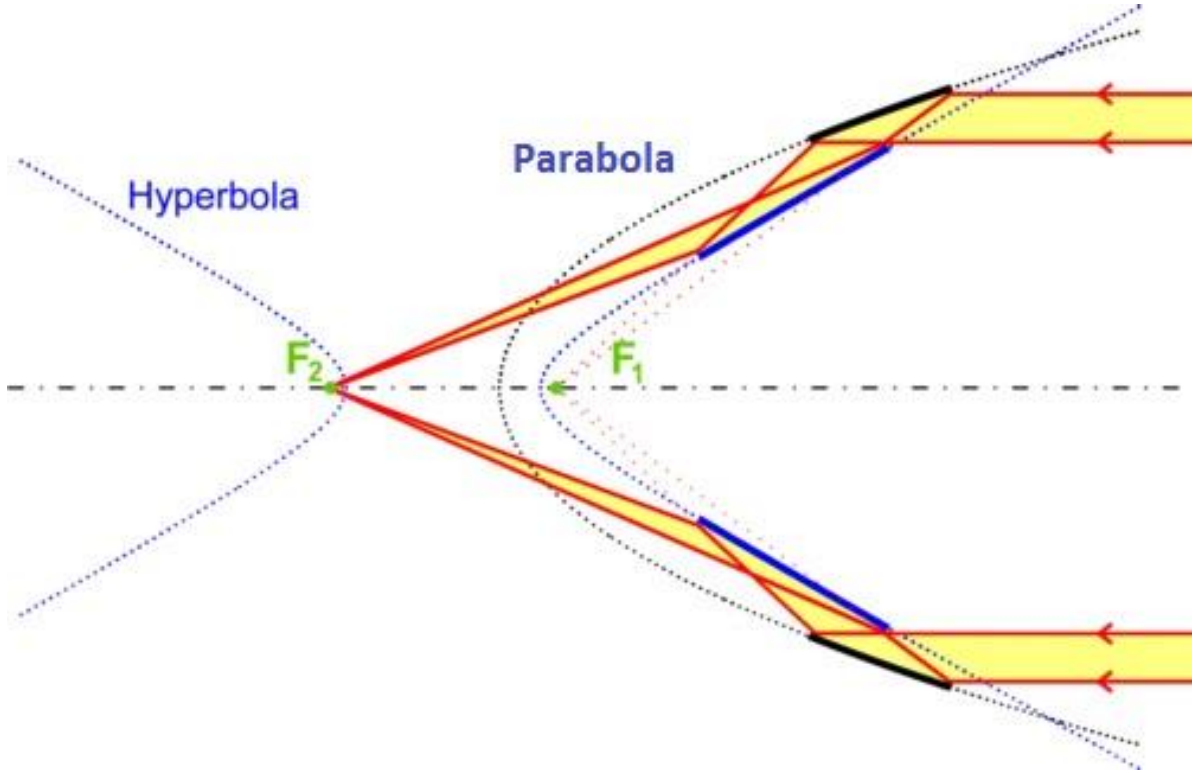
Wolter type I telescope

- Wolter Type I telescope – has a grazing incidence paraboloid followed by a grazing incidence concave hyperboloid.



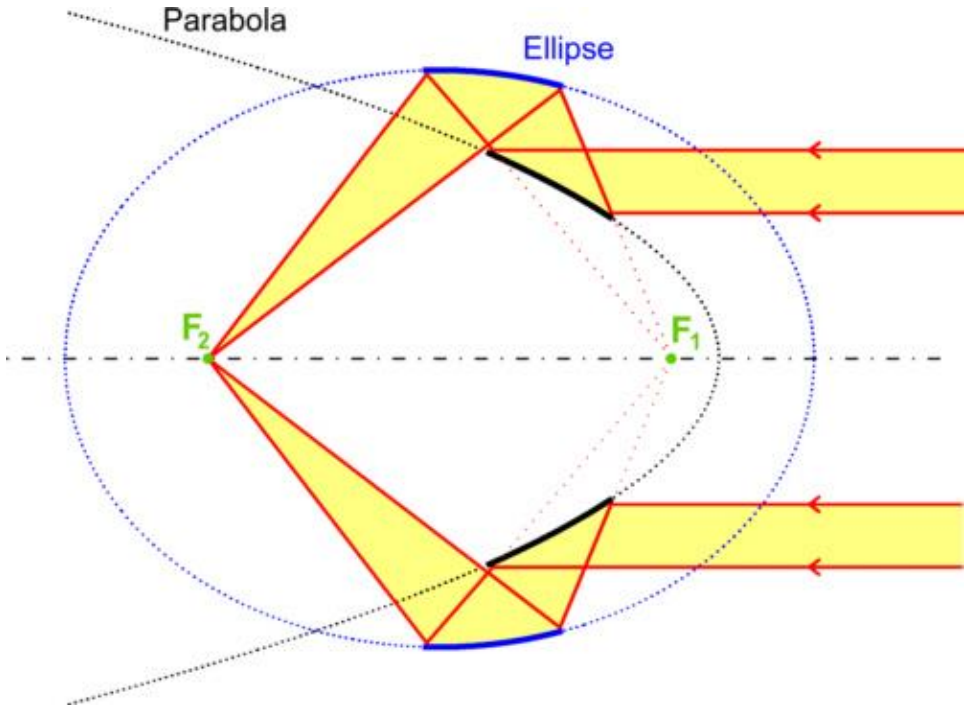
Wolter type II telescope

- Wolter Type II telescope - uses a convex secondary hyperboloid and has a longer focal length (allowing the detectors to be physically larger and easier to make).



Wolter type III telescope

- Wolter Type III - telescope uses a convex paraboloid and a concave ellipsoid, gives an even longer focal length.



Wolter telescopes (summary)

- Wolter Type I telescope – has a grazing incidence paraboloid followed by a grazing incidence concave hyperbolid.
- Wolter Type II telescope - uses a convex secondary hyperboloid instead of a concave one and has a longer focal length (allowing the detectors to be physically larger and easier to make).
- Wolter Type III - telescope uses a convex paraboloid and a concave ellipsoid, gives an even longer focal length.

Wolter telescopes

- Wolter telescopes are **two mirror** telescopes, somewhat analogous to the Cassegrain and Ritchey-Chrétien optical systems in use at longer wavelengths.

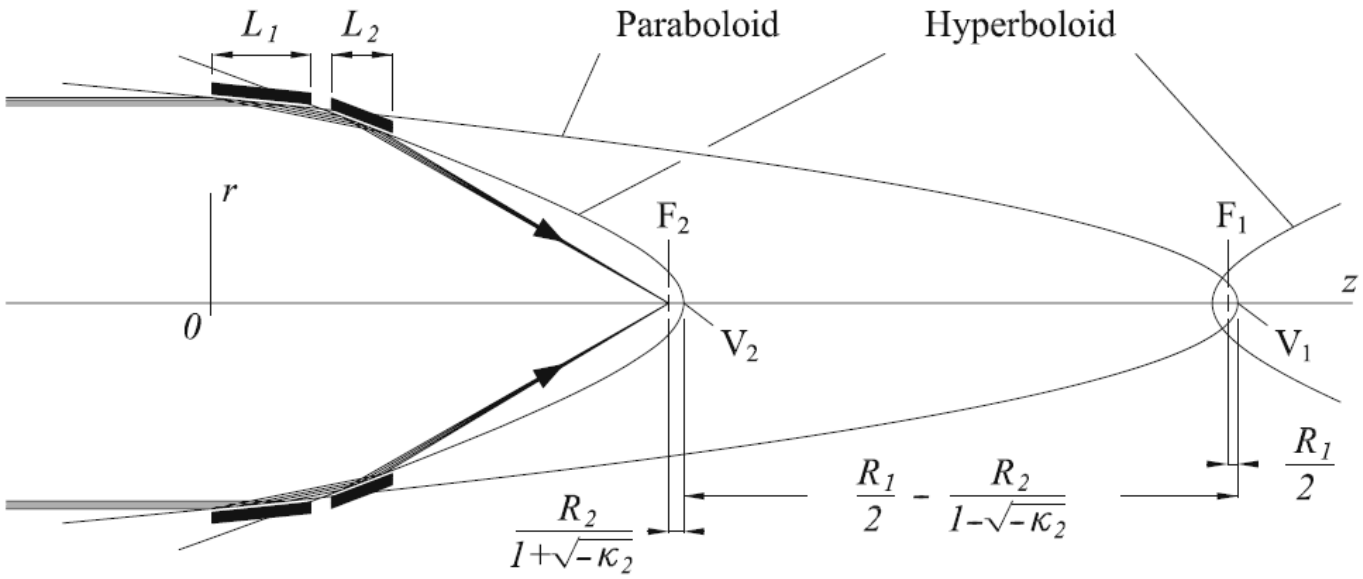
- The **Wolter Type I** is the shortest of these long systems and therefore has been extensively utilized as a *telescope* in X-ray astronomy.

General Imaging Properties

Small grazing angles result in larger focal length:

$d/f = 2 \times \sin 4\alpha$, where d is aperture size, f – focal length, α – the slope angle of the 1st mirror element (approximately the grazing angle for paraxial rays).

$\boldsymbol{d/f}$ **is typically around 1/10** for the energy band up to 10 keV



General Imaging Properties

The field-of-view: Wolter optics with a lower ratio d/f have smaller fields-of-view than those with larger ratios.

For example:

ROSAT – soft X-ray telescope ($d/f \sim 1/3$) – had field of view diameter of 2°.

XMM-Newton has the CCD cameras which cover about 30′.

General Imaging Properties

Angular Resolution: The limit of the resolution is due to **surface irregularities** in the mirrors, rather the diffraction limit of the system.

The lower limit of the irregularities is about 3-4 Å in size for the very best of the current production techniques – comparable with the wavelength of photons of about 1 keV.

Chandra has the resolution of <1", XMM-Newton – 16", Swift – 18".

For the resolution of around 1' the mirror shapes can be simple cones – fabrication costs are much reduced.

General Imaging Properties

Collecting Area: For small slope angles α the geometrical collecting area of a mirror shell is a thin circle with a projected area S of

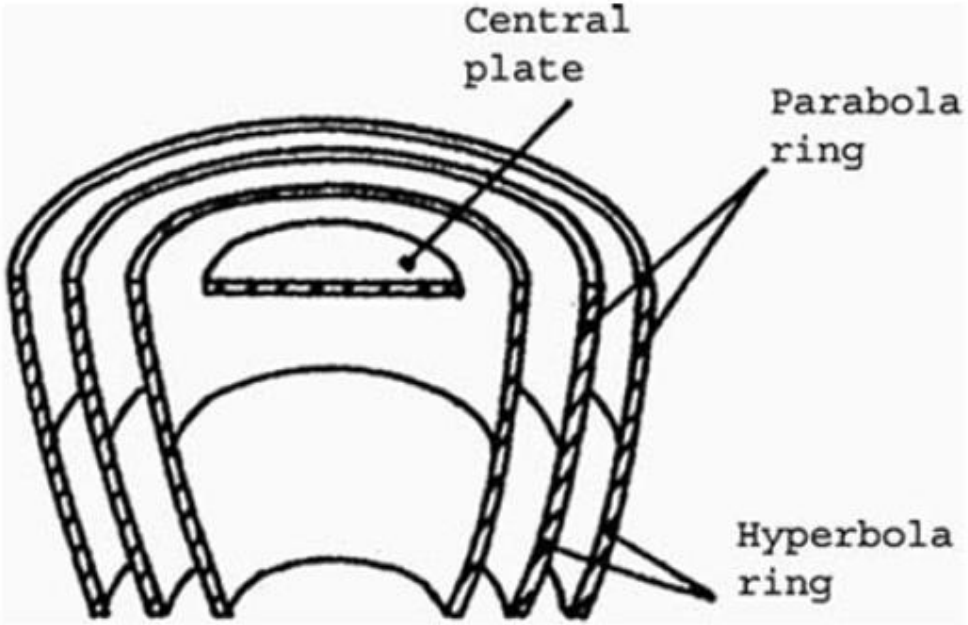
$$S \approx 2\pi r \times l \sin \alpha$$

where r is the mirror radius and l the length of a mirror element.

This is much less than the polished mirror surface.

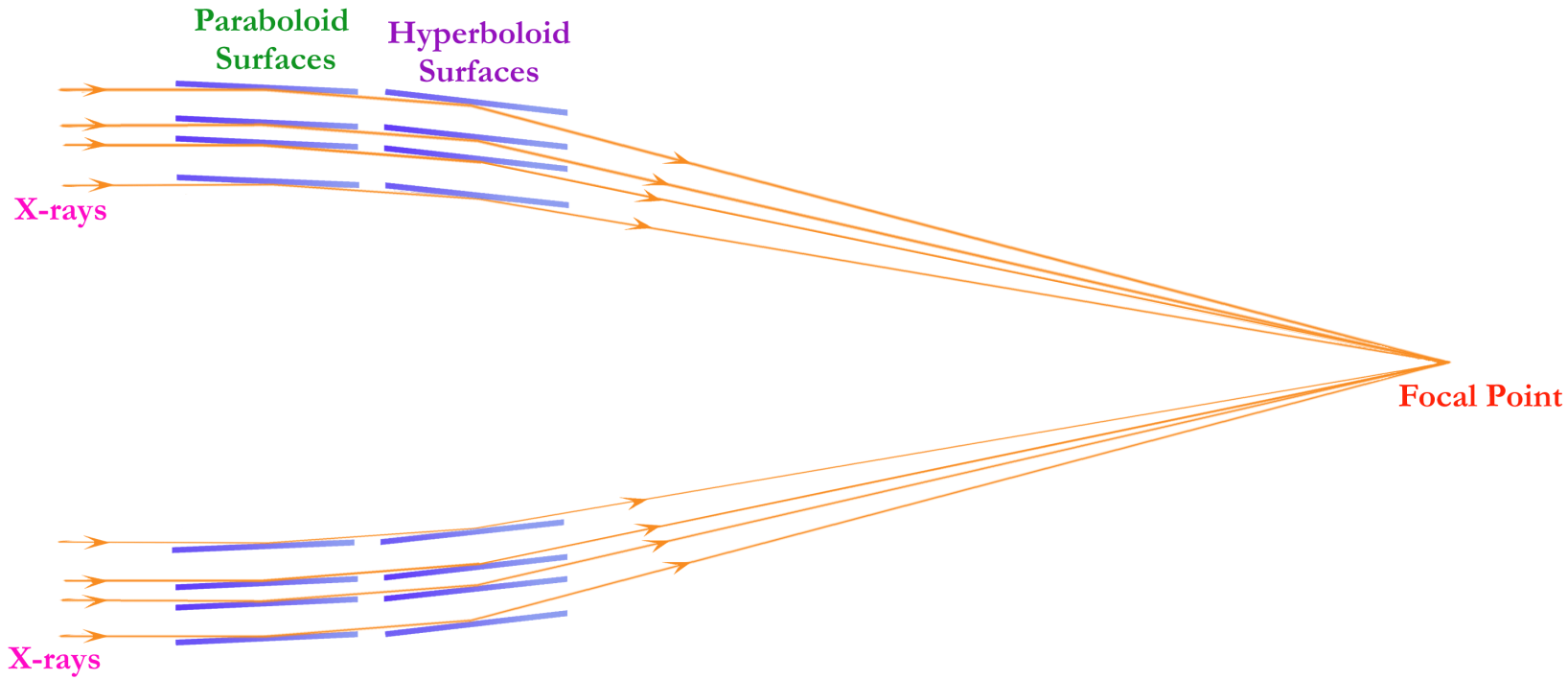
Nested grazing incidence telescopes

- Effective area can be increased by nesting grazing incidence telescopes one inside the other.

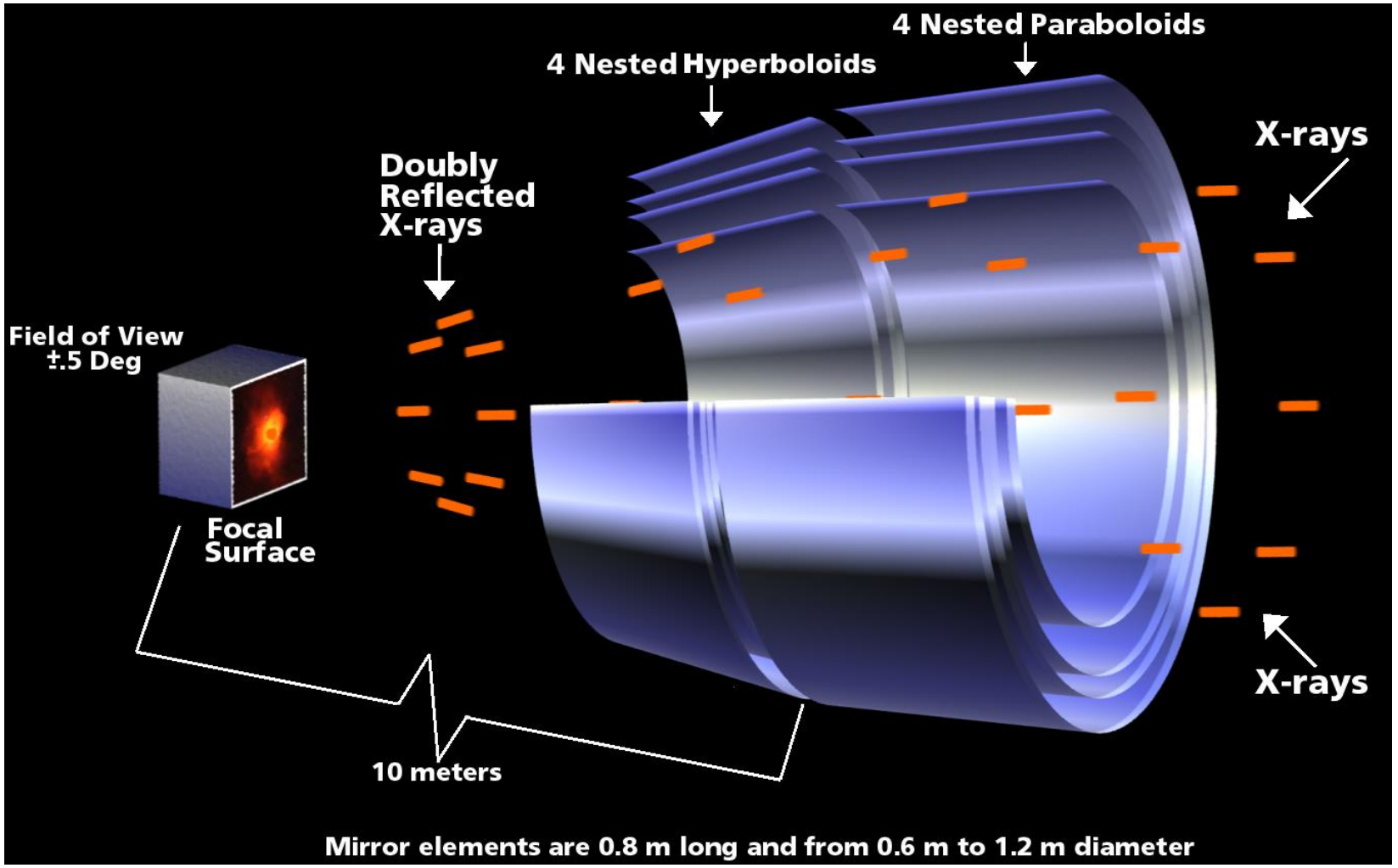


Chandra X-ray telescope

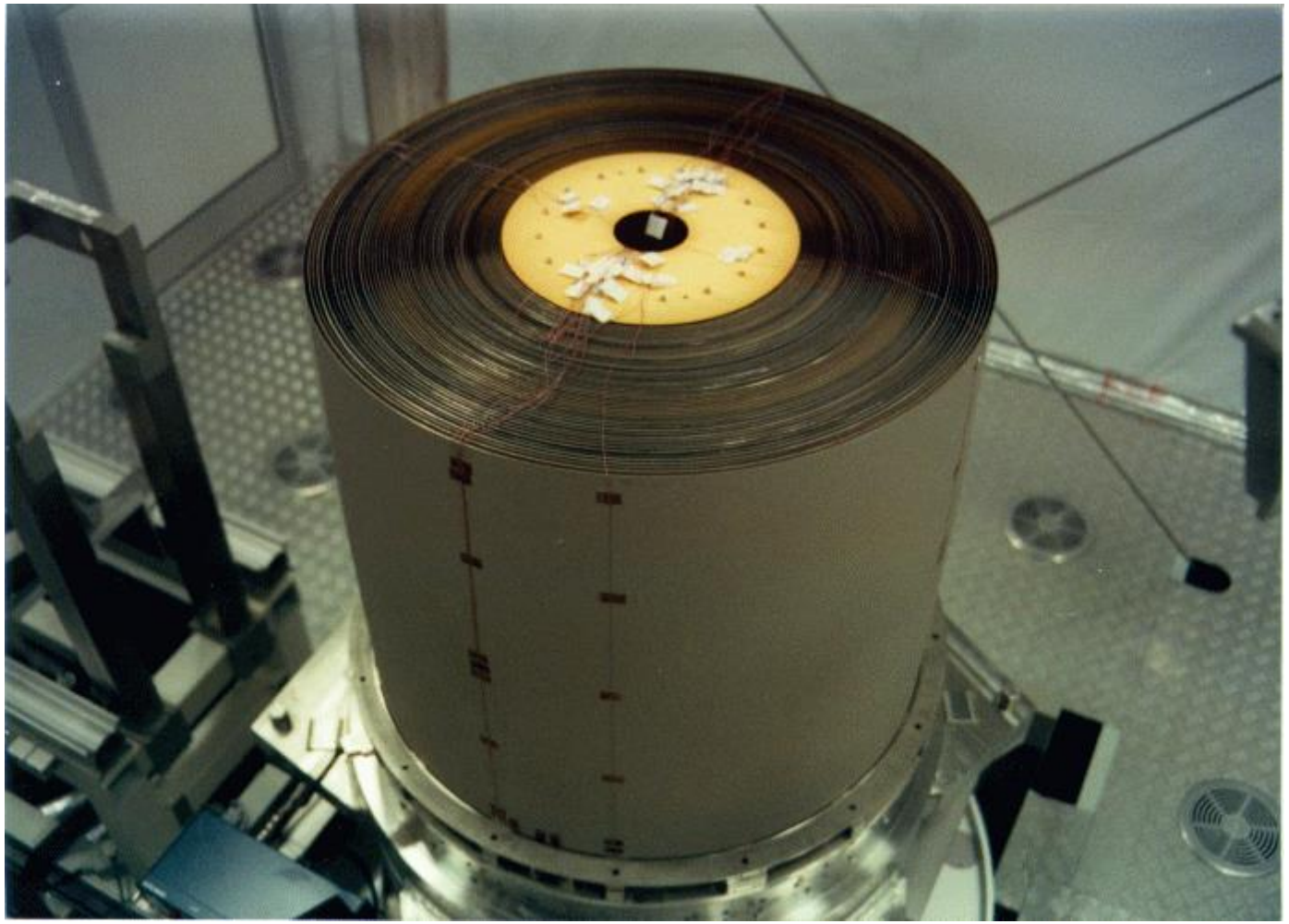
- The basic optical design of Chandra



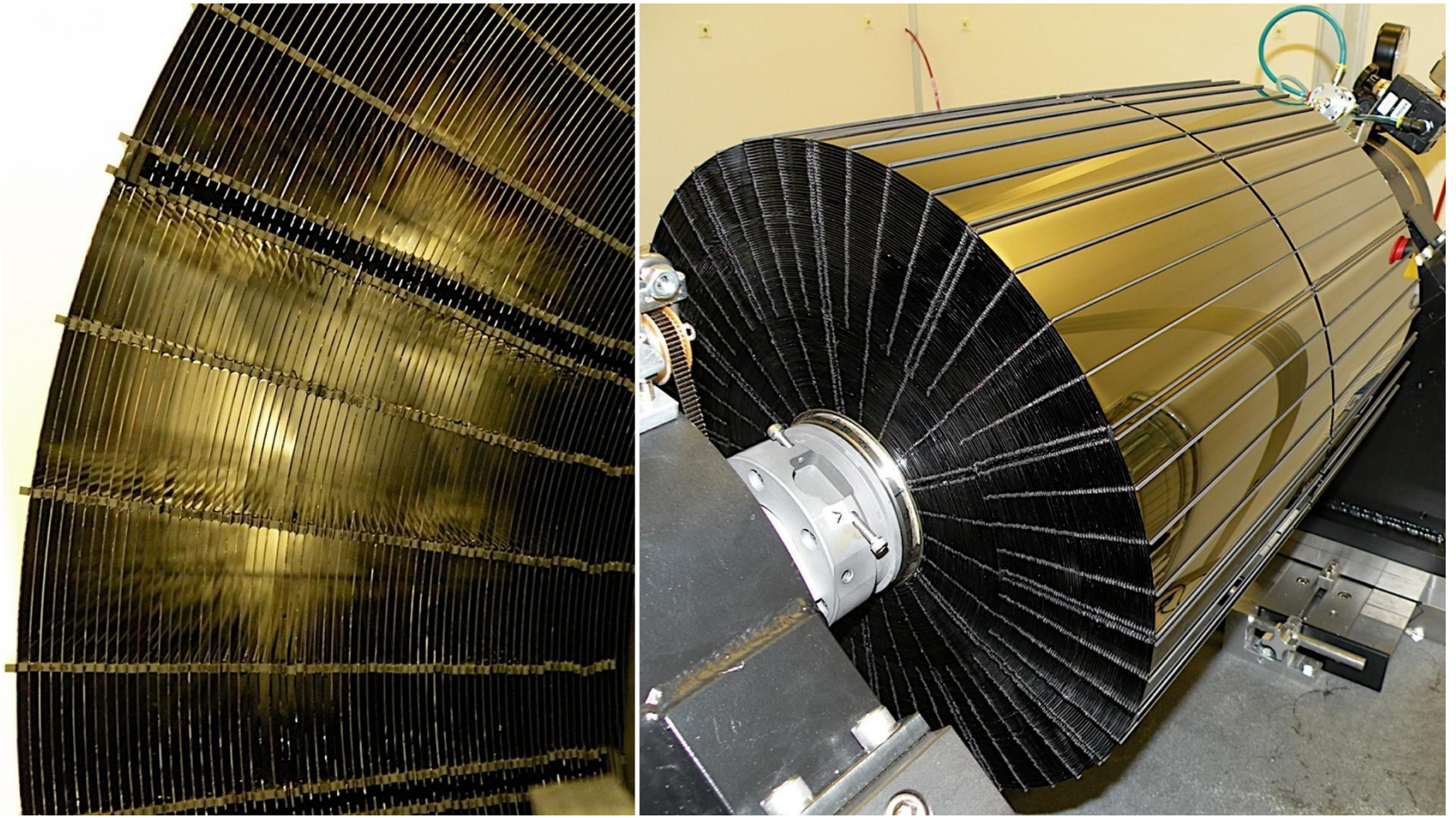
Chandra X-ray telescope



XMM-Newton 58 nested mirrors



The Nuclear Spectroscopic Telescope Array (NuSTAR)



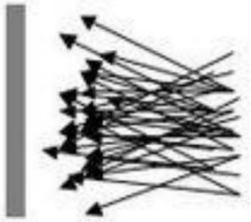
133 nested mirrors (a conical approximation)

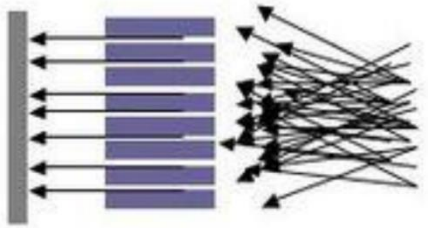
Satellite missions with Wolter telescopes

Mission	Year of launch	Upper Energy limit (keV)	Focal length (m)	Mirror modules	Degree of nesting	Effective area @ 1 keV (cm^2)	On-axis resolution
~~ROSAT~~	~~1990~~	~~0.3-2.5~~	~~2.4~~	~~1~~	~~4~~	~~420~~	~~3”~~
Chandra	1999	10	10.0	1	4	780	<1”
XMM-Newton	1999	15	7.5	3	58	4260	16”
Swift	2004	10	3.5	1	12	130	18”
~~Suzaku~~	~~2005~~	~~12~~	~~4.75~~	~~4~~	~~175~~	~~2250~~	~~120”~~
NuSTAR	2012	3-79	10.15	2	133	847 (9 keV)	45”
AstroSat	2015	8	2	1	41	200 (1.5 keV)	120”
Spektr-RG / eROSITA ART-XC	2019	10 5-30	1.6 2.7	7 7	54 28	2400 450 (8 keV)	16” 45”

High energy imaging by collimation

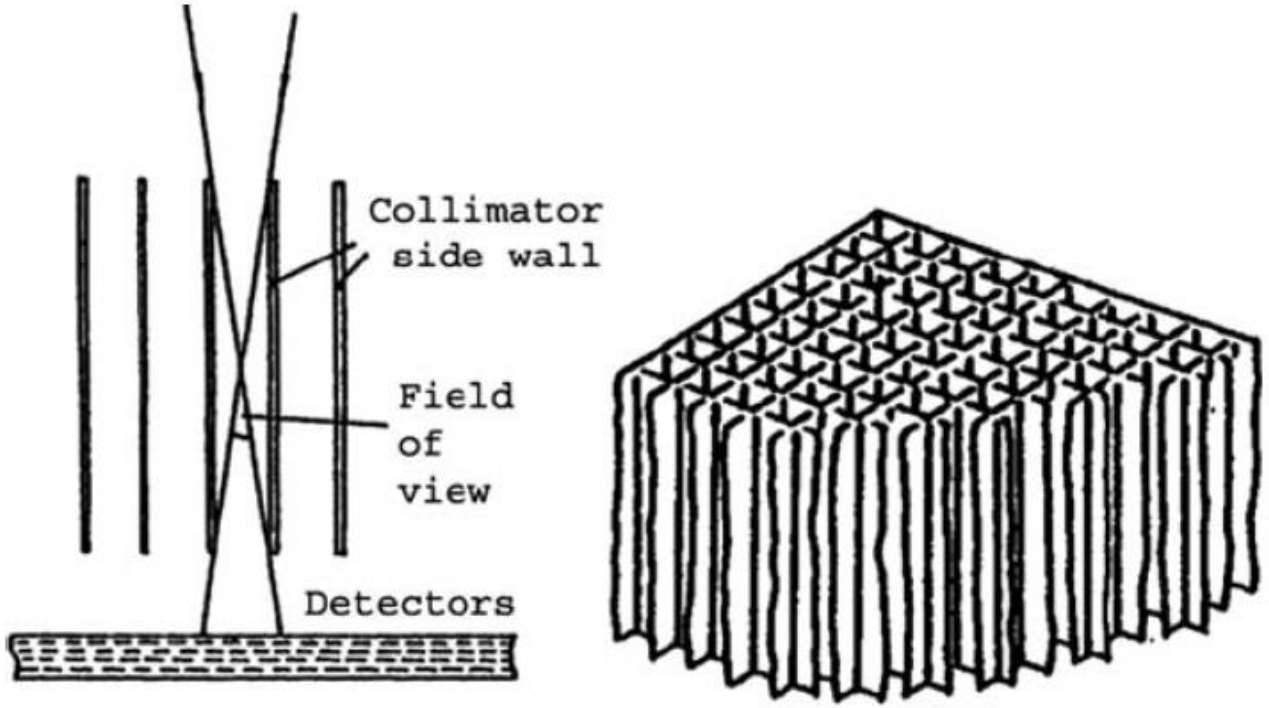
- High energy x-rays and gamma rays cannot be focussed, but some degree of directionality is possible by using a collimator, which restricts the angle of acceptance.





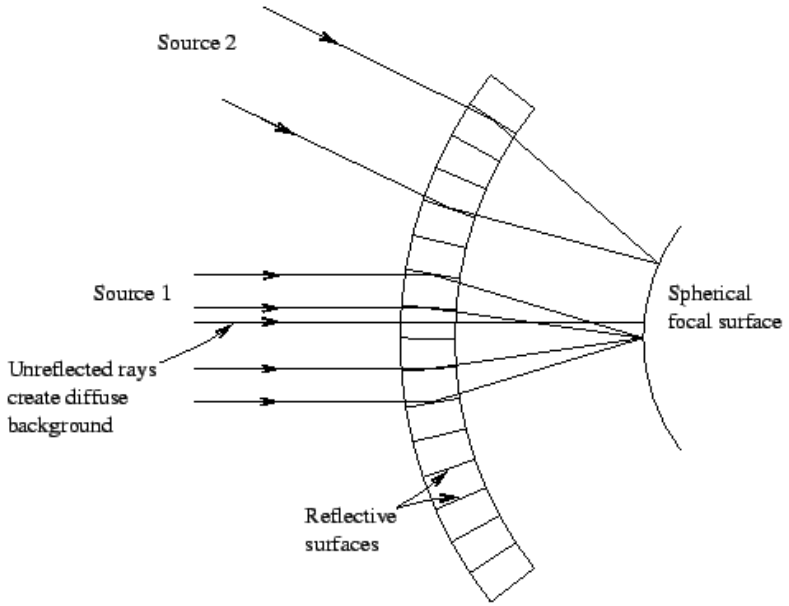
High energy imaging by collimation

- Simplest collimator design is a honeycomb collimator which is a closely packed array of tubes.

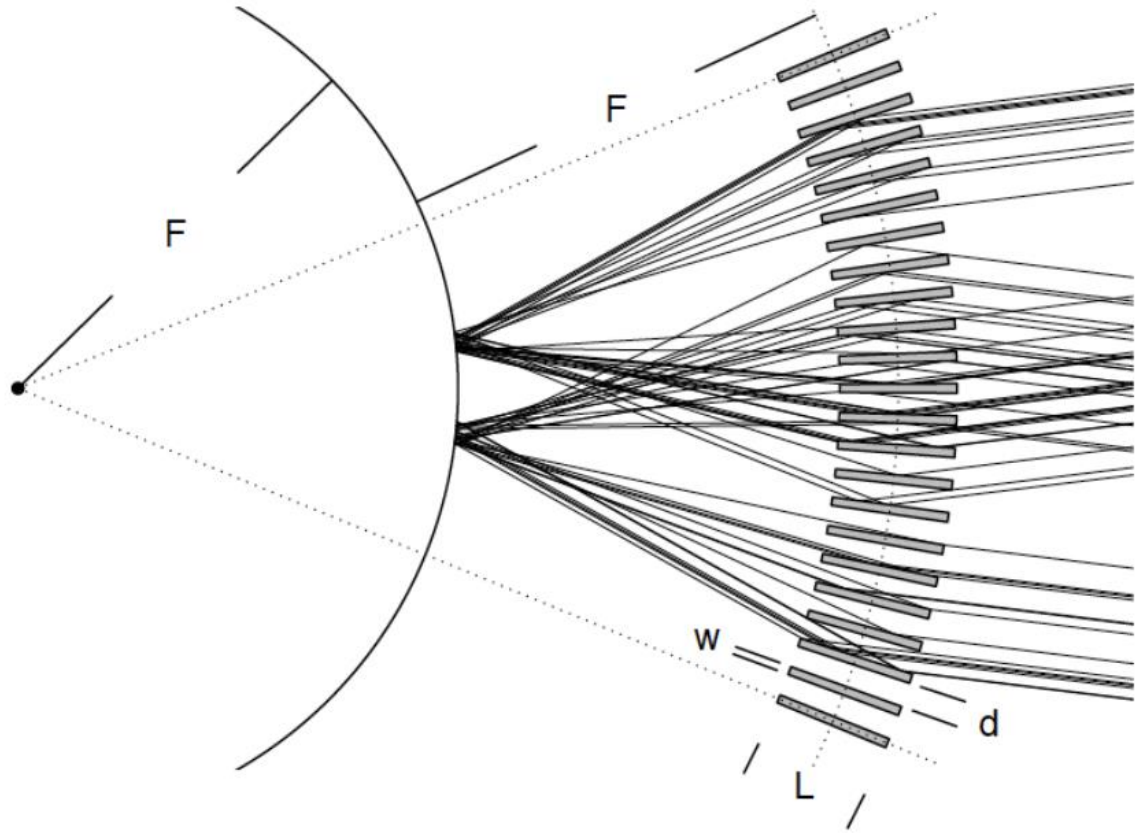


Lobster Eye imaging collimator

- Lobster Eye imaging collimator is essentially a honeycomb collimator curved into a portion of a sphere and with a position sensitive detector also curved onto a spherical surface.
- Lobster eye collimator uses grazing incidence reflection, rays which pass straight through form a background.
- Gives moderate quality imaging over a wide field.



Lobster Eye imaging collimator



The Lobster-eye X-ray Telescope (LXT) is an approved NASA mission for the study of diffuse sources. LXT combines very large field of view with good angular and energy resolution.

Laue Diffraction

- X- and gamma-rays are diffracted by certain crystals, and give rise to a pattern of spots for a normal incidence beam.
- For some crystals (e.g. germanium) the number of spots is quite small.

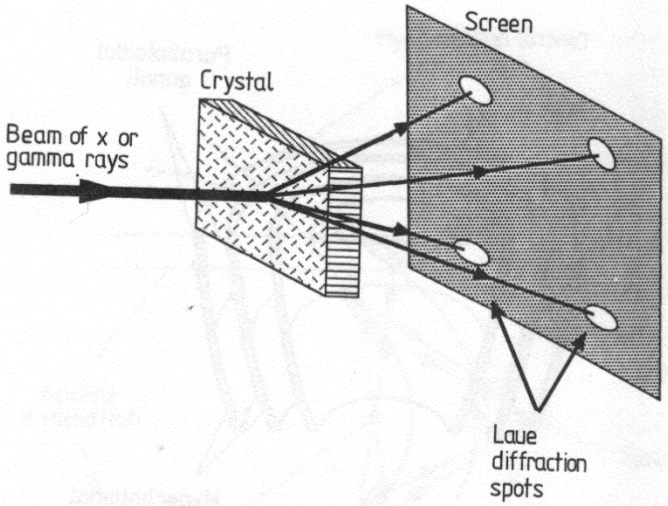
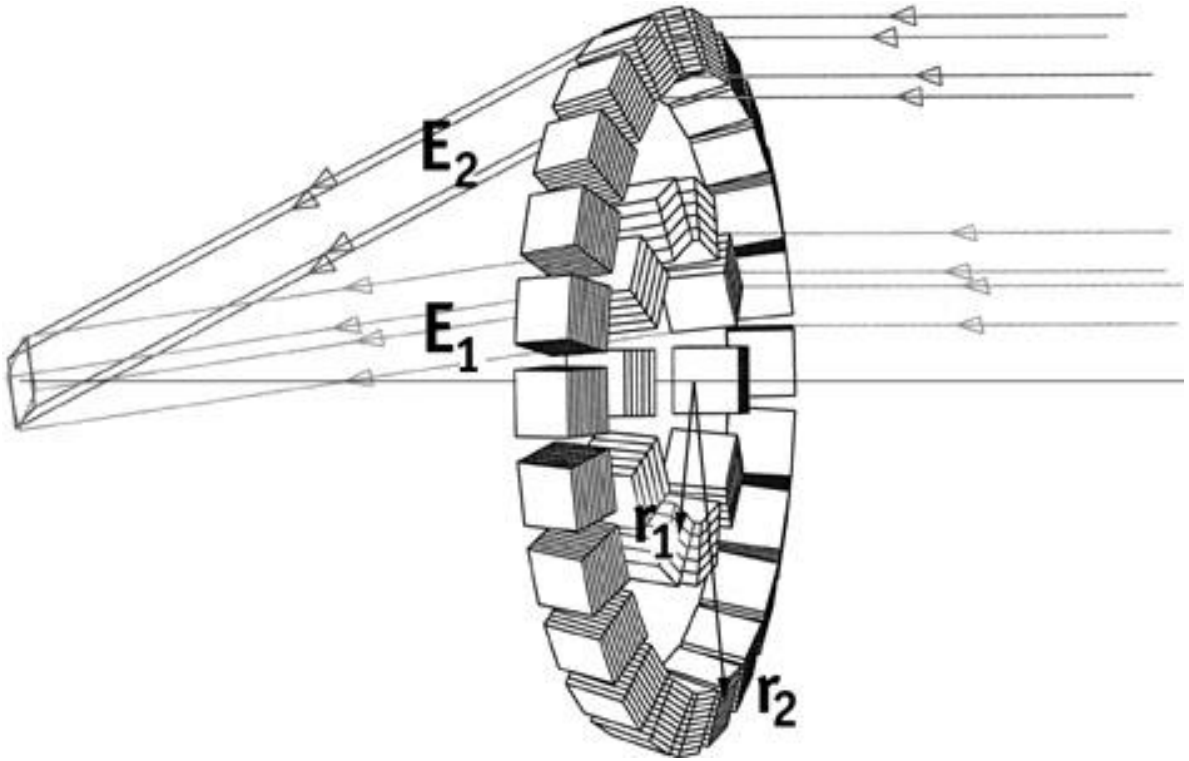


Figure 1.3.12. Laue diffraction of x- and gamma rays.

Laue Diffraction telescope



Crystals are oriented such that the spots from crystals in a ring fall on the same part of the detector.

(Still not in use, only experiments!)



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 4

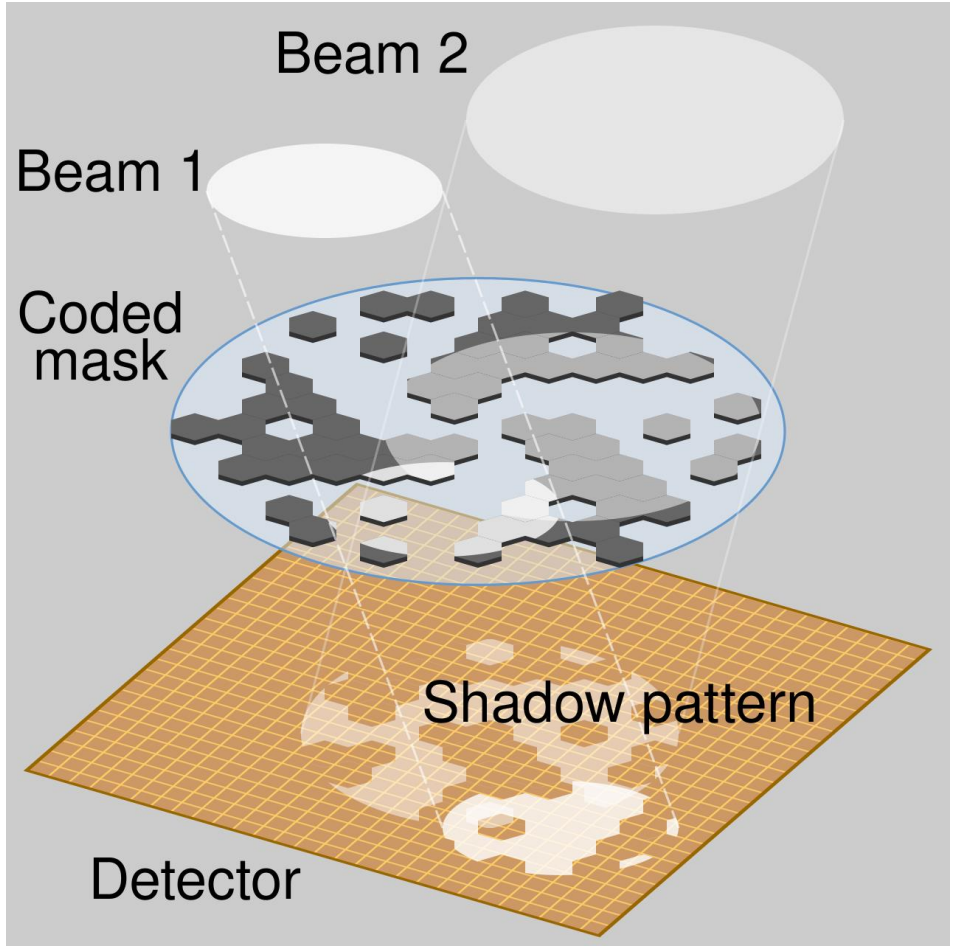
Vitaly Neustroev

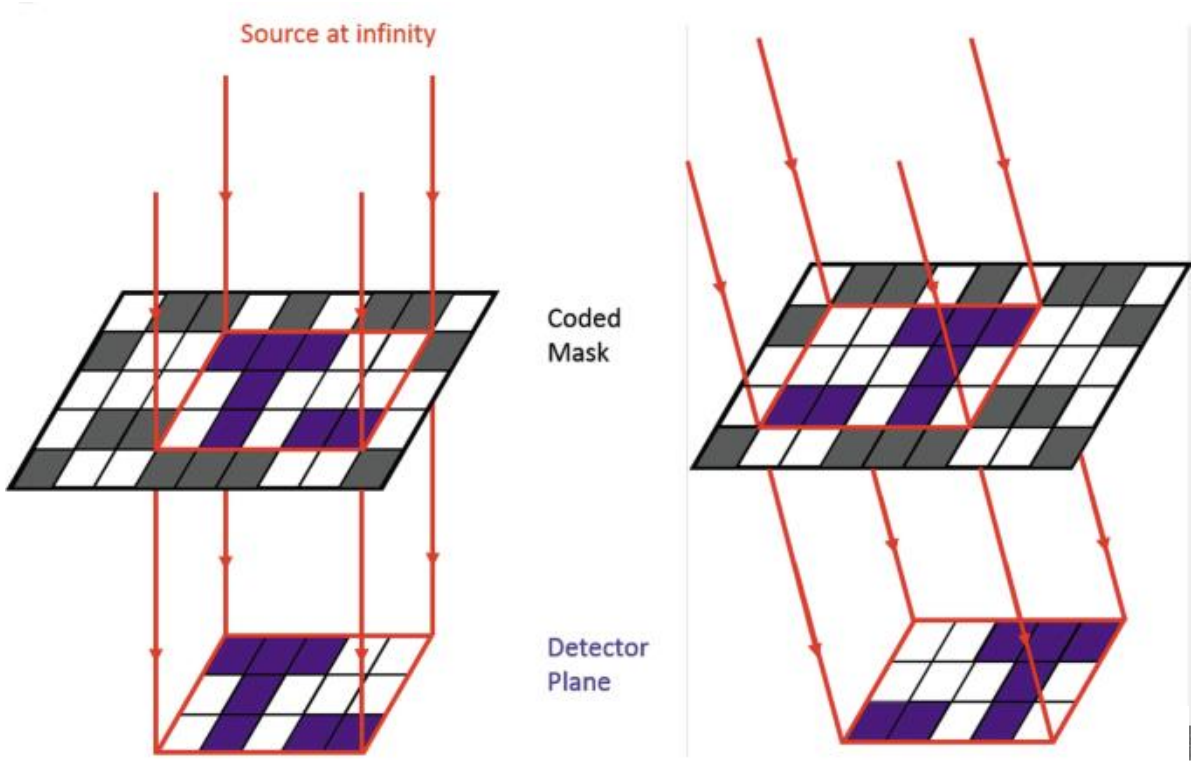
γ-ray and VHE telescopes

The line between a telescope and a detector is blurred.

At these high energies they often work together.

Coded mask imaging

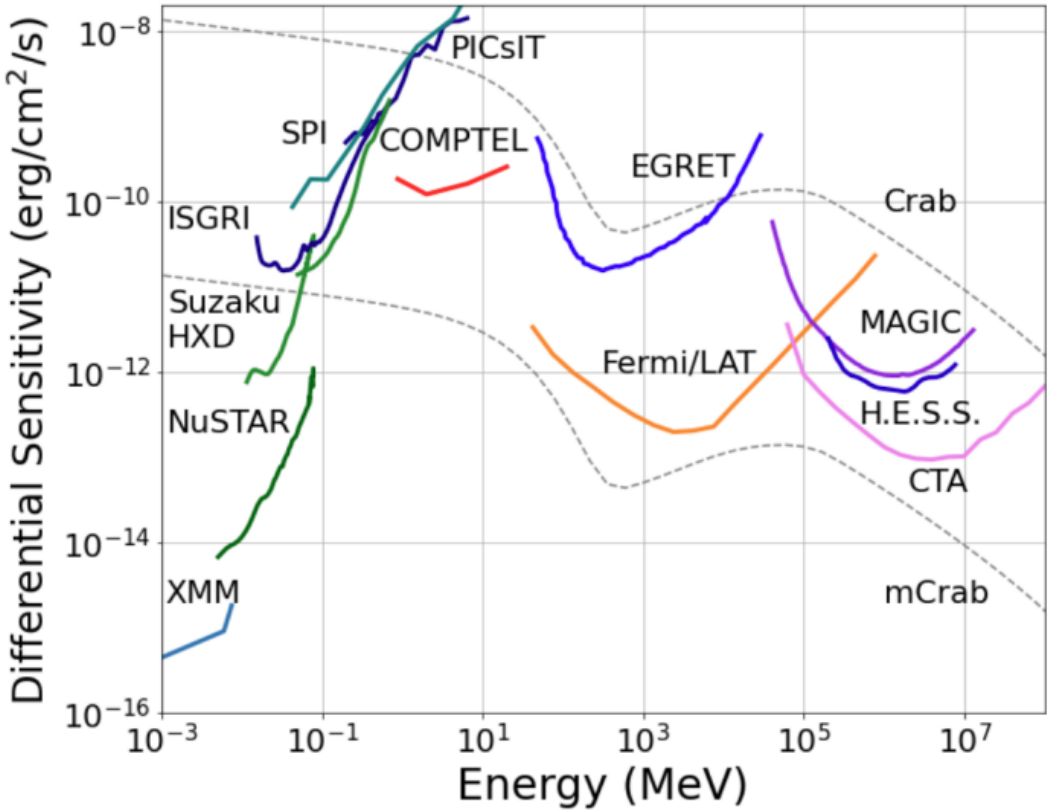




Coded mask design for the Integral satellite

MeV gap

- The soft to medium energy γ-ray regime remains one of the least explored energy ranges in multi-wavelength astrophysics.
- This gap, often referred to as the MeV Gap, is due to sensitivity of detectors.
- The interaction cross-section between photons and matter is lowest in this range, and the dominant mechanism is Compton scattering, which produces long-range secondaries.
- Large detector volumes are needed to stop and contain such photons.



The differential sensitivity for current and past X-ray and γ-ray missions shows the limited performance achieved in the MeV regime. The reduced sensitivity in the range from 100 keV to 100 MeV is referred to as the MeV Gap.

Compton scattering

- In Compton scattering, γ-rays partially transfer their energy to electrons, and by doing so, the electron recoils and the gamma ray is scattered at an angle relative to its initial direction with lower energy. The Compton-scattered photon will then interact a second time, and potentially a third and fourth, before finally absorbing all its energy in the detector.
- Only the complete measurement of the secondary products, i.e. the energized target electron and the de-energized scattered photon, allow for the determination of the initial energy and direction of the incident photon.

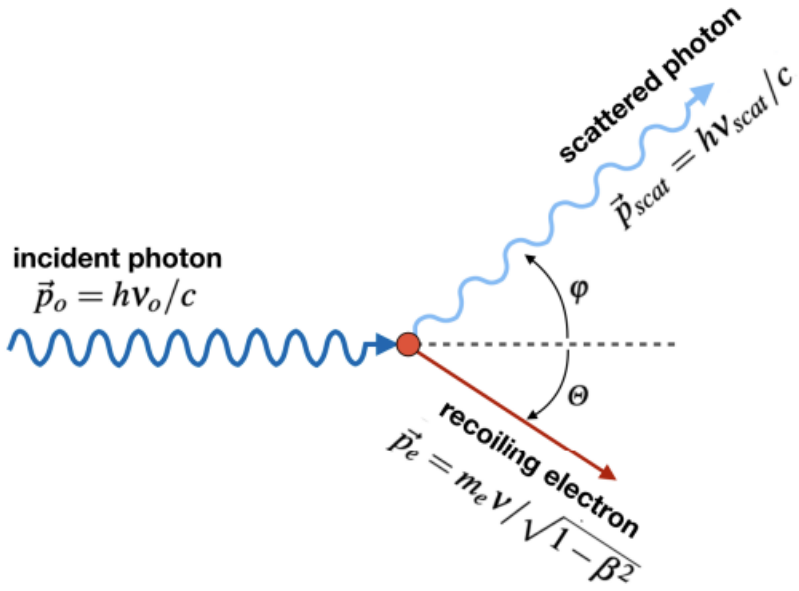
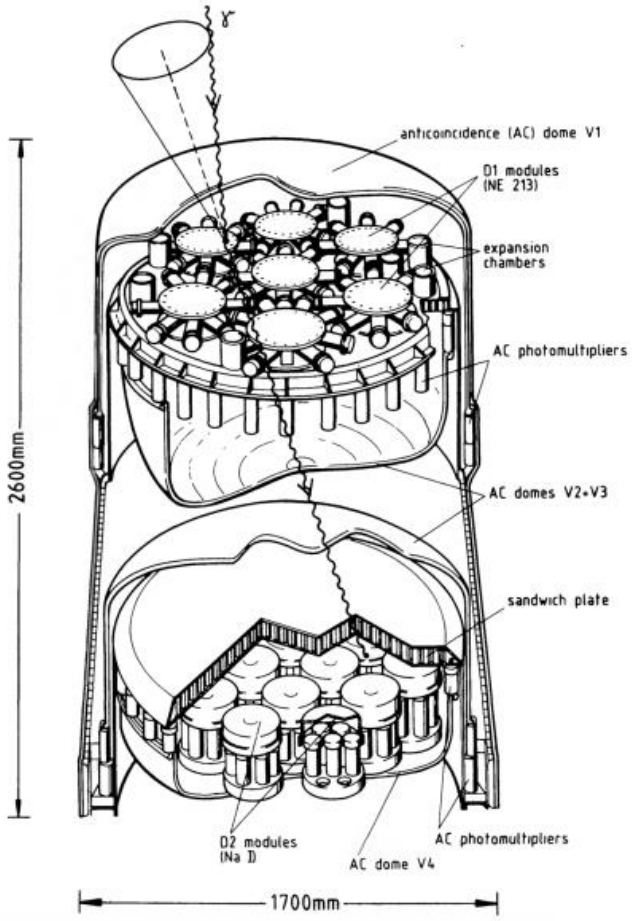
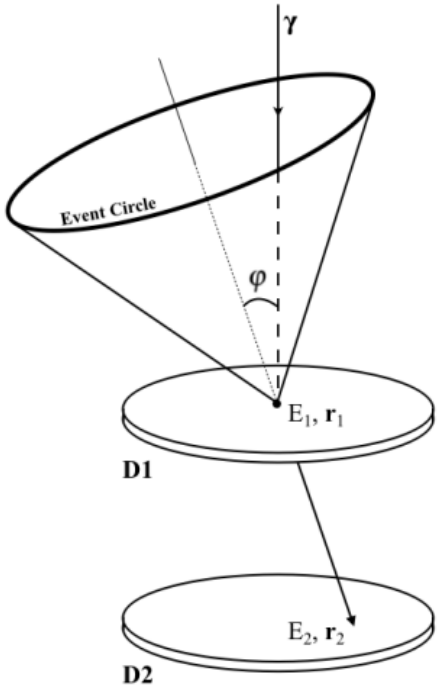


Fig. 2 – A photon with initial momentum $\vec{p}_0 = h\nu_0/c$ scatters off an electron assumed to be unbound and at rest. The photon scatters at an angle φ relative to the initial direction, and has a momentum $\vec{p}_{scat} = h\nu_{scat}/c$. The electron recoils from the scatter at an angle Θ and with momentum $\vec{p}_e = m_e v/\sqrt{1-\beta^2}$. Through the conservation of energy and momentum, the relation between the scattering angle and the final photon energy can be determined.

(from Kierans, Takahashi, and Kanbach, 2022)

Compton Scattering γ-ray telescope



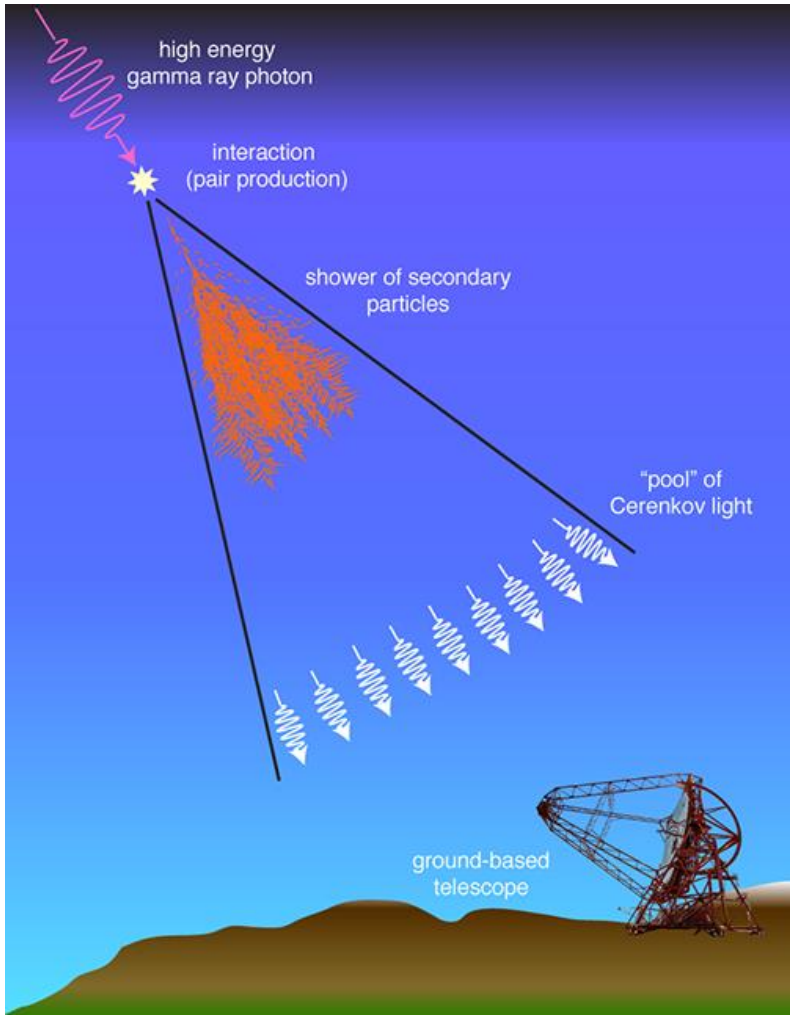


- A classic Compton telescope uses a scattering plane (D1) and an absorption plane (D2) to measure the position and energy of two interactions in a Compton event.
- From these measurements, the original photon direction can be constrained to a circle on the sky called the "event circle."

Čerenkov air shower telescope

- Here the atmosphere acts as a detector for very high energy gamma rays.
- Gamma rays interact in the upper atmosphere to produce secondary particles, these particles are moving at speeds greater than the speed of light in the local medium.
- Čerenkov radiation results when charged particles move at greater than the local speed of light, and because of the rapidly changing electric field cause the local medium to radiate.
- Čerenkov radiation is concentrated in a cone.

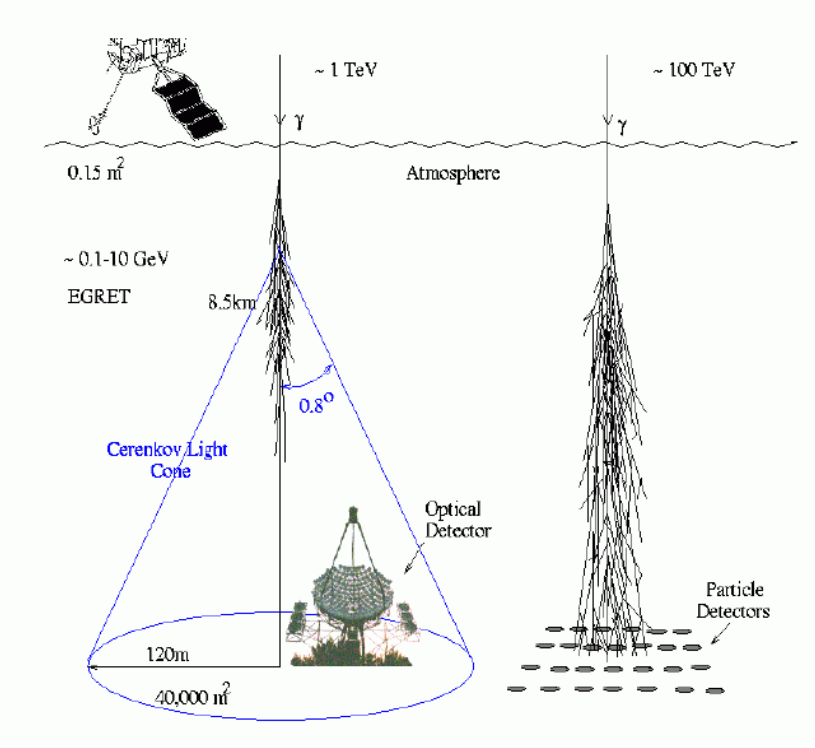
Čerenkov air shower telescopes



- Illustration of the process of detecting a gamma ray using Earth's atmosphere.

(Credit: Diagram by NASA's Imagine the Universe; telescope image from the HESS Collaboration)

Čerenkov air shower telescopes



- Čerenkov telescopes are large collection area telescopes with fairly crude optics, often segmented mirror, usually in the open air without an enclosure.

Čerenkov air shower telescopes



The H.E.S.S. Telescopes

Because the shower and the consequent Čerenkov flash is spread over a wide spatial area, arrays of such telescopes are built to improve the directionality.

The MAGIC telescopes

MAGIC is a system of two 17m-diameter Cherenkov Telescopes located at the Canary island of La Palma Spain).

University of Oulu is a member of the collaboration



The MAGIC telescopes

If you detect a flash in more than one telescope you can determine directionality by accurate timing.



The MAGIC telescopes

The MAGIC telescopes designed to perform gamma-ray astronomy in the energy range from 50 GeV to greater than 50 TeV.

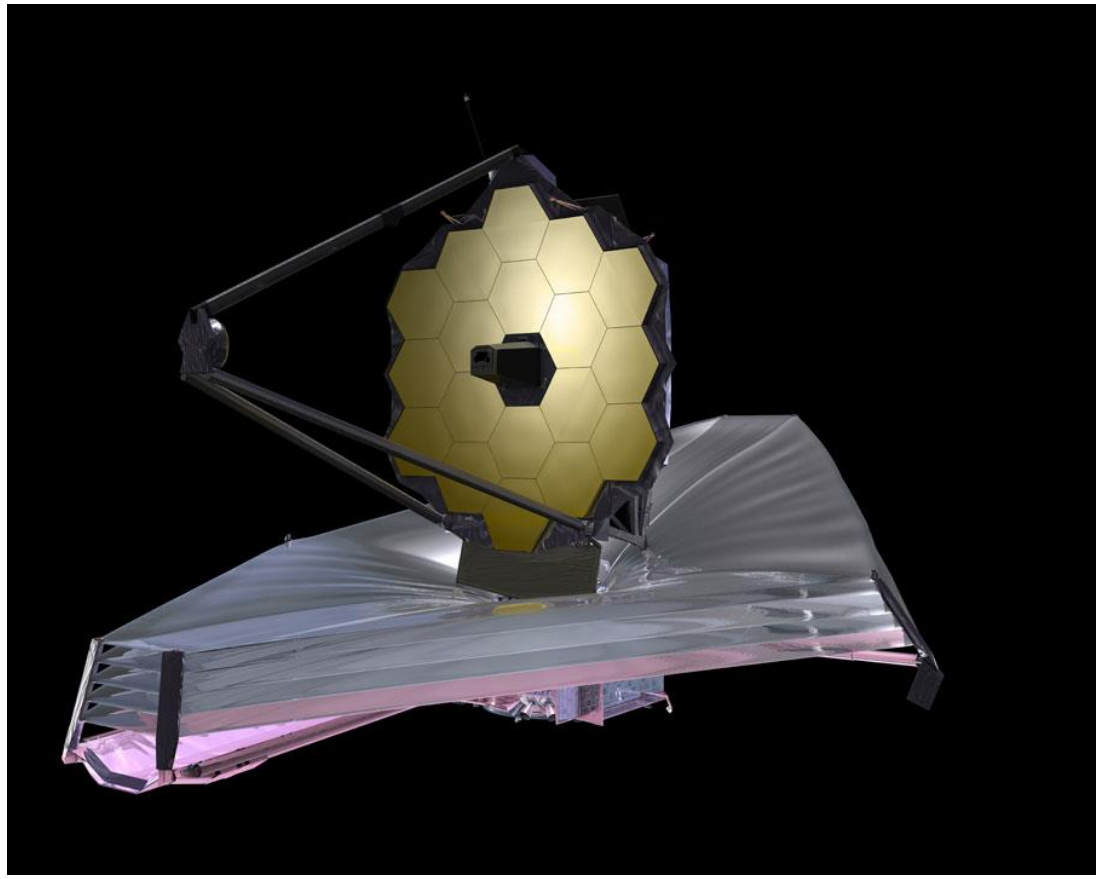


The Cherenkov Telescope Array (CTA)

- CTA is the next generation ground-based observatory for gamma-ray astronomy at VHEs.
- LST-1 (the prototype Large-Sized Telescope) is intended to become the first of four LSTs on the CTA-North site on the island of La Palma.



Homework: James Webb Space Telescope

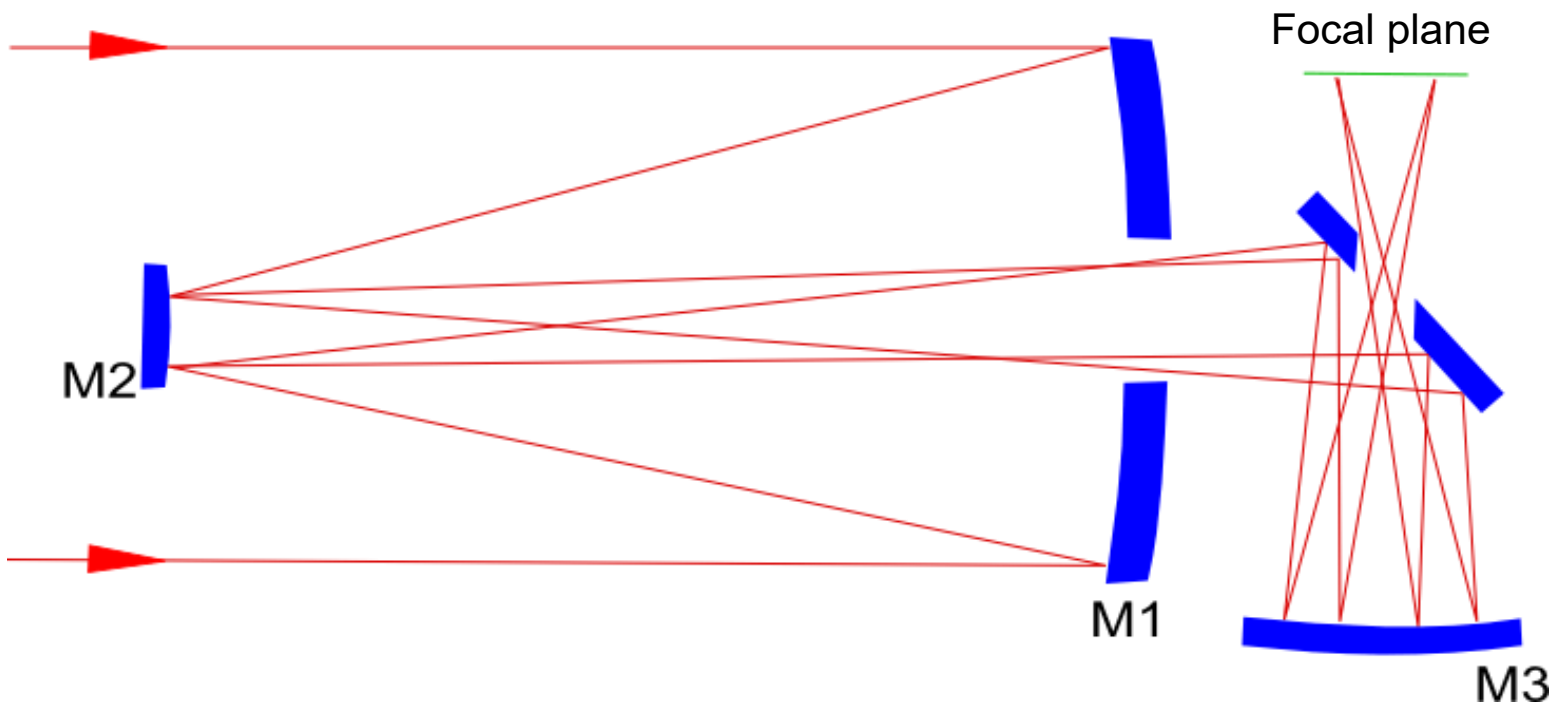


- JWST's optical design is a three-mirror anastigmat (Korsch telescope).

- Homework: Plot a scheme and explain differences with the classical Cassegrain telescope. What are the advantages?

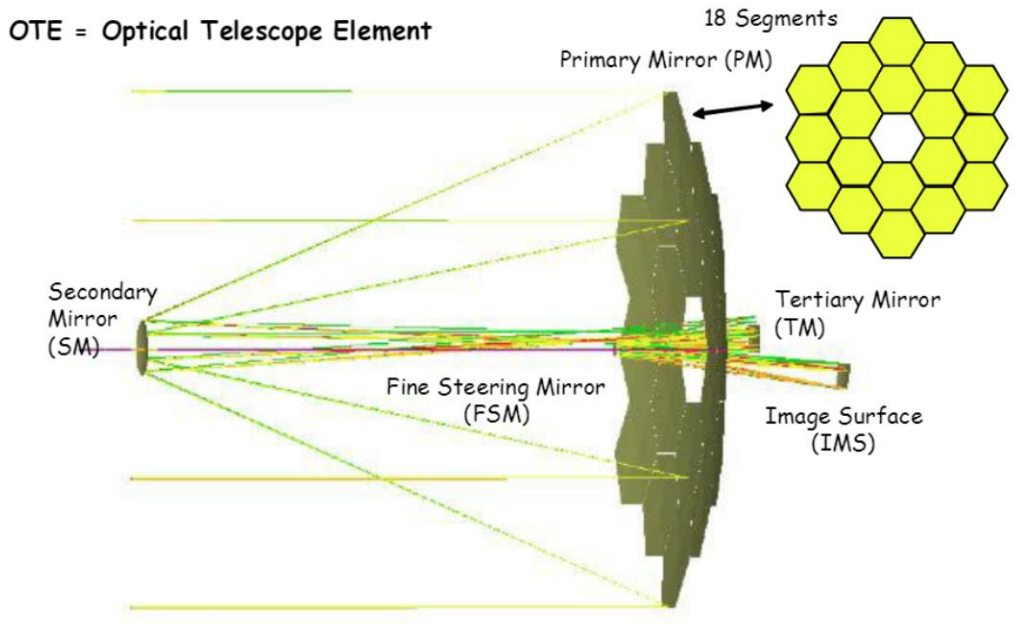
Korsch telescope

Optical design of a Korsch telescope is a three-mirror anastigmat.



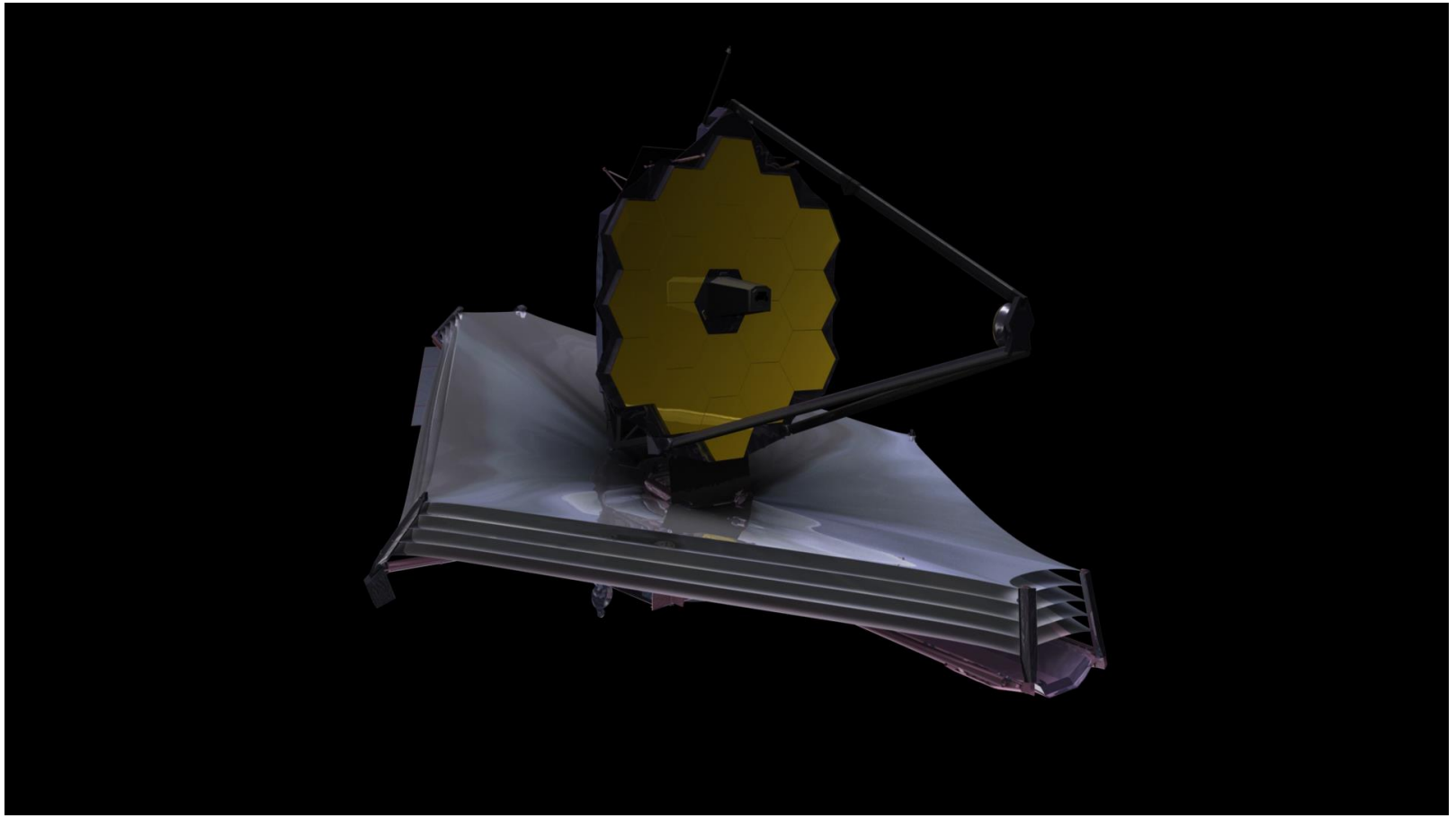
- The primary mirror is concave.
- The secondary is convex, and it works slightly off-axis.
- The tertiary **removes the resulting astigmatism** and it also **flattens** the focal plane.
- This also allows for a **wider** field of view.

JWST's optical design



- JWST's optical design is a three-mirror anastigmat (Korsch telescope).

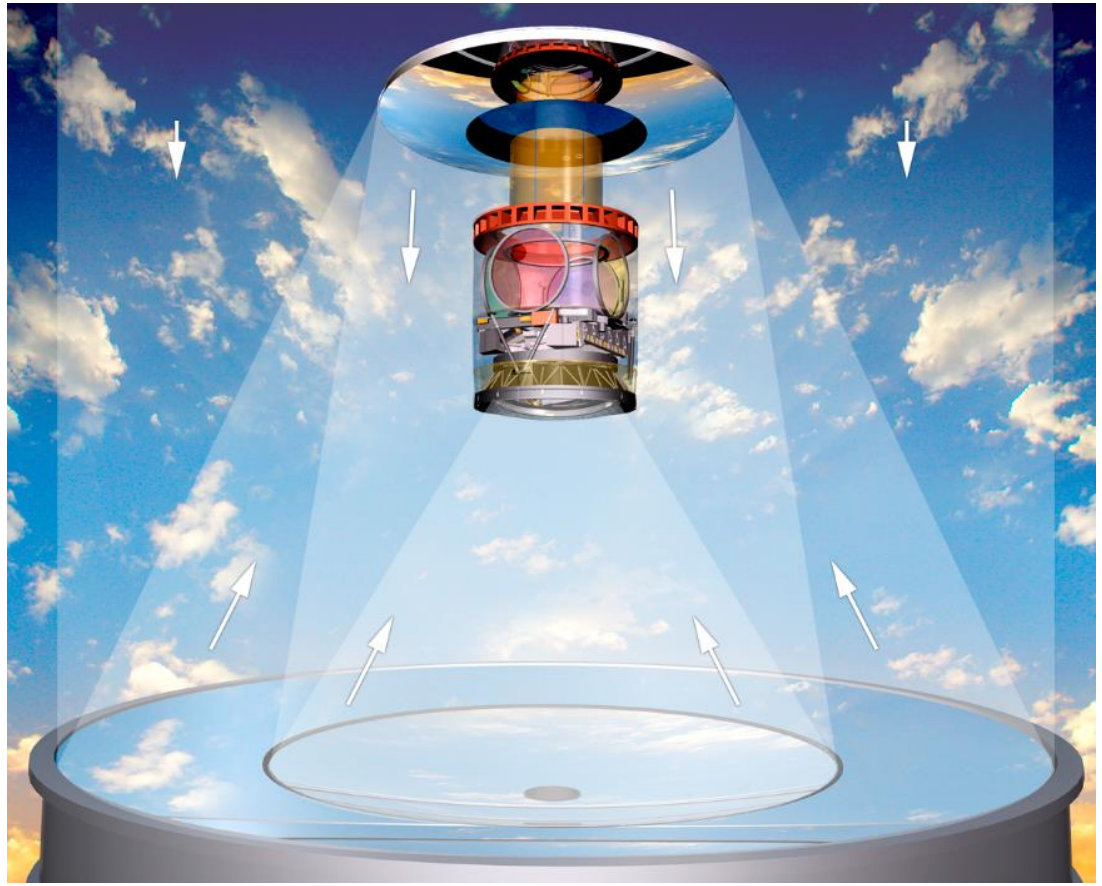
Homework: James Webb Space Telescope



The path which light follow as it hits the primary mirror, and then reflected to the secondary, and then in through the aft optics assembly where the tertiary and fine steering mirrors are.

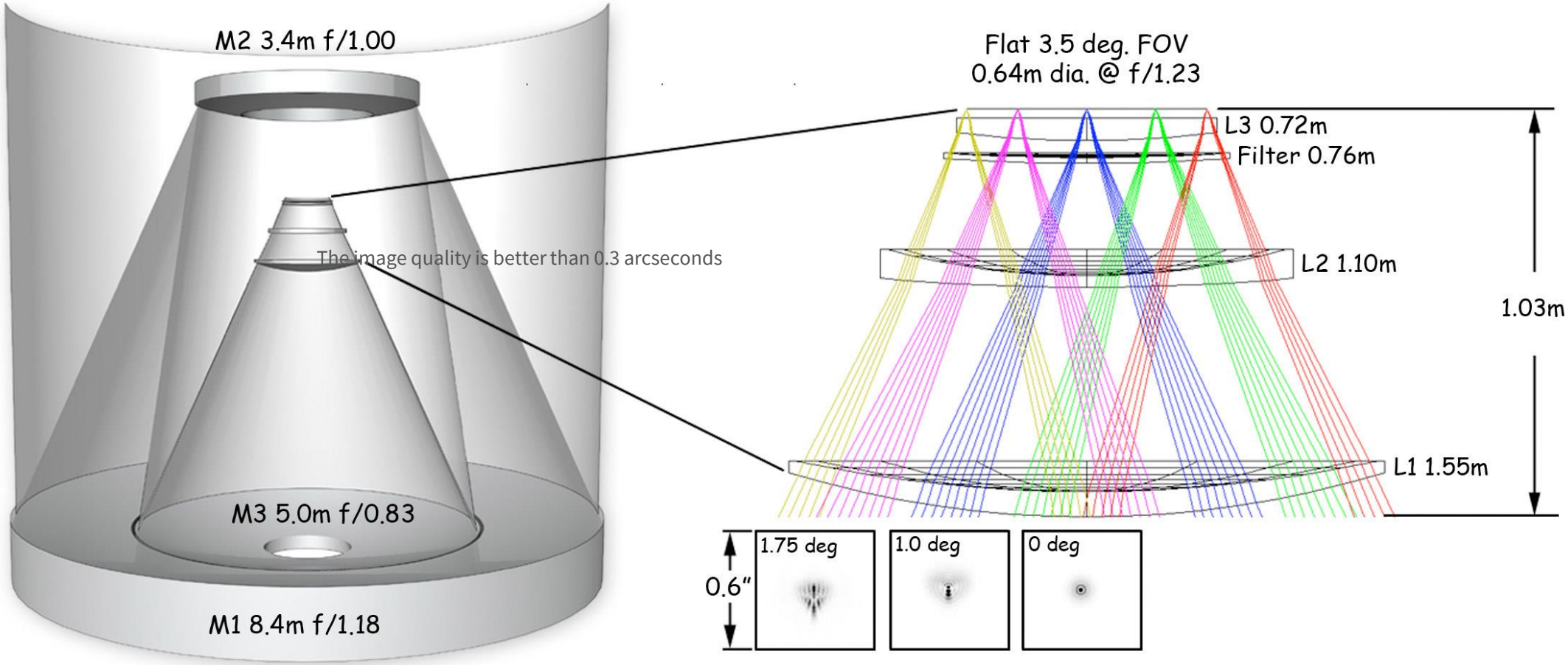
https://youtu.be/y9Z2GbFJWmo

Vera C. Rubin telescope optical design



- Rubin Observatory views a 9.62 square-degree patch of sky, more than 40 times the area of the full moon.
- Optical design: a compact three-mirror telescope augmented by a large refractive corrector.
- Primary mirror is highly annular (ring-shaped), having an outer clear aperture of 8.36 m and an inner diameter of 5.12 m, giving an effective collecting area of a 6.67 m filled aperture.

Vera C. Rubin telescope optical design



Rubin Observatory Optical Design

The image quality is better than 0.3 arcseconds

Diffraction and Spatial resolution

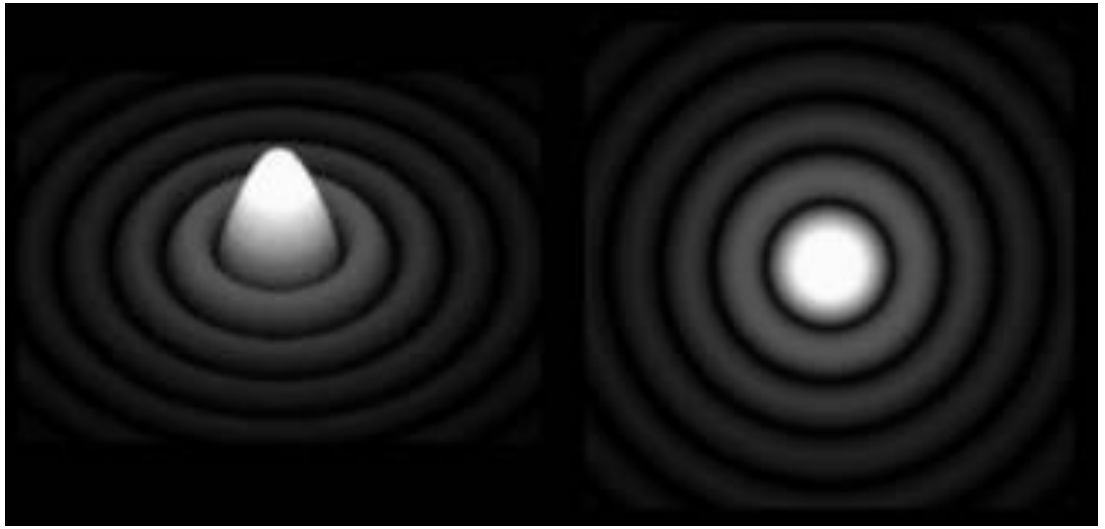
Seeing

Turbulence in atmosphere

Adaptive Optics

Diffraction and Spatial resolution

- Airy disk:



- The first dark ring is at an angular distance of $1.22 \lambda / d$ (radians) from the center. This is often taken as a measure of resolution in an ideal telescope.

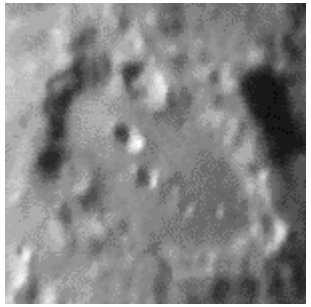
Diffraction and Spatial resolution

- The fundamental limit to the spatial resolution of a circular aperture telescope is set by the size of the Airy diffraction pattern, but **in practice** the resolution is often worse than that, particularly at wavelengths shorter than 10 μm (100000 Å). Causes are:
  - Imperfections in manufacture or support of the reflecting surface;
  - Aberrations;
  - **Distortion of the wavefront by the atmosphere (seeing)!**

Seeing and Scintillation

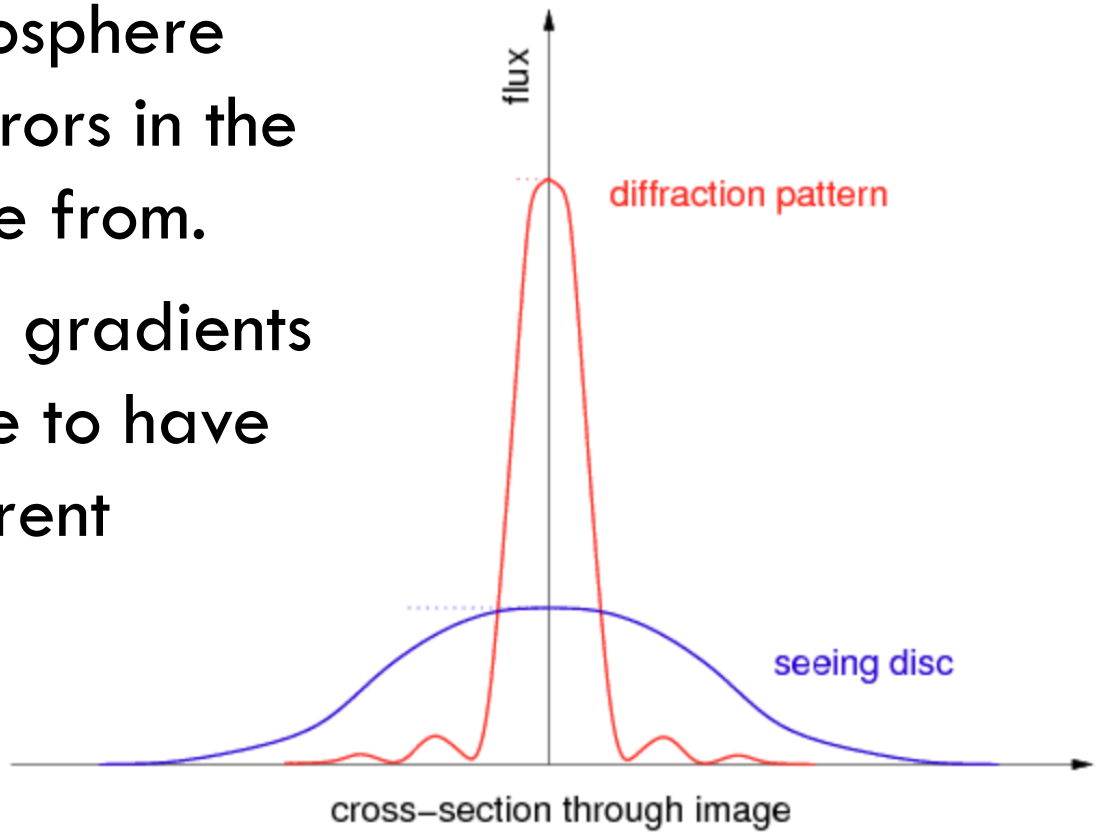
- The Earth's atmosphere is not a quiescent plane-parallel ideal region. It is, in fact, a turbulent mix of gasses, always in motion. Because the index of refraction varies with temperature and pressure motions in the atmosphere result in variations in refraction of the light passing through.
- This results in rapid fluctuations in the apparent brightnesses of stars (scintillation), and motions in the apparent positions or variations in the apparent sizes of stars (seeing).





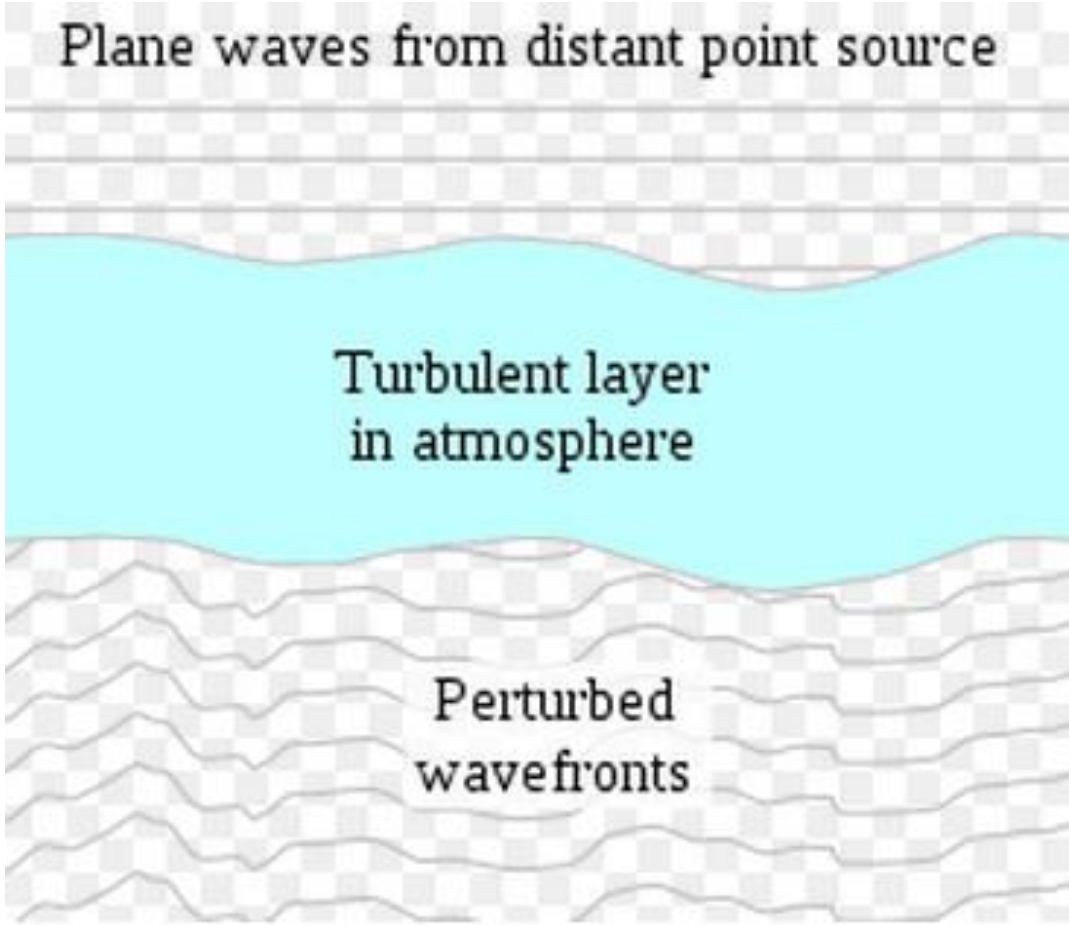
Distortion of the wavefront

- At optical and near infra-red wavelengths the spatial resolution is not set by the diffraction limit.
- Distortion of the wavefront by the atmosphere causes phase errors, therefore small errors in the direction that the light appears to come from.
- Turbulence combined with temperature gradients causes some pockets of the atmosphere to have different temperature and hence different refractive index.



Turbulence in atmosphere

How optical wavefronts from a distant star may be perturbed by a layer of turbulent mixing in the atmosphere.



Adaptive Optics

- In **Adaptive** Optics, mirrors are deformed to correct for distortions of the wavefront by the atmosphere.
- Not to be confused with Active Optics, in which the distortions corrected are those caused by mechanical or thermal deformation of the telescope.
- Adaptive optics corrections need to be calculated and applied at a frequency up to hundreds of Hz (c.f. ~ 1 Hz for active optics).

Kolmogorov theory

- The theory of atmospheric turbulence is due to the Russian physicist Kolmogorov.
- The turbulence above a telescope occurs at different scales and heights.
- It can be characterised with a single scale length parameter, usually called the Fried parameter, denoted by r_0.

Kolmogorov theory

- Fried's coherence length r_0 is the distance over which the phase difference is one radian:

$$r_0 \approx 0.114\,\lambda^{6/5}\left(\frac{\cos z}{550^2}\right)^{0.6}\ \text{m}$$

λ is the operating wavelength in nm
z is the zenith angle

- Fried's parameter defines the maximum diameter of a telescope before it becomes seriously affected by atmospheric turbulence:
- In visual wavelengths, telescopes with a diameter >11.5 cm is always have their images degraded by turbulence.

The seeing disc

- Full Width Half Maximum (FWHM) of the point spread function due to atmospheric turbulence (the seeing) is given by:

$$\beta = 0.98\ \lambda\ /\ r_0 \quad \text{(radians)}$$

- This is quite close to the diffraction radius of a telescope of diameter $D=r_0$:

$$\vartheta \cong 1.22\ \lambda\ /\ D \ \text{(radians)}$$

- For larger telescopes we need to use **adaptive optics**.

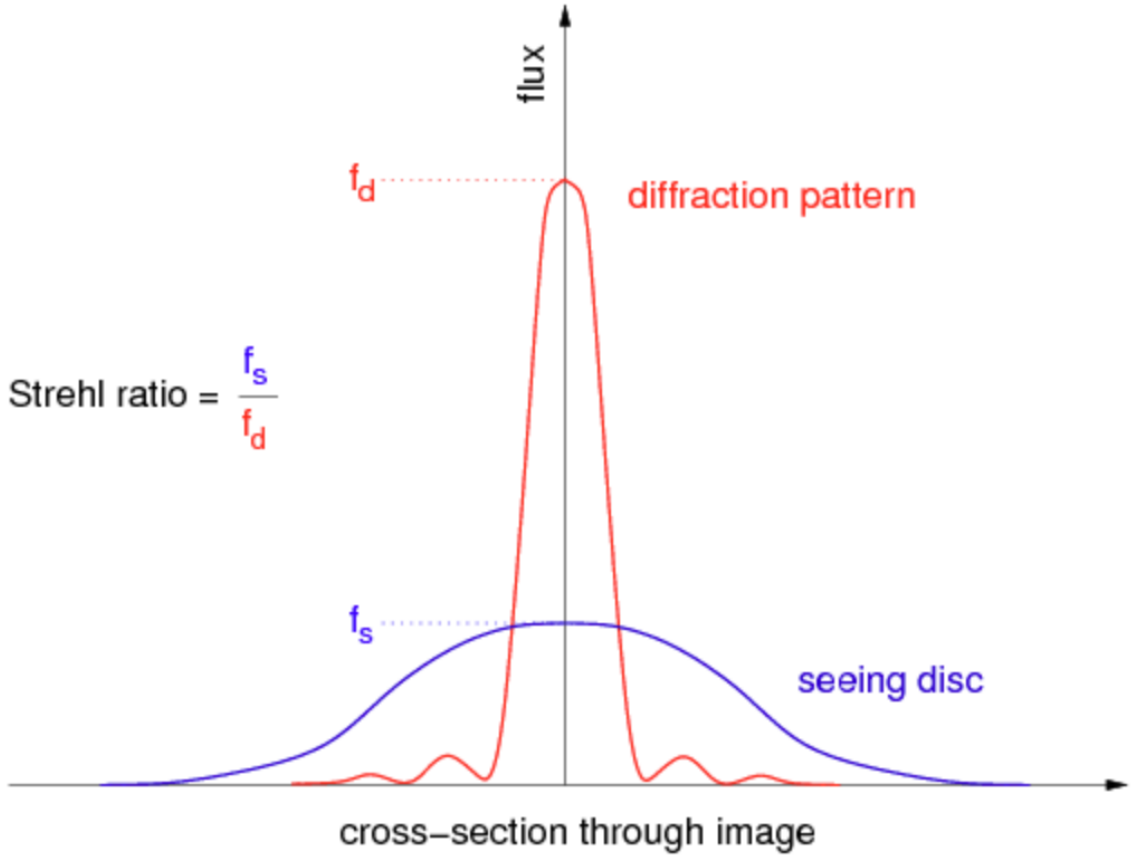
The Strehl Ratio

- Performance of an adaptive optics system is characterized by the **Strehl Ratio**, which is the ratio of the central intensity in the real point spread function, to that in the ideal point spread function for that telescope
  (i.e. the Airy or diffraction function).

- For a seeing limited telescope the Strehl ratio is given by:

$$S = (r_0/D)^2$$

D is the telescope diameter.



The seeing disc of a star superposed on the theoretical diffraction pattern (figure from Vik Dhillon's lectures).

Adaptive Optics

- The aim of an adaptive optics system is to produce a Strehl ratio as close as possible to 1.
- The higher the Strehl ratio, the more the image is concentrated and hence the higher the spatial resolution.
- The Strehl ratio recorded by a telescope without adaptive optics is typically only a few per cent.
- In general, the higher the ratio of the telescope diameter D to r_0 the more difficult it is to correct the wavefront.
- If D/r_0 is not too high (2-5) then considerable improvement in the image quality can be obtained using a **tip-tilt** corrector.
  - This is the lowest order adaptive optics corrector, which is a flat mirror which can be tilted in two orthogonal planes, to keep the image centred.

Adaptive Optics

- Adaptive optics is more difficult at shorter wavelengths.
  - r_0 is dependent upon wavelength, in the sense that it is larger for longer wavelengths (slide 189):

$$r_0 \propto \lambda^{6/5}$$

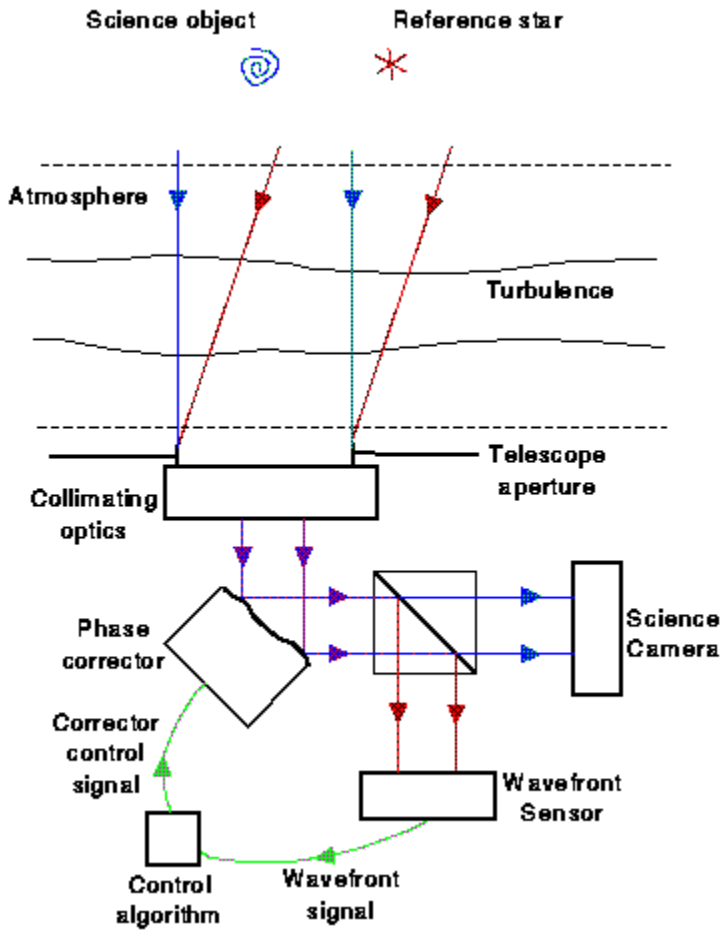
  - The timescale on which the properties change is faster at shorter wavelengths.

$$\tau_0 = 0.31\ r_0 / \langle V \rangle$$

  $\langle V \rangle$ is the altitude averaged wind velocity, typically 10 m/s.
- For this reason, most practical adaptive optics systems work in the near infra-red.

r_0 at 20000 Å is 5.9 times what it is at 5000 Å.

Practical Adaptive Optics system



- **Collimating optics** – to ensure phase correction is correct for all parts of the wavefront.
- **Phase corrector** – deformable mirror with a number of piezoelectric actuators.
- **Beam Splitter** – either dichroic filter (to reflect blue and transmit red) or a pickoff to send only the reference star to the wavefront sensor.
- **Wavefront sensor** – usually a fast frame transfer CCD to detect the wavefront from the reference star.
- **Control loop** – to calculate the phase corrections from the output of the wavefront sensor.
- **Science camera** – CCD or IR hybrid array.

The Isoplanatic Angle

- You can see from the diagram that the light from the reference star travels through the atmosphere at a slightly different angle to that from the science object.
- How far off does this have to be before the reference star does not any more sample the relevant path through the atmosphere?
- This depends upon a number of things including the height of the turbulent layer in the atmosphere.

$$\vartheta_0 \approx 0.31\ r_0 / \langle h \rangle$$

<h> is some average turbulence altitude. Typically <h> ≈ 5 km.
So, if r_0 = 20 cm, then ϑ_0 = 8.25 arcseconds.

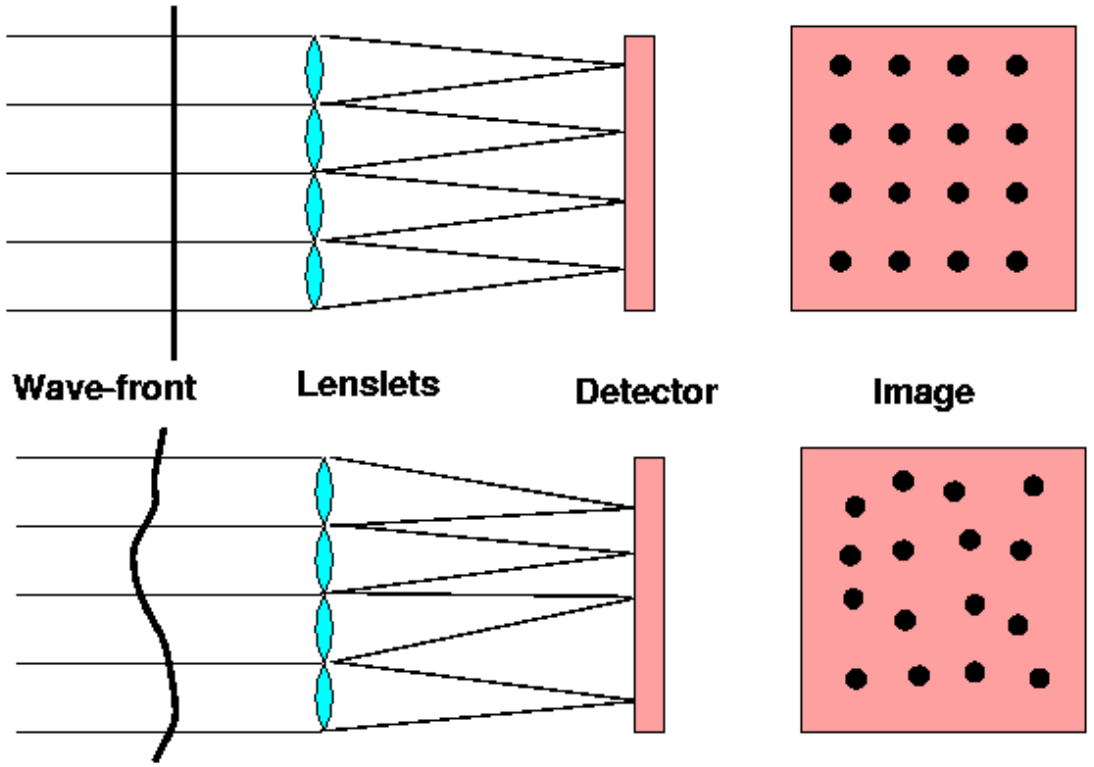
The field of view is very small! Again, much larger in the infrared.

Wavefront sensors

Simplest type is

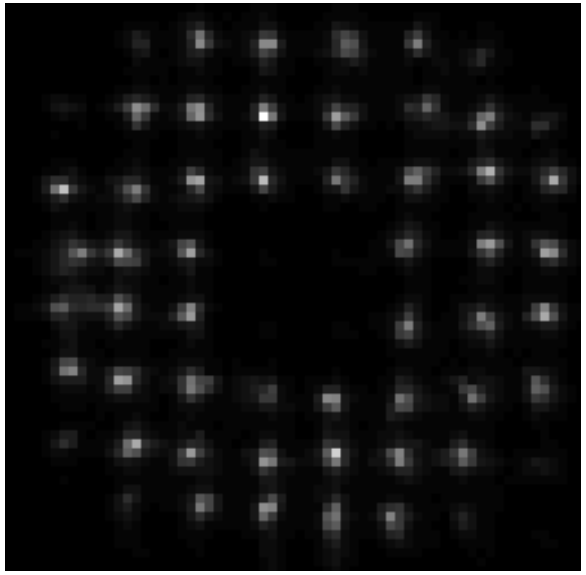
a Shack-Hartmann wavefront sensor.

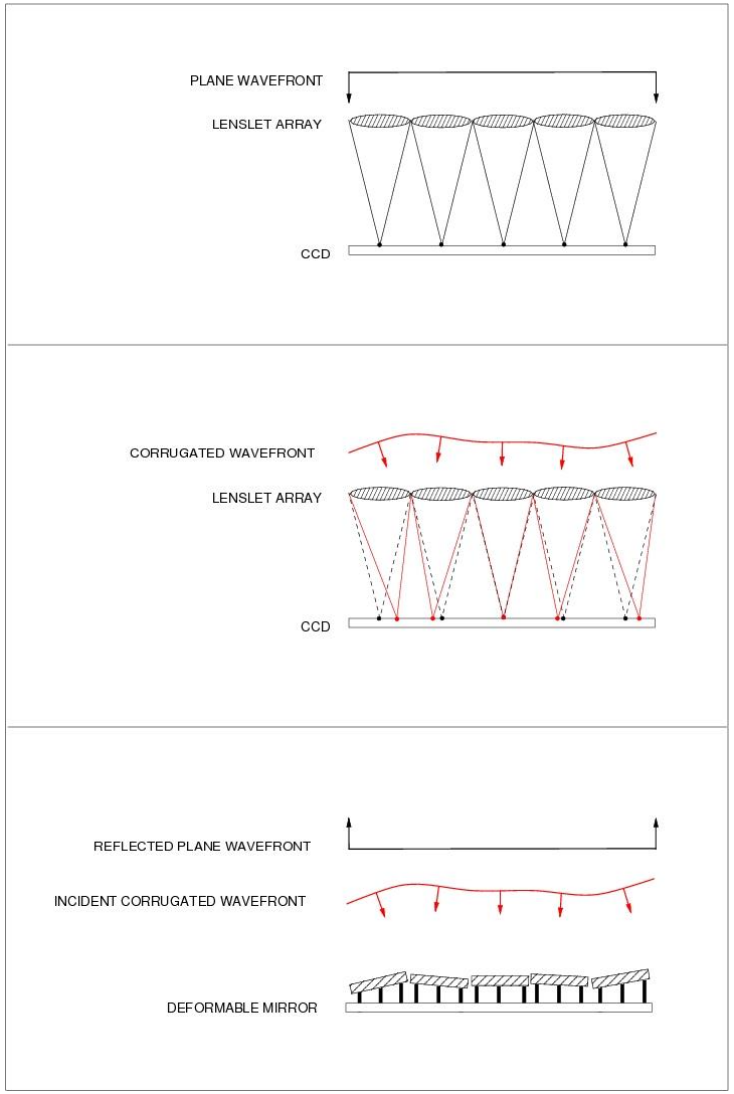
- Series of lenslets which each image part of the pupil plane onto a detector.
- Spots from the lenslets move around as the wavefront changes.



Shack-Hartmann Wavefront Sensor

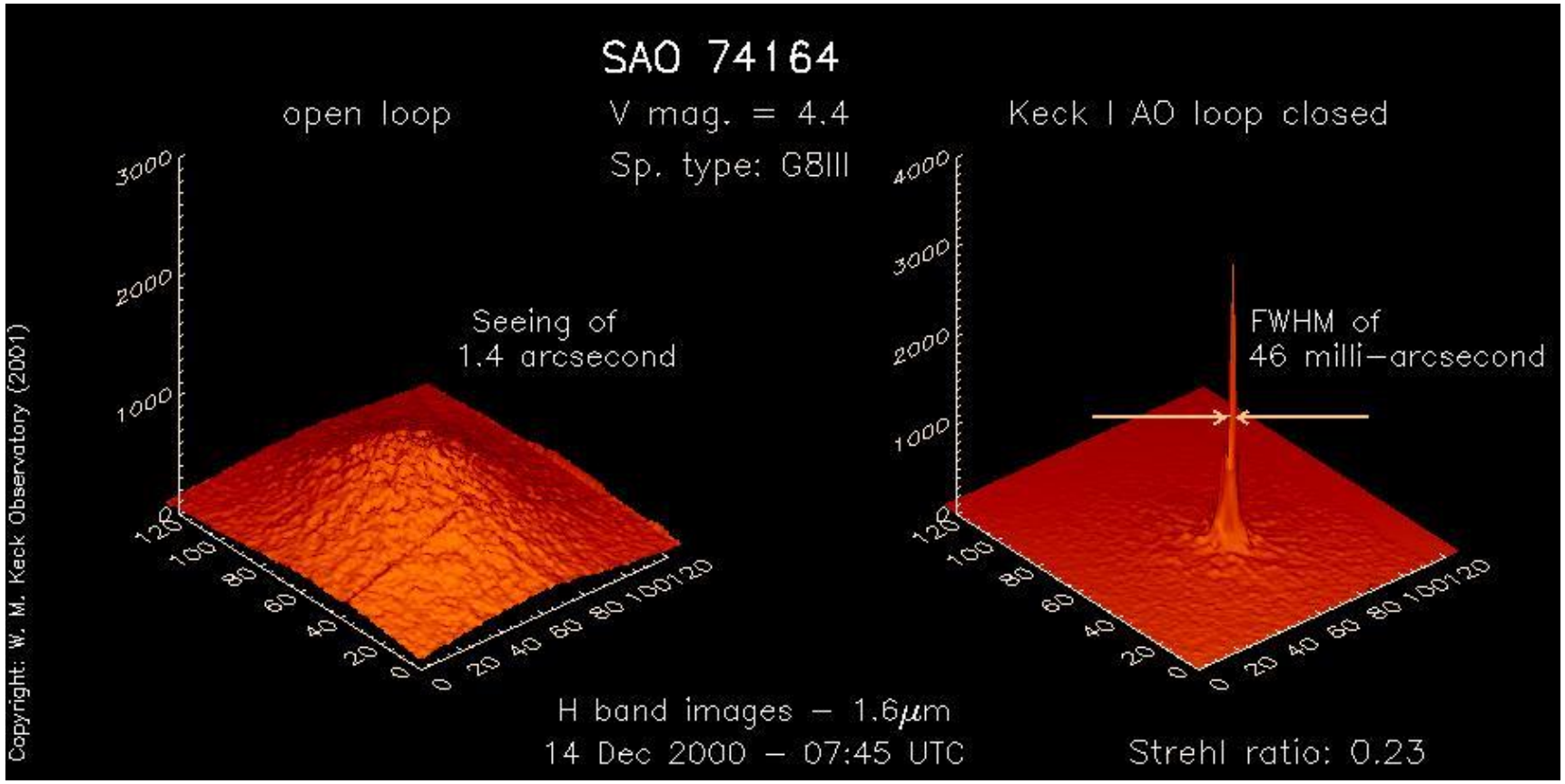
- Software detects the centroid of each spot.
- Then it models the wavefront distortions using a set of functions called Zernicke polynomials.
- Feedback to a membrane mirror to alter shape to correct the wavefront distortion.





Figures from Vik Dhillon's lectures.

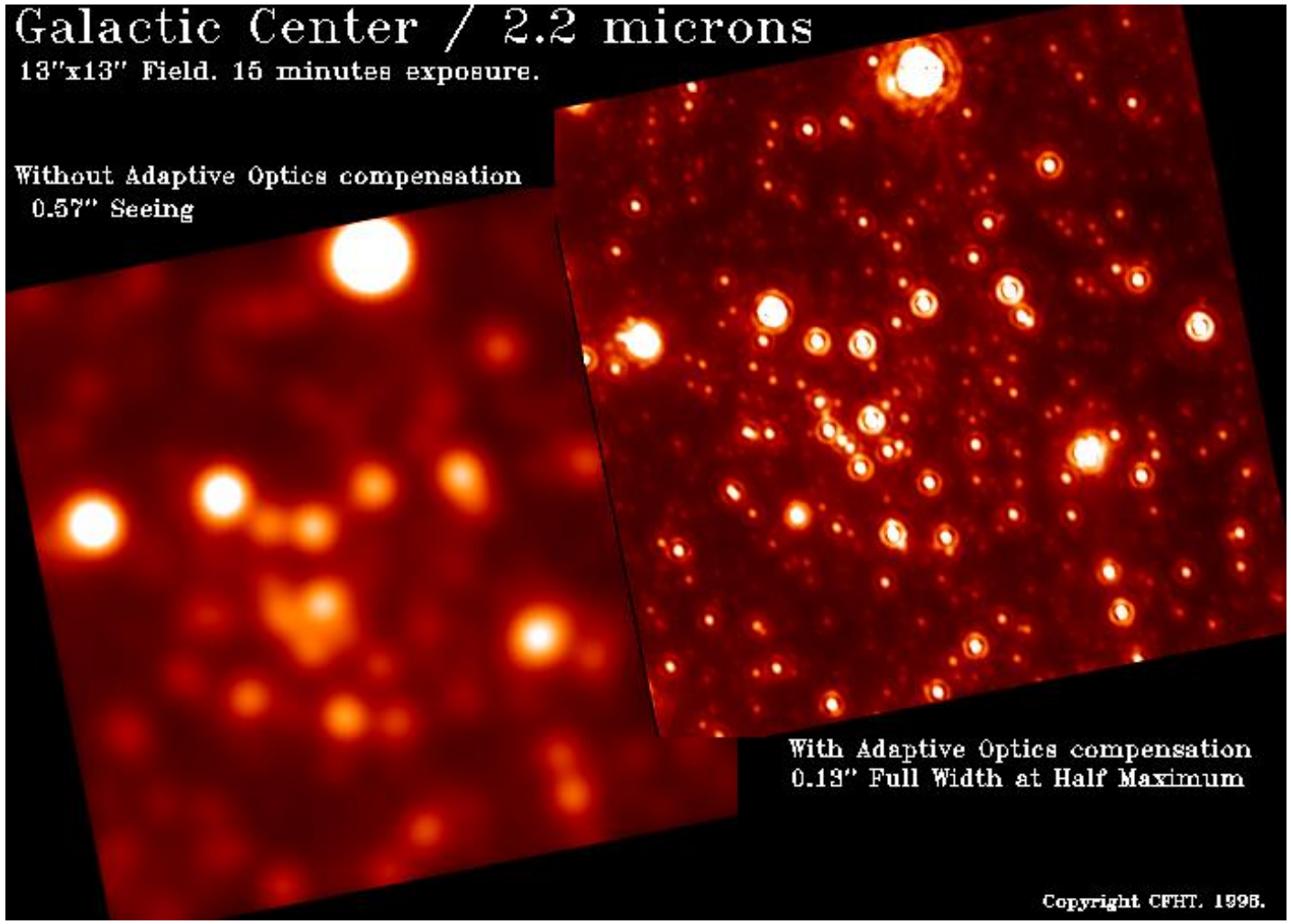
Effect of adaptive optics on a star image at the Keck telescope



This is cheating a bit because the star is its own guide star!

Effect of adaptive optics (real, not cheating)

Infrared image of the centre of our galaxy, without (left) and with (right) adaptive optics



Problems of adaptive optics

- You need a bright star as a guide star
- The wavefront is only corrected for objects near that star, within the isoplanatic angle.
- So, if there isn't a star near the objects you are interested in, why not put one there?

Laser guide stars

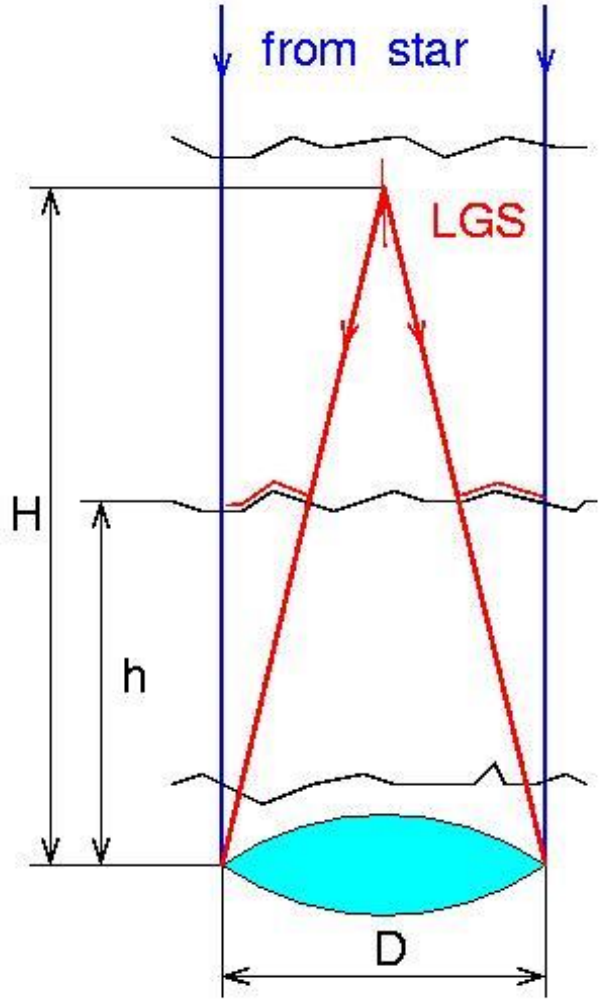
- Two types of laser used to create artificial guide stars:
  - Sodium lasers – Sodium layer at about 90 km altitude is produced from micrometeorites. This can be excited by a laser beam tuned to one of the D lines (usually D2 at 5890Å).
  - Rayleigh Scattering Lasers – light from the laser is back scattered by air density fluctuations. The laser beam is focused at an altitude of ~20 km, in addition the beam is pulsed and the return is gated with a fast shutter, so that the altitude is defined by the light travel time to and from the layer.

Rayleigh Scattering lasers

- Rayleigh scattering is more efficient in blue and ultraviolet wavelengths (hence the sky is blue). The return beam is therefore bluer than the outgoing beam.
- There is a tradeoff between the rate of the pulses and the altitude that the laser is gated to. If the altitude is 20 km, then the light travel time there and back is $(4 \times 10^4) / (3 \times 10^8)$ seconds, and the laser guide star cannot be used for atmospheric fluctuations faster than this.

The cone effect

- The Laser Guide Stars are not at infinite altitude, and therefore the light from the guide star does not pass through the same atmosphere as the beam from the astronomical target.
- The beam from the star suffers different wavefront distortions to the beam from the laser guide star.
- This problem is worse for lower altitudes and for larger telescopes.



The cone effect

- The maximum size of telescope that can be used with a laser guide star at altitude H is given by:

$$d_0 \approx 2.9\ \theta_0\ H$$

where ϑ_0 is the isoplanatic angle.

- For realistic values of the isoplanatic angle at optical wavelengths, d_0 is about 1 metre for Rayleigh scattering lasers (gated at 20 km altitude) and 4-8 metres for a sodium laser.

The tilt problem and its solution

- The laser beam passes through the atmosphere twice, once on the way up and once on the way down.
- The lowest order terms in the wavefront correction, called Tip and Tilt, cancel out and measurement of the Laser Guide Star does not sense them.
- Use a natural guide star for tip and tilt only. You can do this in the infrared as the isoplanatic angle is larger.
- Use a polychromatic Laser Guide Star. Sodium can be excited to higher levels to emit other lines (in the ultraviolet and infrared). Using the wavefronts at more than one wavelength the atmospheric tilt can be calculated.

Practical problems of laser guide stars

- Sodium lasers are expensive and unreliable.
- Require high power consumption.
- Can make yourself very unpopular with other telescopes on the mountain looking in the same direction.
- Want to be very sure you have no planes flying overhead as you can blind a pilot.



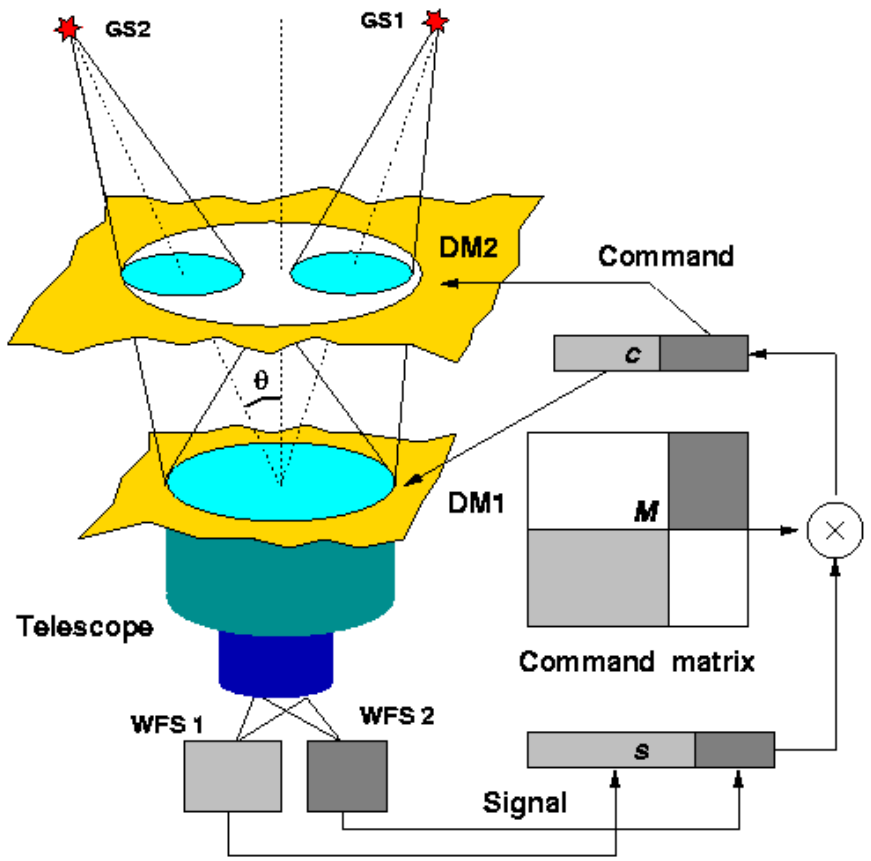


Sodium (left) and Rayleigh (right) laser guide stars in action.

Multi Conjugate Adaptive Optics

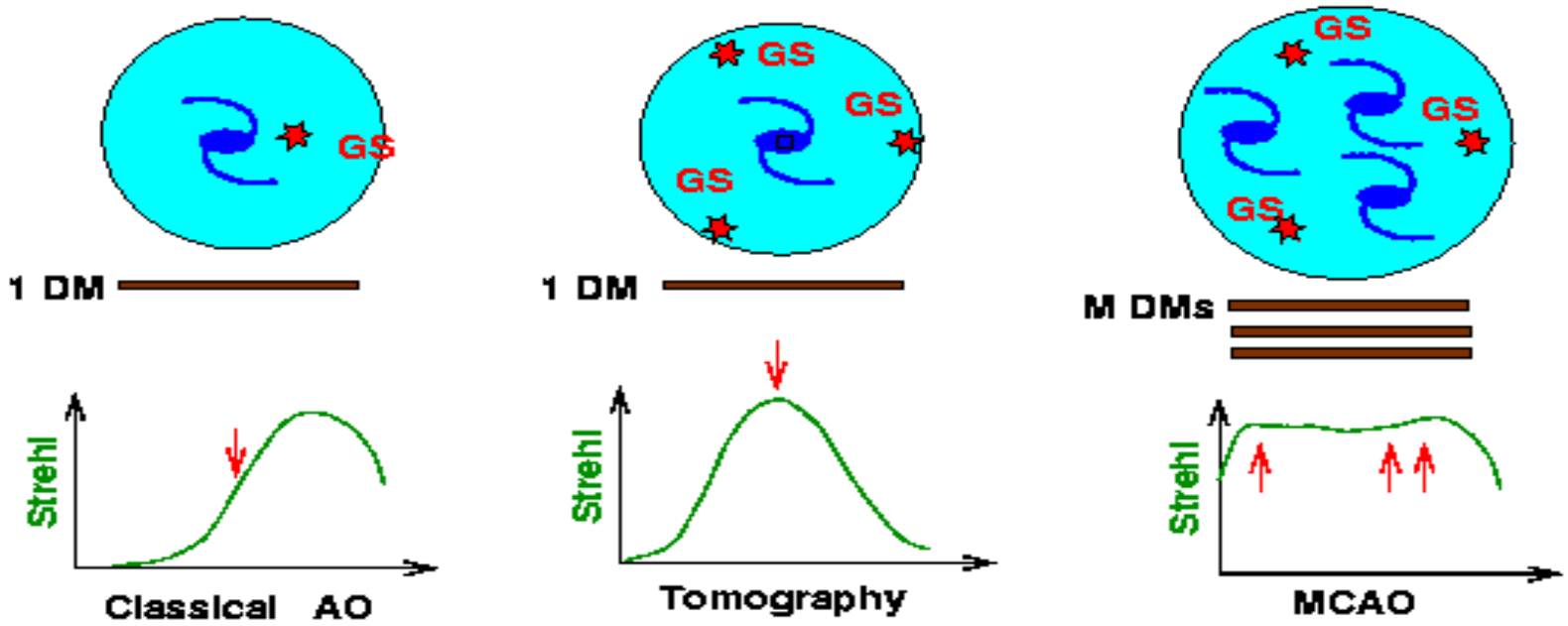
- With only one guide star, we can only model a height averaged wavefront distortions.
- With more guide stars, we can begin to make a three dimensional model of the turbulence, and correct the turbulence in different layers.
- This also helps correct the cone effect, as you can now model each layer and know which part of that layer a particular beam passes through.
- In Turbulence Tomography you use a number of guide stars to model the 3 dimensional turbulence and work out better what to do with your deformable mirror.
- In Multi Conjugate Adaptive Optics you have a number of deformable mirrors which correct the distortions for a number of turbulent layers.

Multi Conjugate Adaptive Optics



- MCAO with 2 laser guide stars and 2 deformable mirrors. In the high layer the light from each LGS passes through a different parft of the atmospheric layer.

Tomography and MCAO



Horizontal axis is across the image in each case, vertical axis is Strehl ratio.

Tomography and MCAO

- In classical AO your correction is strictly valid for the position of your guide star, and is less valid for your objects
- In tomography, you use a three dimensional model to make the right corrections for the object you are interested in.
- In MCAO you have many deformable mirrors, and your corrections are valid over a much wider field.
- MCAO greatly increases the isoplanatic angle.



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 5

Vitaly Neustroev

Detectors

The central theme of this course is the detection and characterization of photons with both ground-based instruments and instruments aboard spacecraft.

Detector parameters (1)

- **Quantum Efficiency (QE):** Ratio of the number of detected photons C and the number of incident photons N, as a function of wavelength, $\eta(\lambda)$:

$$\eta = C/N$$

- **Detective Quantum Efficiency:** Square of the ratio of the output signal-to-noise ratio to the input signal-to-noise ratio.

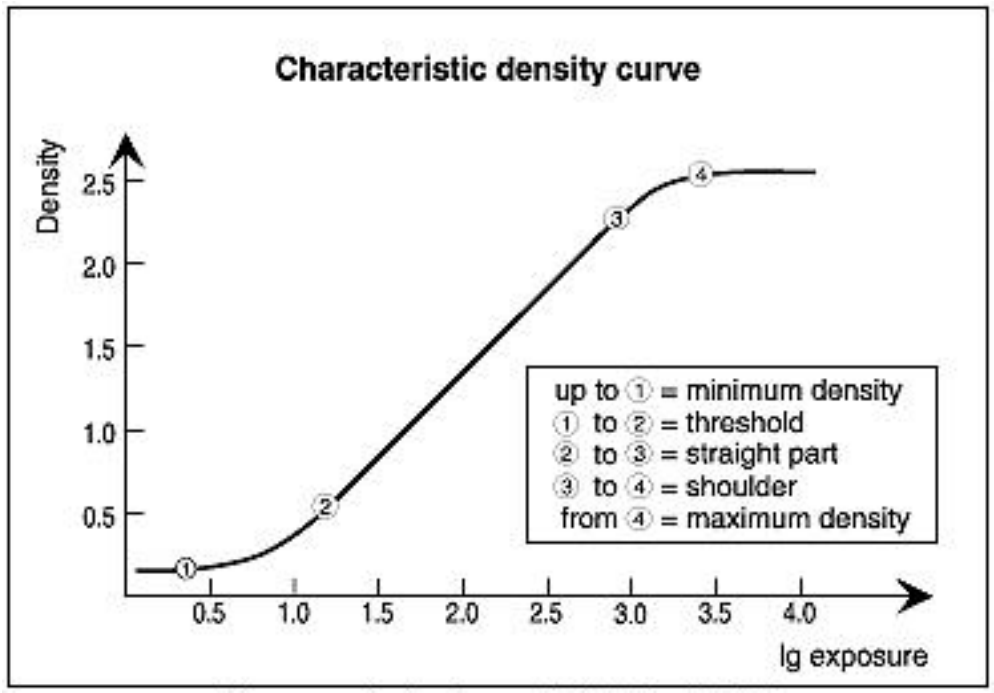
- DQE is always less than or equal to the QE because of the effect of noise.

Detector parameters (2)

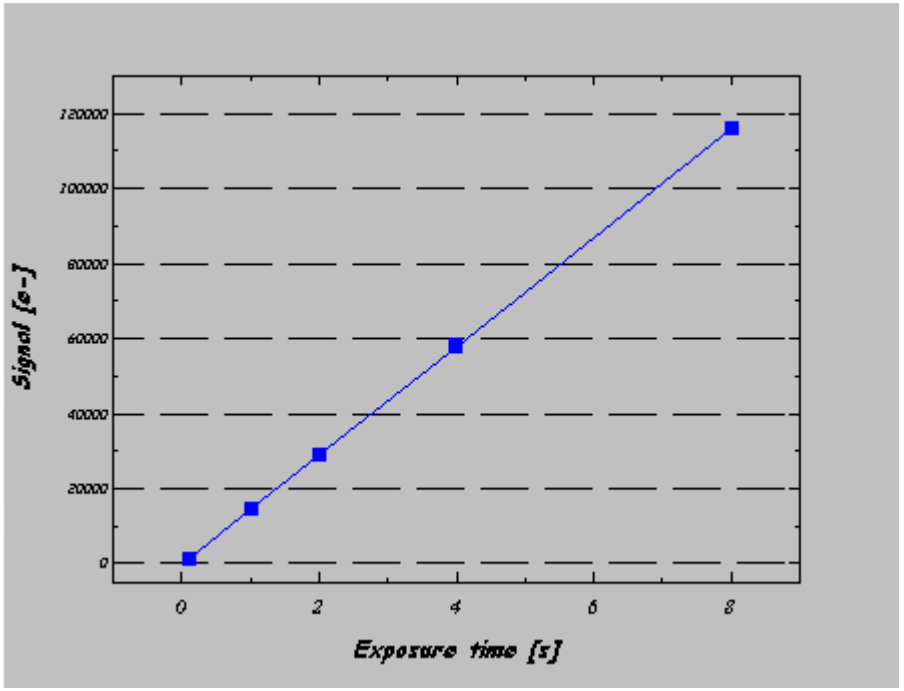
Detector Linearity: The range over which the detector response is linearly related to the stimulus:

$$R = Flux \times time$$

Photographic emulsions are not linear (note log-log scale).



CCDs and Photomultiplier Tubes (PMTs) are linear over a large range.



Detector parameters (3)

- **Dynamic Range:** The ratio of the largest measured value to the smallest value should be as large as possible.
- **Spectral Bandwidth:** The range of wavelength over which the detector is useful.
- **Angular Resolution** (spatial resolution) – describes the ability of any image-forming device to distinguish small details of an object. It should be well matched to the telescope and instrument.
- **Ability to integrate:** The ability to collect photons for an extended period of time.

Detector parameters (4)

- **The response time** is the time a device takes to react to a given input.
- **Digital output:** In order for calculations to be carried out, the data received or recorded by an astronomical detector must be made available as numbers.
- **Noise:** Photon statistics limited noise, Readout noise, electronic noise, dark noise, etc.
- **Cooling.**

Types of Detectors

- **Quantum Detectors:**
  respond to incident photons
  (photoemissive, photoconductive, photovoltaic, photochemical)

- **Thermal detectors:**
  respond to temperature rise due to absorption of radiant energy
  (thermocouple/thermopile, bolometer/thermister, pyroelectric device, Golay cell)

The primary interaction processes

- **Photo effect** is one of the most important interaction processes of photons with opaque matter in photon detectors. A photon with energy $h\nu$, impinging on the surface of a material, can release an electron with a maximal energy of

$$W_{max} = h\nu - W_A ,$$

where W_A is an electron work function characteristic of a given material. Typical values are between $W_A \approx 2$ eV and 6 eV.

The primary interaction processes

- **Solid-state detectors:** In solid-state detectors, with donor and acceptor regions forming n-p junctions, the internal photo effect creates electrons and electron holes that can be collected and counted.

The primary interaction processes

- **Pair Production:** At energies above 1.02 MeV, i.e., twice the equivalent rest mass of an electron, m_e, a photon can produce an electron-positron pair in the electric field of a nucleus. The pair production efficiency increases with the photon energy and the interaction cross-section is proportional to Z^2, with the Z charge number of the nucleus. The electron-positron pairs then have to be recorded by appropriate means.

The primary interaction processes

- **Scintillation counters:** In the X-ray and γ-ray regimes, scintillation detectors are in frequent use. The high-energy photons, through the internal photo effect, ionize atoms in certain crystallic materials. The recombination events of the atoms and electrons then produce a large number of optical photons that can be collected by standard photomultipliers.

The primary interaction processes

- **Photon detectors based on superconductivity.** Many modern photon detectors employ superconductive material. *For example:*
  - **Cooper pairs.** Cooper (1956) suggested that electrons (as fermions) in superconducting metallic materials form pairs—later to be called Cooper pairs—due to an electron-phonon interaction at low temperatures, which could overcome the Coulomb repulsion of the electrons. The resulting bosons would be responsible for the superconductive properties, but might easily split into two electrons by an external excitation. Even photons with an energy of a few millielectronvolts can break up a Cooper pair into the electrons, which then have to be collected by suitable circuitry. This could, for instance, be a superconducting tunnel junction (STJ).

“Ancient” and modern detectors

Photographic emulsion

Photomultiplier

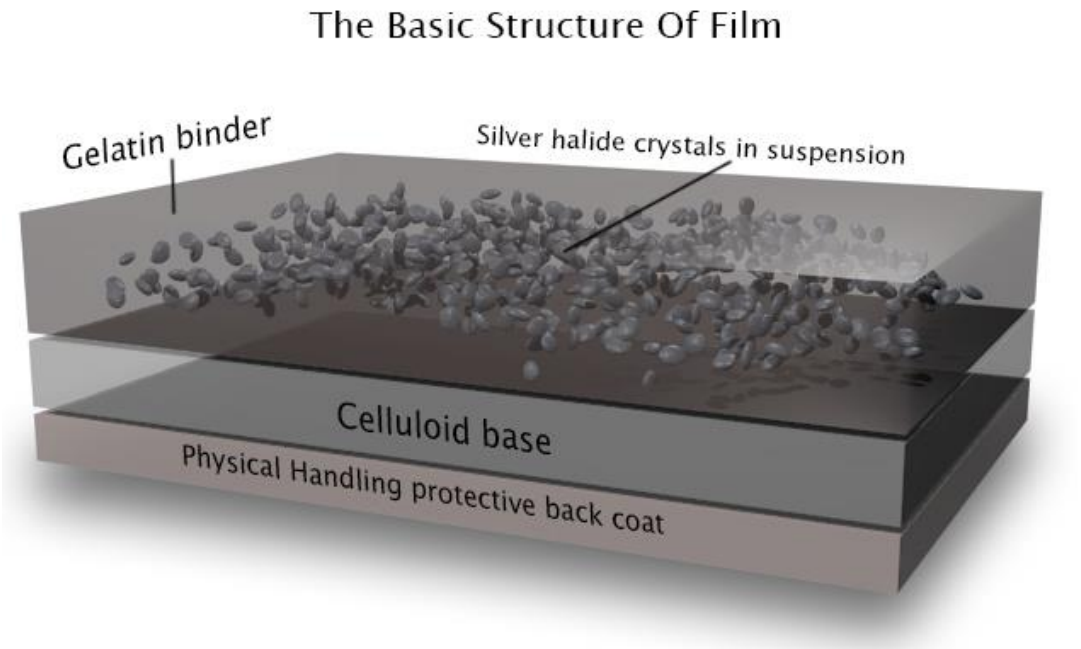
Micro-Channel Plate

Charge Coupled Device detector (CCD)

Photographic emulsions (1)

The first technological advance in astronomical detectors (1840-60).

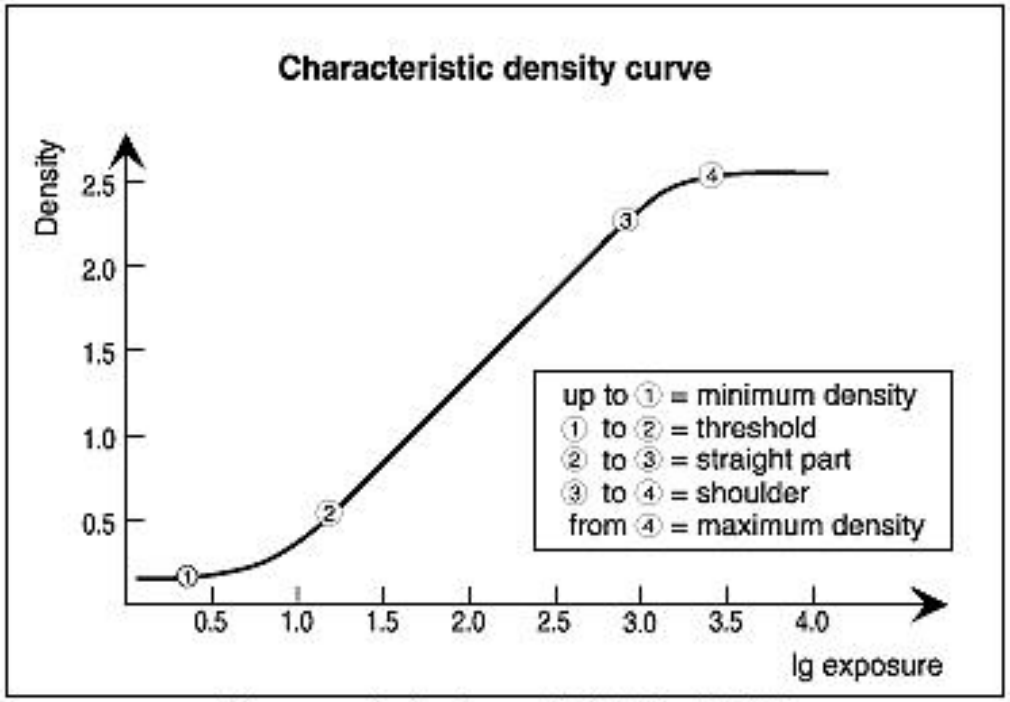
- Typical photographic film / plate contains tiny silver halide crystals suspended in a gelatin.
- When radiation of appropriate wavelength strikes one of the silver halide crystals, a series of reactions begins that produces a small amount of free silver in the grain.
- The free silver produced in the exposed silver halide grains constitutes what is referred to as the "latent image," which is later amplified by the development process.



Thus, photoemulsions can make a permanent record of astronomical objects imaged by telescopes.

Photographic emulsions (2)

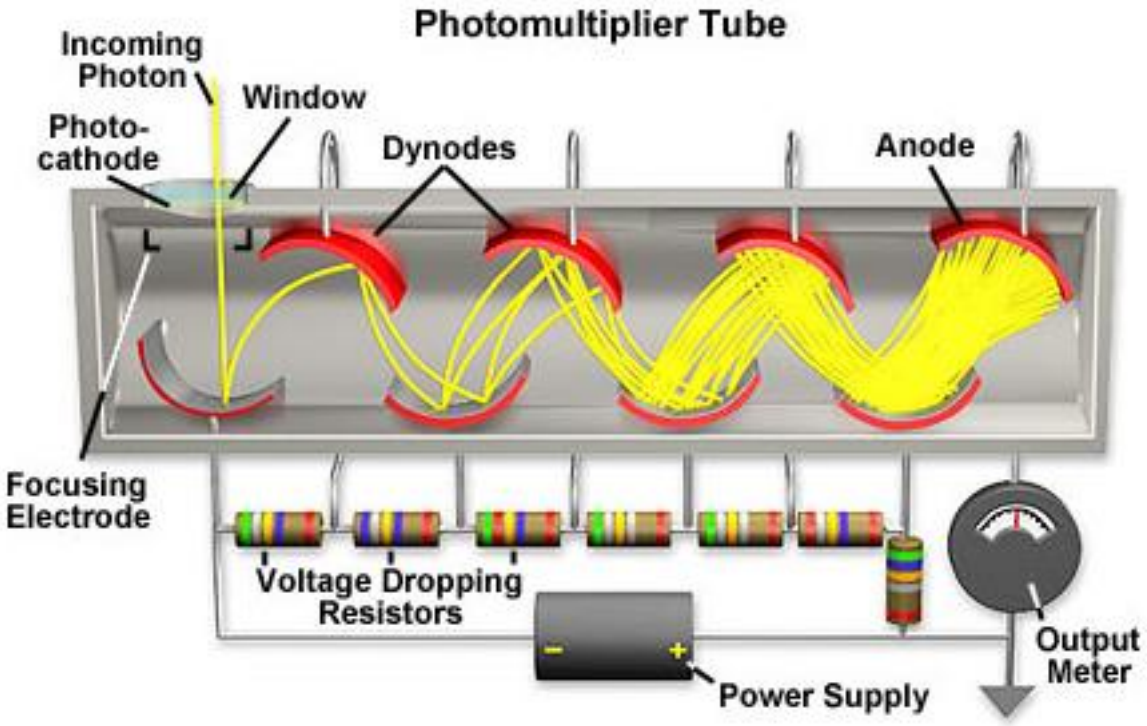
- **Disadvantages:**
  - QE is just a few %.
  - Because of the analog (rather than digital) nature of the image record on an emulsion, it is difficult to make quantitative measurements of star brightnesses.
  - Photoemulsions are also nonlinear with input light.
  - Photoemulsions are also nonlinear with increasing exposure time – an exposure of 2 minutes does not give twice the output of a one-minute exposure.
- One **advantage** of the photographic plate is that it can be made very large (40×40cm).
- Photoemulsions are not used for decades.



Then why do we discuss them? → **Archival Data** (e.g., DSS)!

Photomultipliers (1)

- Photomultipliers detect photons through the photoelectric effect: the absorption of a photon results in the emission of an electron.
- The detectors work by amplifying the electrons generated by a photocathode exposed to a photon flux.
- Once a photoelectron has escaped from the photoemitter, it is accelerated by an electric potential until it strikes a second electron emitter.
- Dynodes are successively more positive than the cathode by ~100 V for each stage.
- The final signal pulse may contain 10^6 electrons for each incoming photon.

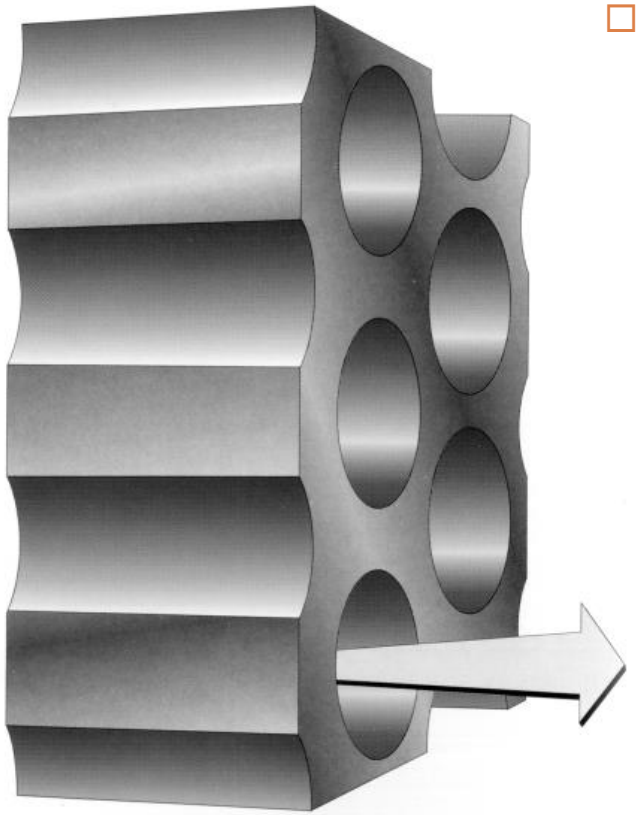


Photomultipliers (2)

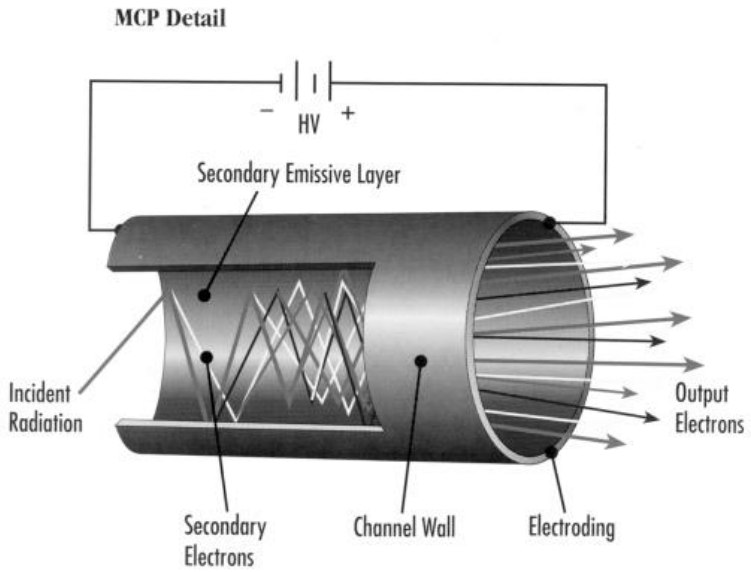
- **Advantages:**
  - Have much higher QE of around 20%.
  - Have a digital output.
  - Are linear with input light and exposure time.
  - Can be used in UV.
- **Disadvantage:**
  - It is essentially a single channel device, there is no positional information in the signal. The output signal does not depend on where on the cathode the photon hit, so we get only a measure of all the light that fell on the photocathode.

- Still in use as a photon detector for the Čerenkov telescopes (e.g., in MAGIC), and on old small optical telescopes.

Micro-Channel Plate (MCP)



- MCPs are basically microscopic photo-multiplier tubes!
- An MCP is simply a matrix of resistive glass tubes constructed using fiber technology. A photon knocks an electron out of the resistive material (lead oxide) coated on the inside of the tube, and is accelerated down the tube by the applied voltage.

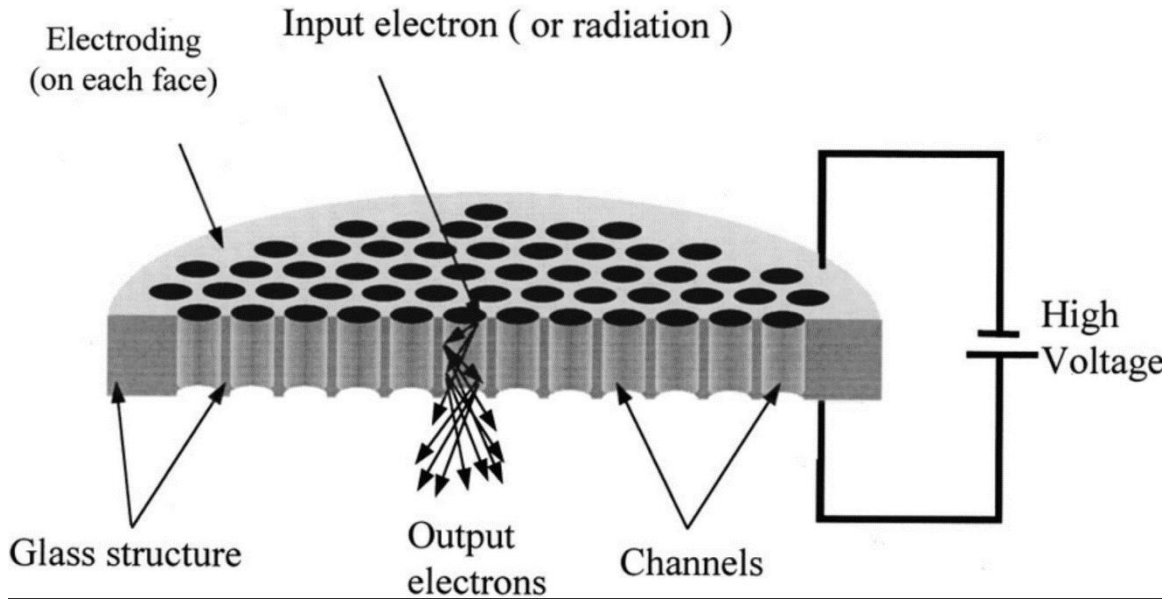


This creates a cascade of events which can then illuminate a detection screen.

Micro-Channel Plate

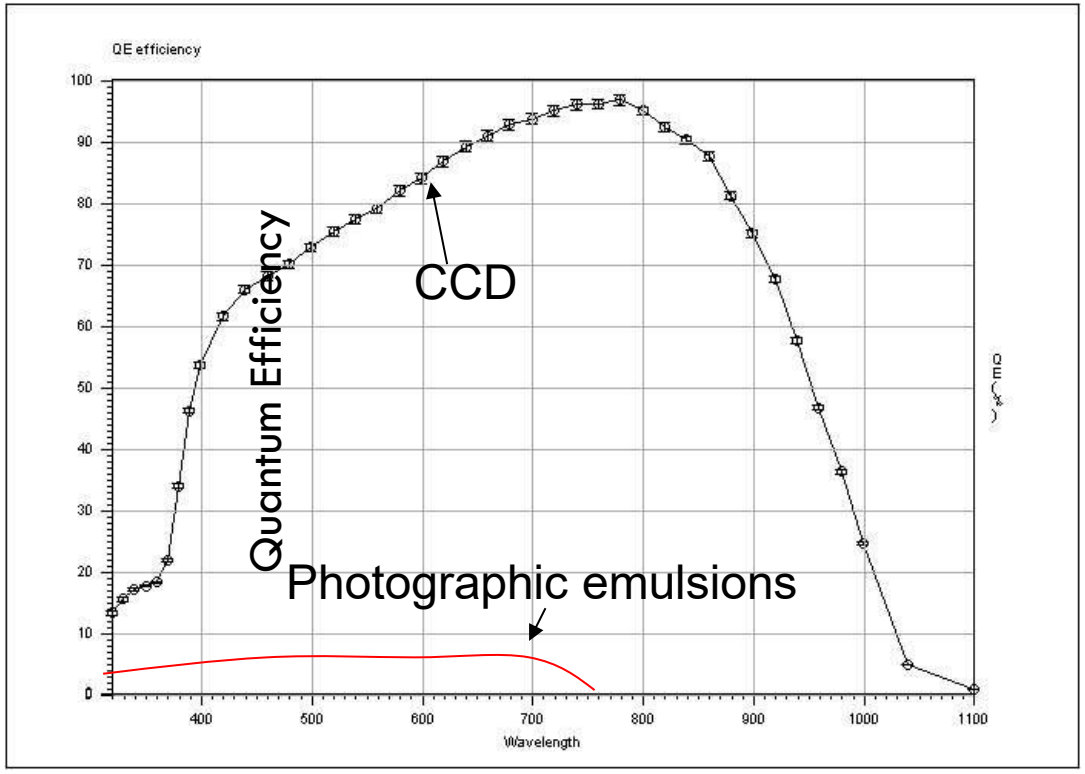
- **Advantages:**
  - QE is up to 50%.
  - Can have excellent spatial resolution and microsecond time resolution possible.
  - Can be used in X-ray, UV and optical wave.

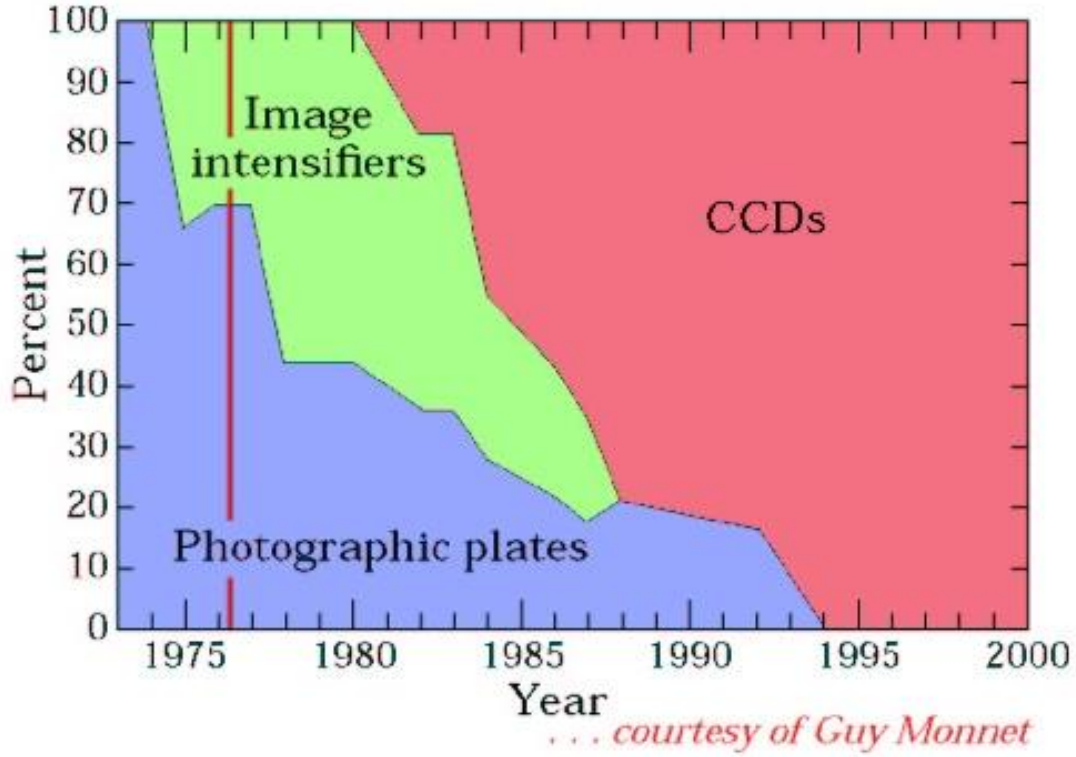
- **Disadvantage:**
  - Lack energy resolution
    (a made-up, not real problem)



Charge Coupled Device detectors (CCDs)

CCDs – a replacement for photoplates: QE is more than an order of magnitude better than photographic plates –
an astronomer's dream for decades!





Advantages of CCDs

- Good spatial resolution
- Very high QE of up to 95% (and even higher)
- Large spectral window:
  By far the most common detector for wavelengths
  4000 Å <λ<10000 Å
- Very low noise
- High photometric precision
- High dynamic range
- Very good linearity
- A reliable rigidity (no physical distortion, etc) –
  robust enough to fly on space missions.

How does a CCD work?

- In order to produce an image, a CCD accomplishes four functions:

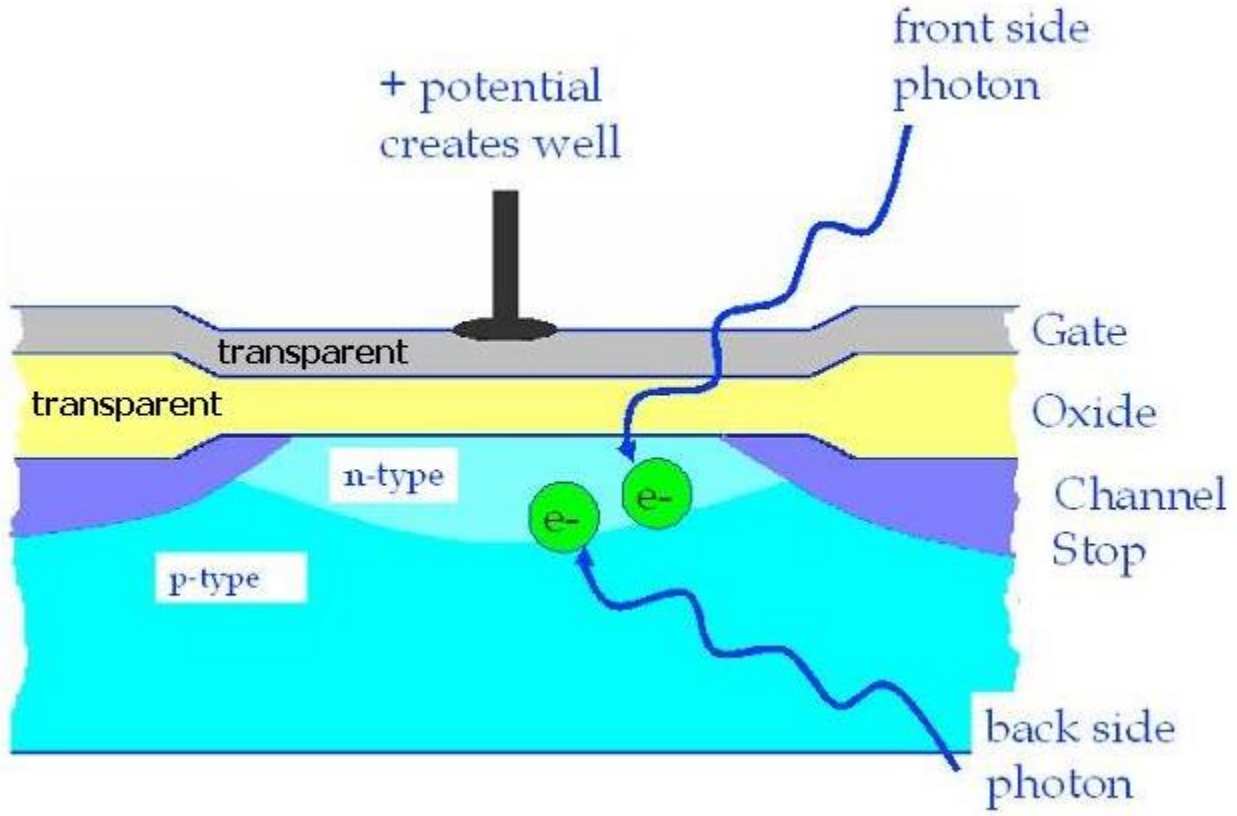
1) generates photoelectrons
2) collects electrons
3) transfers the collected charges
4) reads the charges

How does a CCD work? (2)

- The first function is based on the photoelectric effect.
- A CCD is a silicon wafer which is exposed to radiation.
- The light absorption in the silicate network of the CCD generates these photoelectrons, **in proportion to the number of incident photons**.
- The latter are immediately collected in "picture elements" so-called pixels, closest to where the photons fell on the chip. The picture elements (pixels) of the CCD are defined and fixed by the electrode structure which is applied to this wafer.
- The electrodes form potential wells to prevent the collected charges from escaping.

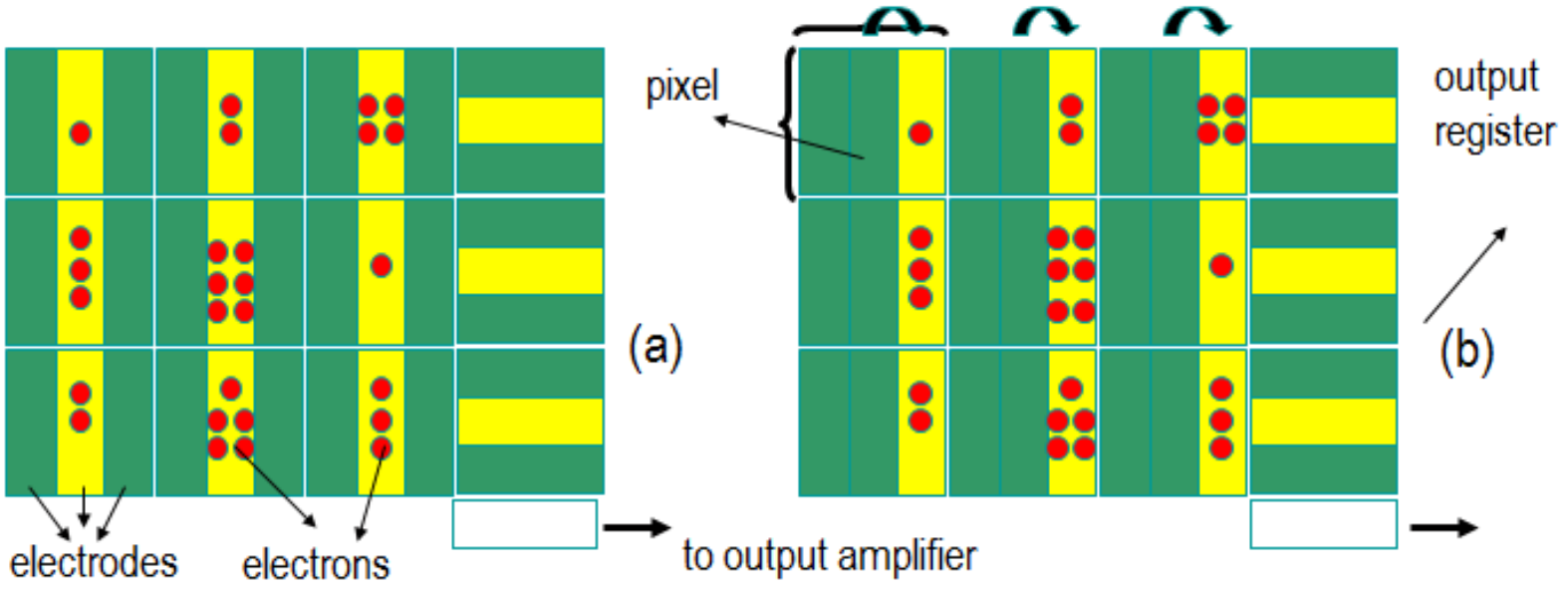
How does a CCD work? (3)

- As you expose the CCD to radiation, electron-hole pairs are generated, and the electrons build up in the electron storage areas immediately below the positive potential electrodes.
- After a while (exposure time – seconds to minutes) the CCD contains an electrostatic representation of the pattern of incident radiation on it.
- This somehow must be read out and stored in digital form.



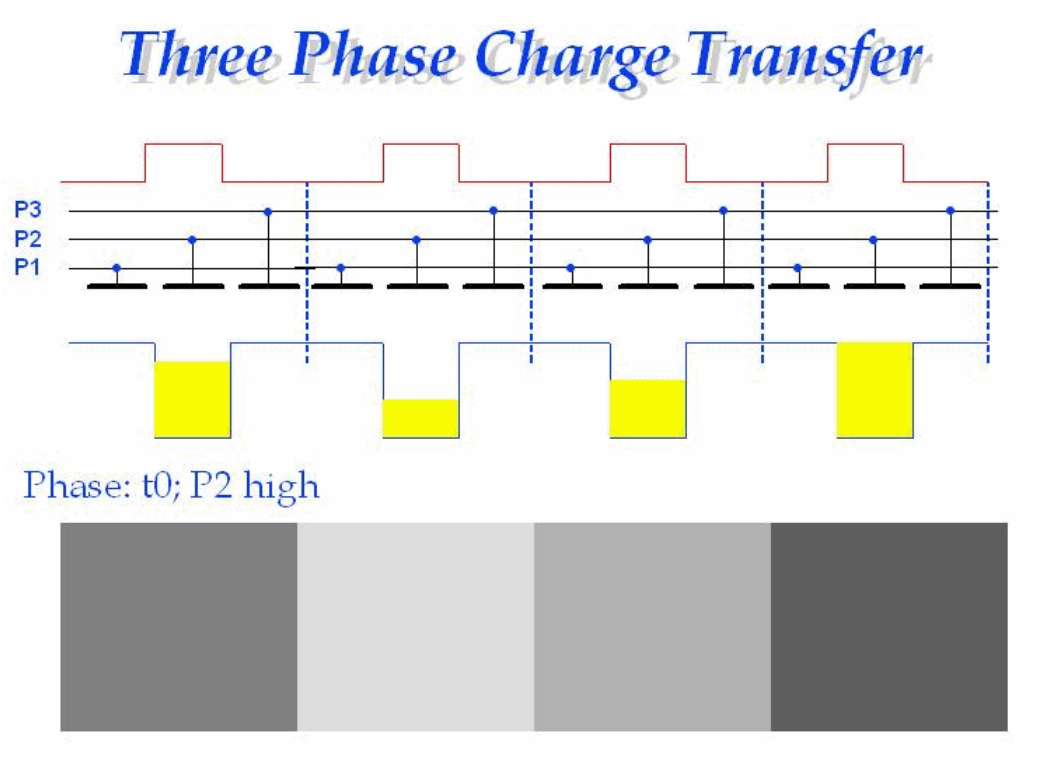
How does a CCD work? (5)

- Each pixel is divided into 3 regions (electrodes who create a potential well). For the charge collection process during an exposure the central electrode of each pixel is maintained at a higher potential (yellow) than the others (green).



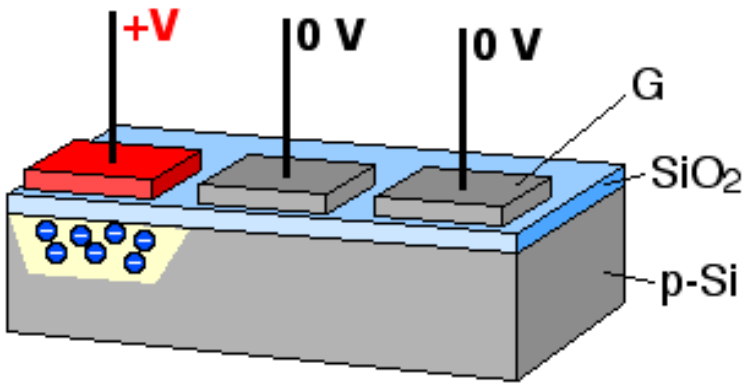
How does a CCD work? (6)

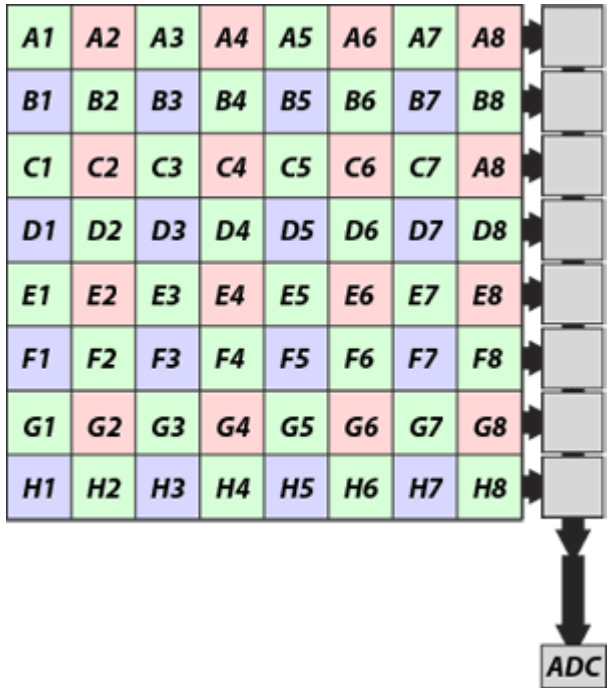
- By changing the potential of the electrodes in a synchronized way, electrons are transferred from pixel to pixel. Charges on the right are guided to the output register



How does a CCD work? (7)

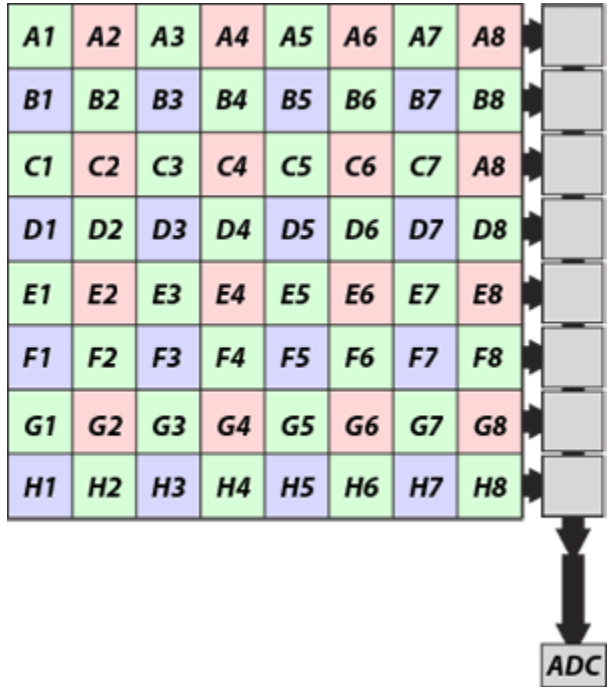
- The readout register is shifted to the right by one pixel, and the pixel at the bottom right is shifted into a readout capacitor.





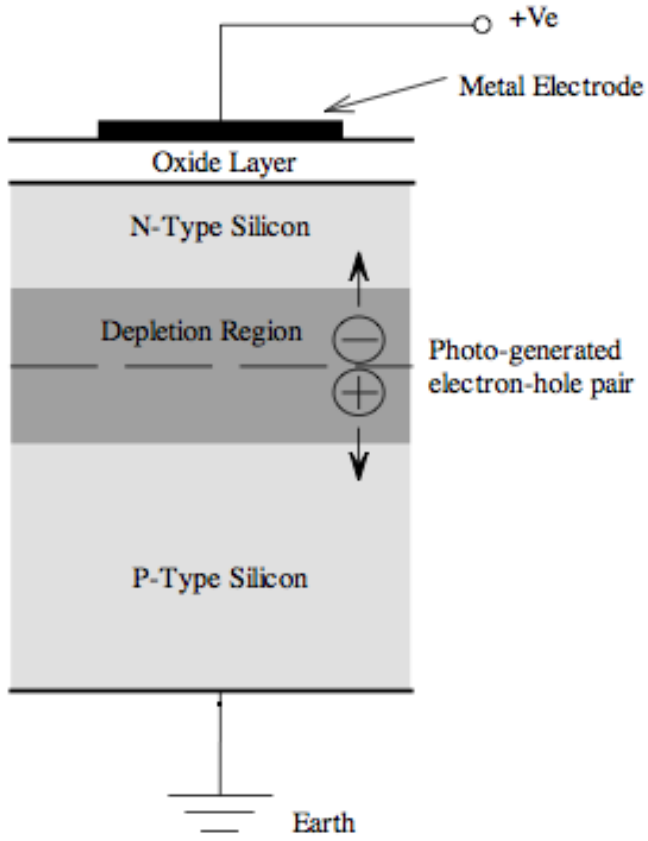
CCD readout

- This process is then repeated until each pixel in the readout register has been digitized.
- The image is then shifted right horizontally by one more pixel, the next collumn is shifted into the readout register, and this is digitized in the same way.
- The whole process is repeated until the entire image is read out.
- For a 2048 x 2048 pixel CCD it takes approximately 10-60 seconds to read out the whole chip.
- A CCD read out this way is a line transfer CCD.
- The channel stops between rows are permanent as charge does not move vertically except in the readout register.
- These channel stops are biased to negative potential by doping, hence charge cannot leak horizontally.



Buried channel CCD

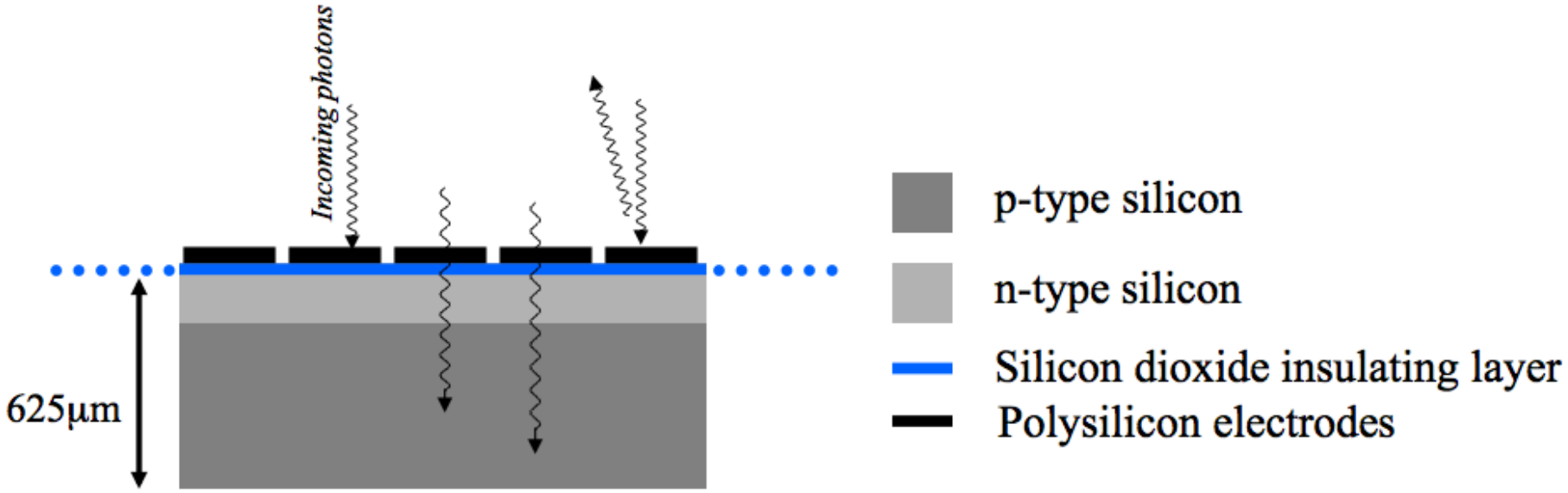
- CCDs described before are surface channel CCDs, but in these the charge is being shifted along in a thin layer just below the oxide insulator.
- Surface layer has crystal irregularities which can trap charge, causing loss of charge and image smear.
- By grounding the p-type silicon substrate and applying a positive voltage to the n-type layer, the p-n junction is reverse biased, widening the depletion region and creating a potential well in the n-type silicon that confines the electrons in the vertical direction
- This is called a buried-channel CCD, and suffers much less from charge trapping.
- Thus, photoelectrons are collected deep in the well beneath the positive potential set up by the gate electrode, isolated from the surface by the n-type silicon channel.
- Almost **all** scientific CCDs are buried-channel devices.



Courtesy of Vik Dhillon

Front-illuminated CCD

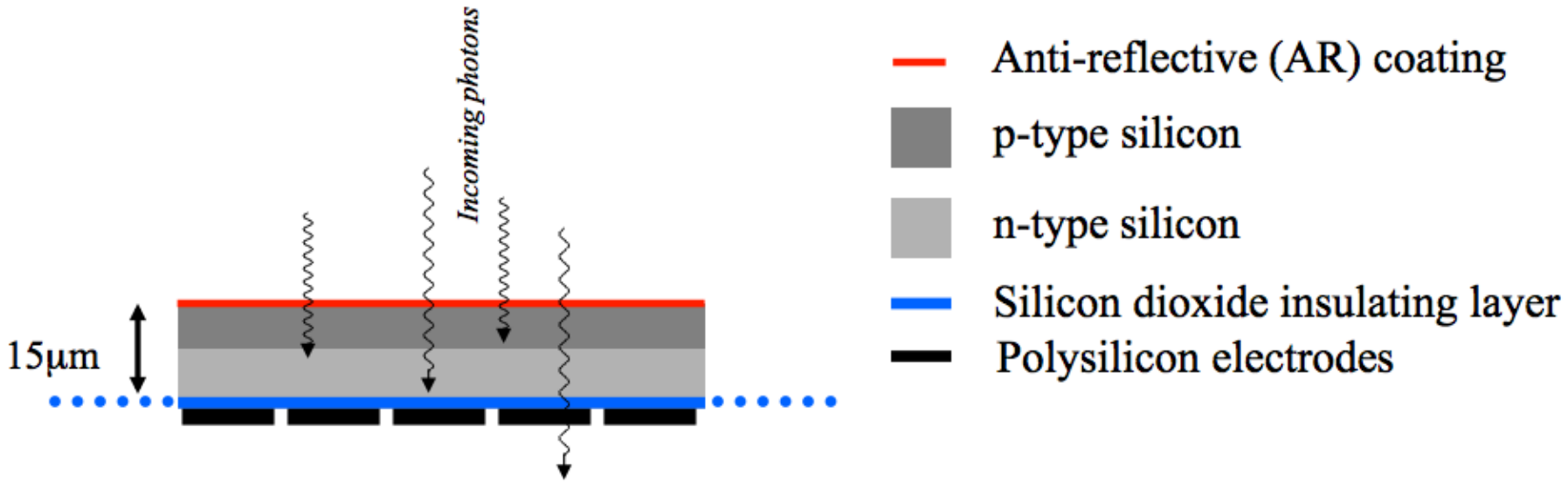
- As described to now, the CCDs are illuminated through the electrodes. Electrodes are semi-transparent, made of a transparent material known as polysilicon, but some losses occur: the electrode structure absorbs and reflects many of the incident photons, particularly at blue wavelengths, preventing them from producing electron-hole pairs in the depletion region.
- They are non-uniform losses, so the sensitivity will vary within one pixel.



Courtesy of Vik Dhillon

Thinned back-illuminated CCD (1)

- Solution is to thin the CCD, either by mechanical machining or chemical etching, to about 10μm, and mount it the other way up, so the light reaches it from the back.
- Thinning is a way of improving sensitivity, especially at blue wavelengths.

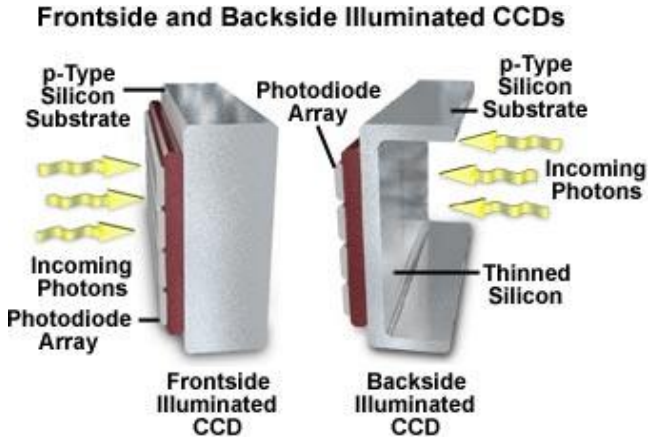


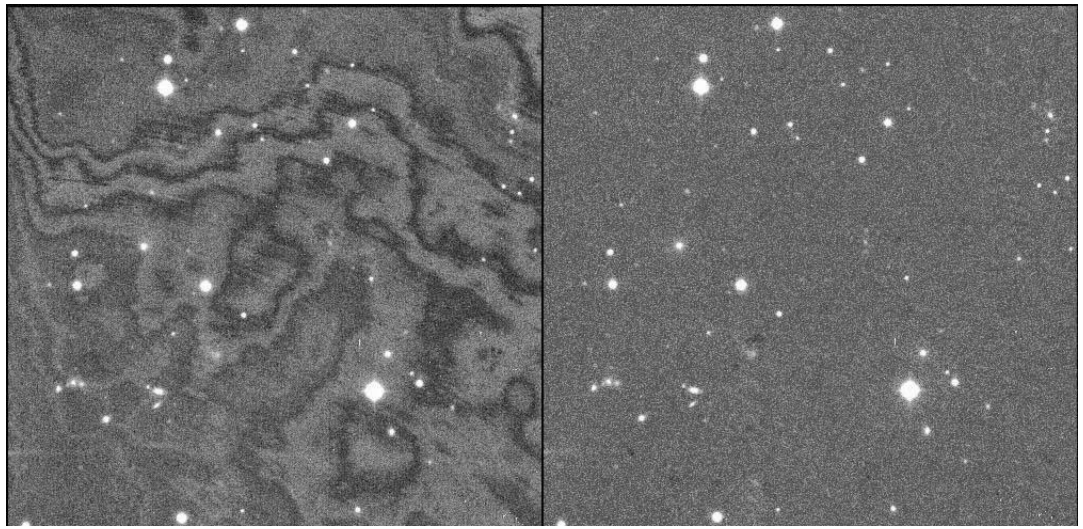
Courtesy of Vik Dhillon

Thinned back-illuminated CCD (2)

Thinned back-illuminated CCDs do **have** a number of disadvantages:

- Thinning can reduce the red response because red photons need more absorption length and if this is not there they will pass right through the silicon.
- Thinned CCDs are also mechanically fragile, prone to warping and expensive to manufacture compared to thick CCDs.
- Another problem is that the thin silicon layer produces interference fringing in the red part of the spectrum, which can limit the accuracy of measurements.





Frame transfer CCDs

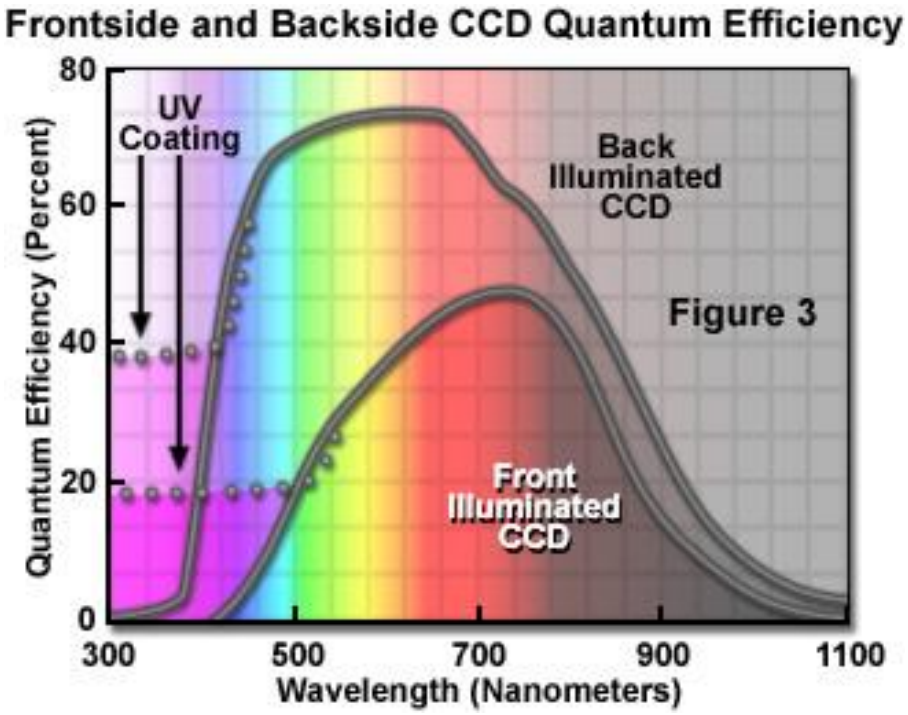
- Instead of reading the CCD out line by line as described before, a Frame transfer CCD has half of its area masked off to stop light reaching it. On readout, the whole CCD is clocked vertically so that the image area is transferred to the storage area.
- The image can then be read out from this storage area whilst the image area is being exposed again.

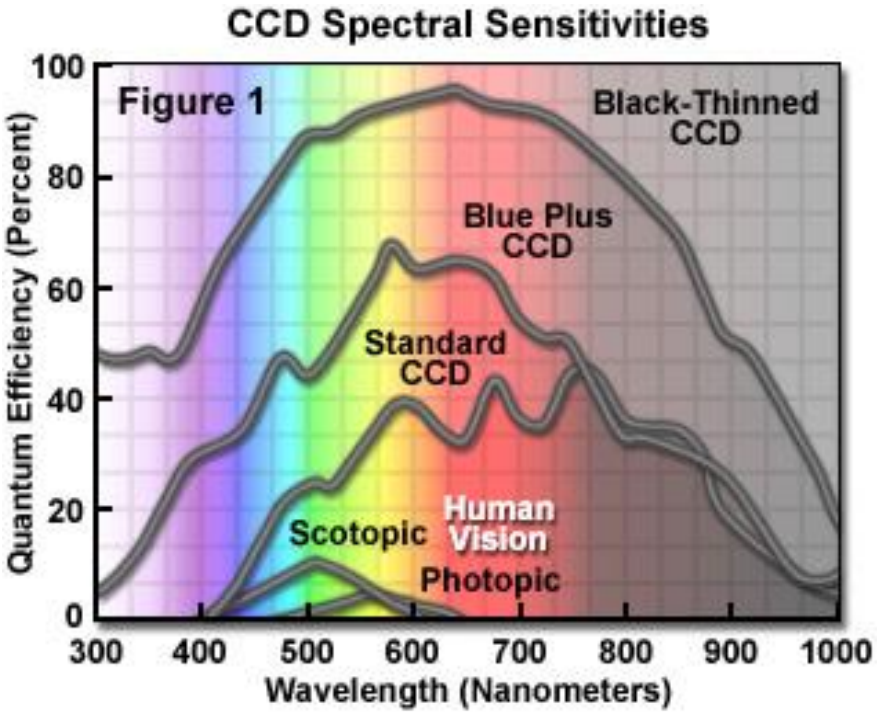
Advantages of CCDs (2)

- **Good spatial resolution:**
  - Today, most common CCDs have 2048 × 2048 pixels. But there exist even larger CCDs with 4096 × 4096 pixels or 4096 × 8192 pixels (10 k x 10 k will be build soon).
  - For realizing even larger chips (and since larger CCD chips are very expensive), several small chips can be placed together resembling a CCD mosaic.

Advantages of CCDs (3)

- **Quantum Efficiency:**





Advantages of CCDs (4)

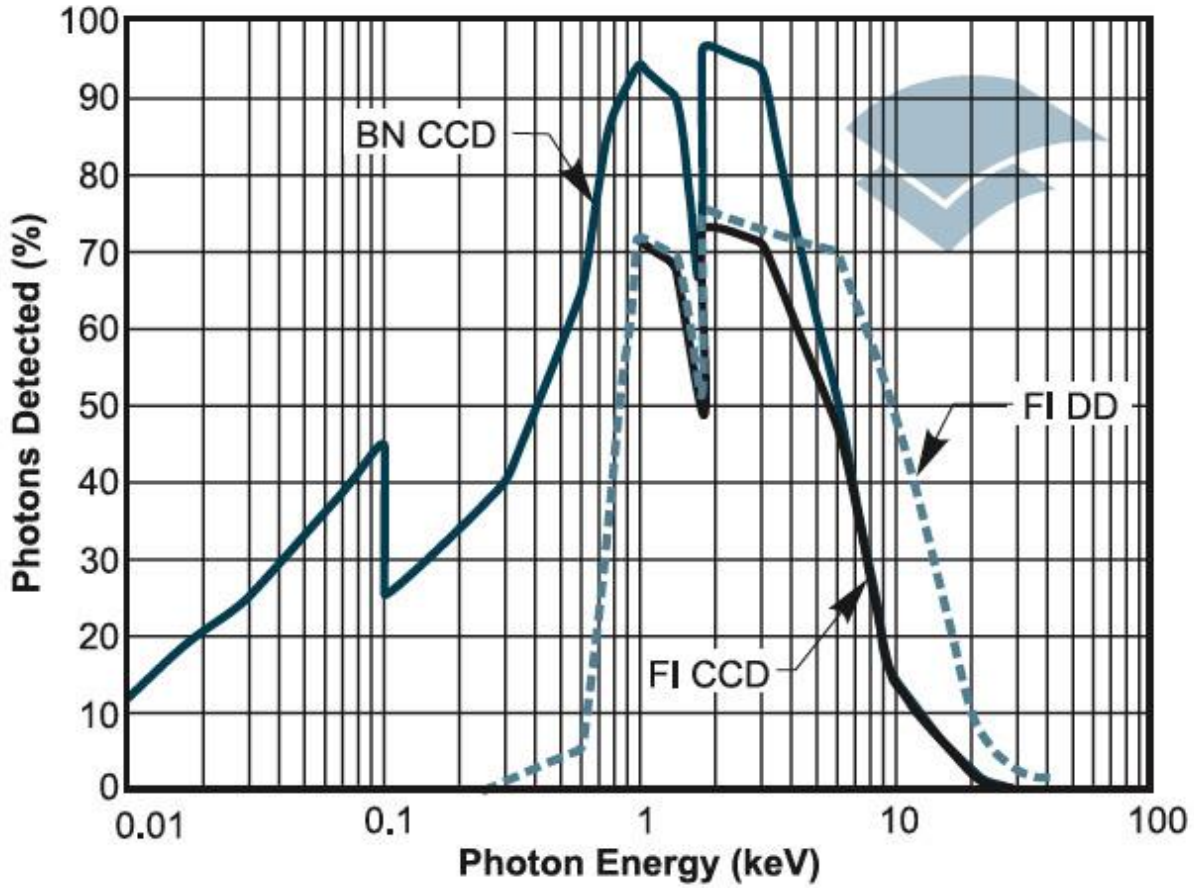
Spectral Range

Goes to X-rays

FI: front illuminated

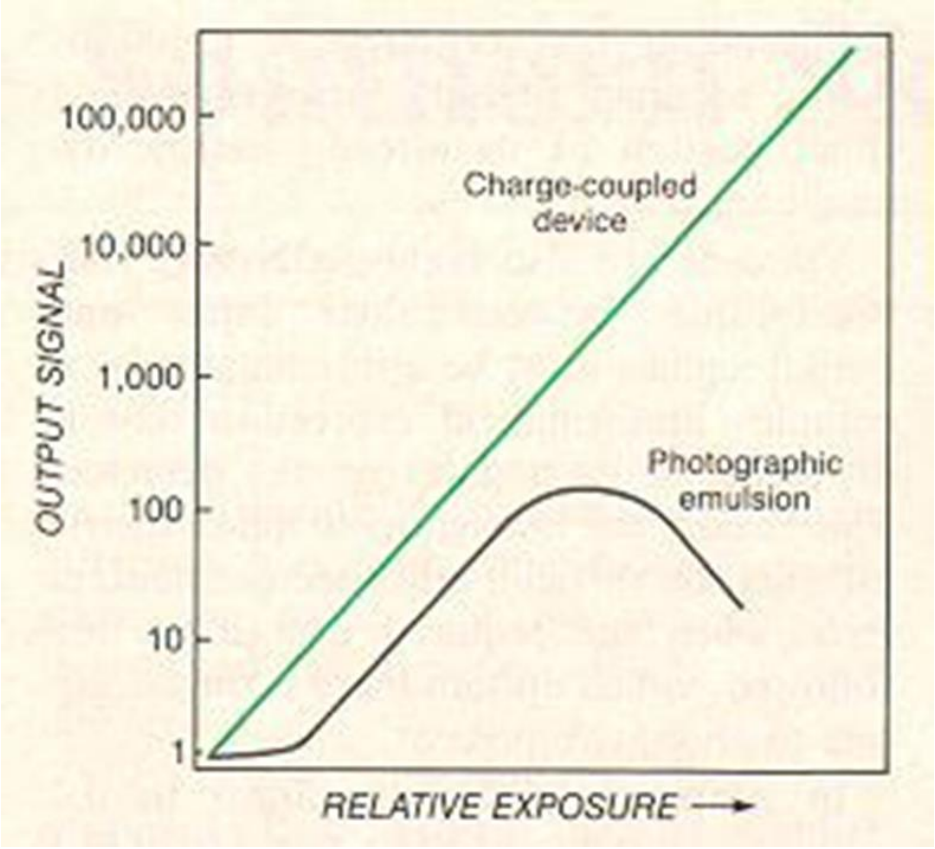
BN: back illuminated, no coating

DD: deep depletion



Advantages of CCDs (5)

- **Linearity and Dynamic Range:** CCDs are extremely linear detectors. Therefore CCDs enable the simultaneous detection of both very faint and very bright objects. The dynamic range of CCDs is about 100 times larger compared to photoemulsions.



Disadvantages of CCDs (1)

- **Size:** The size of a single pixel is in the order of 8 × 8, 15 × 15 or 25 × 25 microns. Therefore the size of CCD chips remain quite small, especially by comparing CCDs to classical photographic plate images. E.g., a CCD with 2048 × 2048 pixels of 15 microns measures only 3 × 3 cm^2. In contrast a photographic plate for a Schmidt telescope can be as big as 30 × 30 cm^2, equivalently to a CCD chip with 400 million pixels!

Disadvantages of CCDs (2)

- **The dark current** (not a real problem) is background signal generated by thermal effects. Because of the dark current CCDs are run cooled, to reduce the possibility of thermal excitation of electrons across the band gap.
- CCDs are operated at temperatures of around 140K, to reduce thermal effects.
- Dark current at 140K is typically 10^{-4} electrons/s/pixel, i.e. negligible.

Disadvantages of CCDs (3)

- **Cosmic rays, X-rays, and particle radiation:**
- There are a number of types of radiation which can interact with the silicon to produce several tens of electron-hole pairs in a cluster, which appears as a bright spot (if the radiation is normal to the detector) or a streak if it is steeply inclined. These radiation events are:
  - Secondary muons in cosmic ray air showers.
  - X rays emitted by UV transmitting glass in the optics of the instrument.
  - Radioactivity from heavy metal impurities in the cryostats.
- These events are identified, classified and rejected by splitting the CCD exposure into two or more equal parts, the hits don't occur in the same place.

Disadvantages of CCDs (4)

- **Saturation** (not a real problem): Typically the full well capacity of a CCD pixel 25 µm square is 500,000 electrons. If the charge in the well exceeds about 80% of this value the response will be non-linear. If it exceeds this value charge will spread through the barrier phase to surrounding pixels.
- This charge bleeding occurs mainly horizontally, as there is little vertical bleeding because of the permanent doped channel stops.
- Readout register pixels are larger, so there is less saturation effect in the readout register.

Disadvantages of CCDs (5)

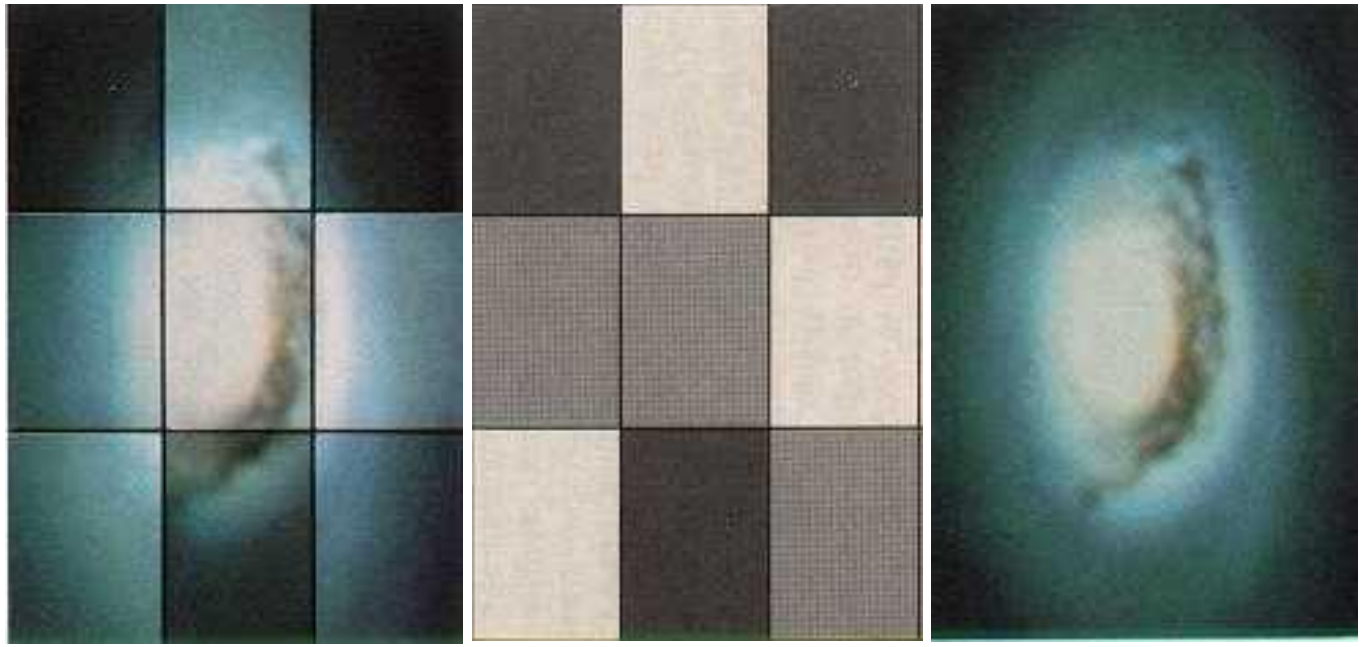
- **Charge Transfer Efficiency:** When the wells are nearly empty, charge can be trapped by impurities in the silicon. So faint images can have tails in the horizontal direction.
- Modern CCDs can have a charge transfer efficiency per transfer of 0.9999995, so after 2000 transfers only 0.1% of the charge is lost.

CCDs: readout noise

- CCDs suffer from **readout noise** which has a variety of sources:
  - The output Field Effect Transistor. This is the ultimate limit to the readout noise, at a level of 2-3 electrons.
  - Transfer loss fluctuations. During transfer an amount of charge is left behind, but this amount varies. Transfer noise is given by: $\sigma_{tr} = \sqrt{(2\zeta n N_0)}$ where $\zeta = 1\text{-CTE}$ is the fraction of charge not transferred, n is the number of transfers and N_0 is the original charge. For faint sources (≈100 electrons) this noise is less than 1 electron.
- The readout noise is the dominant source of random noise.

CCDs: Other noise sources

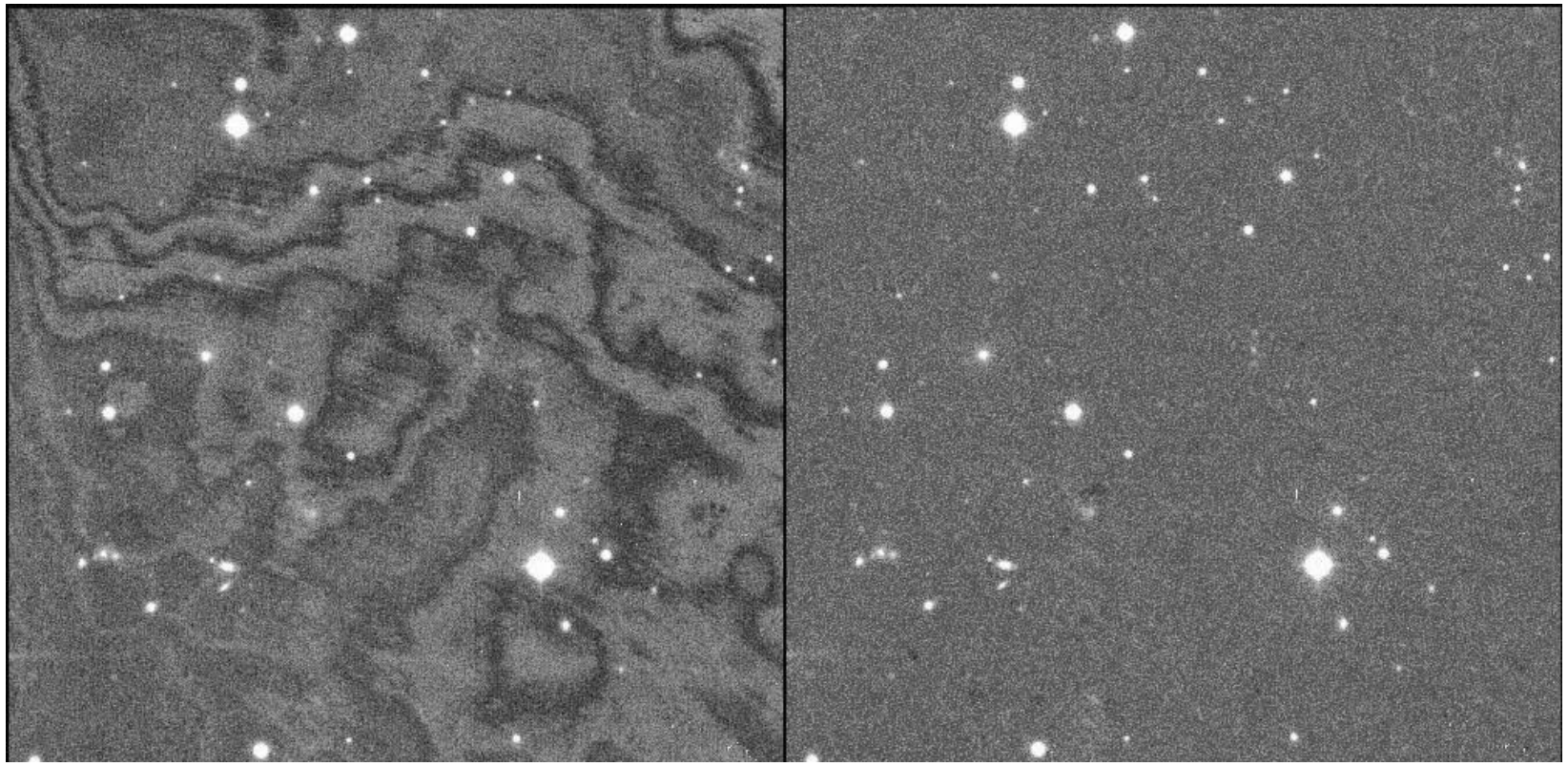
- **Fixed pattern noise.**
  The sensitivity of pixels is not the same, for reasons such as differences in thickness, area of electrodes, doping.
  However these differences do not change, and can be calibrated out by dividing by a flat field, which is an exposure of a uniform light source (**Flat field technique**).



- **Bias noise.**
  The bias voltage applied to the substrate causes an offset in the signal, which can vary from pixel to pixel. This can be removed by subtracting the average of a number of bias frames, which are readouts of zero exposure frames. Modern CCDs rarely display any fixed pattern bias noise.

CCDs: Interference Fringes (1)

- In thinned CCDs there are interference effects caused by multiple reflections within the silicon layer, or within the resin which holds the CCD to a glass plate to flatten it.



CCDs: Interference Fringes (2)

- These effects are classical thin film interference (Newton's rings).
- Only visible if there is strong line radiation in the passband, either in the object or in the sky background.
- Visible in the sky at wavelengths > 7000Å.
- Corrected by subtracting off a scaled exposure of blank sky.
- Fringing can dominate the noise in the redder photometric bands, or in narrow bands, and can sometimes force us back to using thick CCDs despite the loss in QE.



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 6

Vitaly Neustroev

X-ray Detectors

Proportional counters

X-ray CCDs

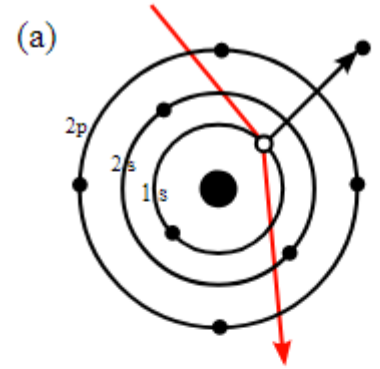
Microchannel plates

X-ray detectors

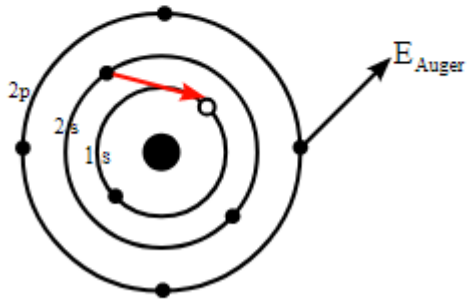
- X-ray photons are few, but are very energetic.
  X-ray detectors can measure the energy as well as
  the spatial position of X-ray photons,
  by counting the secondary photoelectrons they produce.
- X-ray detectors are chosen to have good spatial resolution
  as well as some level of intrinsic energy resolution.
  The three types in serious use today are:
  - Proportional Counters
  - CCDs
  - Microchannel Plates

Auger effect

- The principles that X-ray detectors work on are all pretty similar. The primary photon is very energetic, and produces a number of secondary electrons by the Auger effect.
- The **Auger effect** is a physical phenomenon in which the transition of an electron in an atom filling in an inner-shell vacancy causes the emission of another electron.
- The number of electrons released is proportional to the energy of the incident photon. In the gas proportional counter these electrons are produced by interaction of a photon with an atom of inert gas.



Electron collision
or ionization by an X-ray photon



Auger electron emission

Geiger counters

- Two electrodes inside an enclosure are held at such a potential difference that a discharge in the medium filling the enclosure is on the point of occurring.
- X-ray ionizes gas. Produced electrons accelerate toward anode and cause further ionization, producing more electrons – electron avalanche.
  The amplification factor (gain) is about 10^8.
- Amplifier records a charge pulse.
- Operates in saturated regime:
  no energy resolution.

An analogue of photomultipliers:

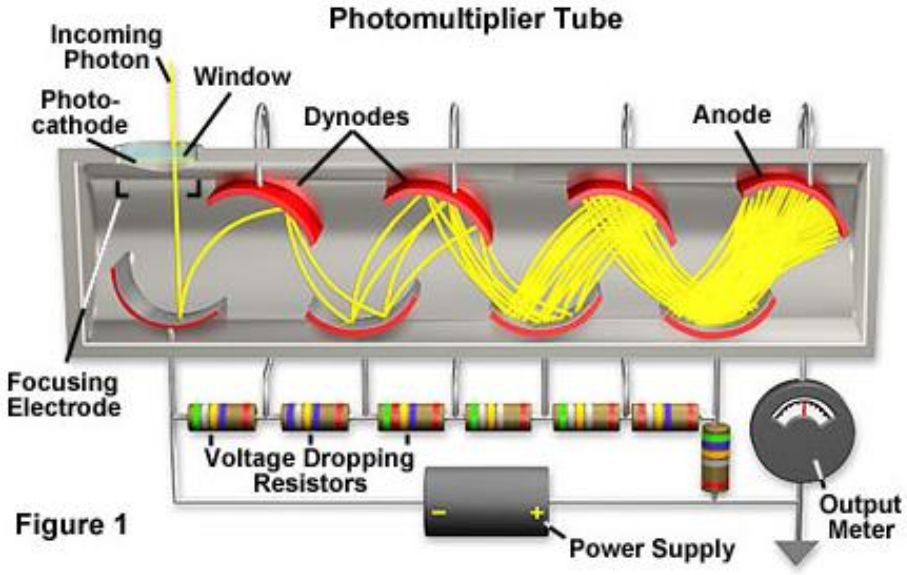
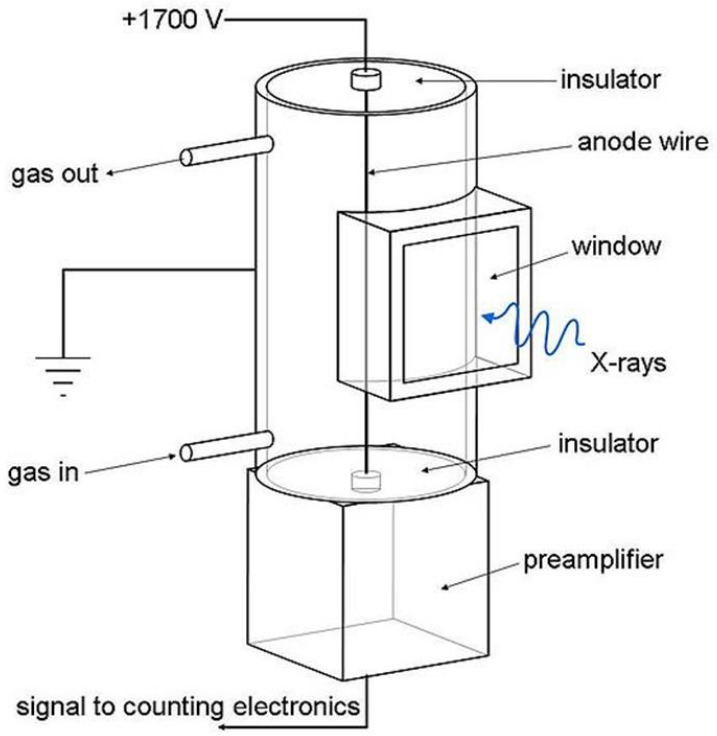


Figure 1

Proportional counters (1)



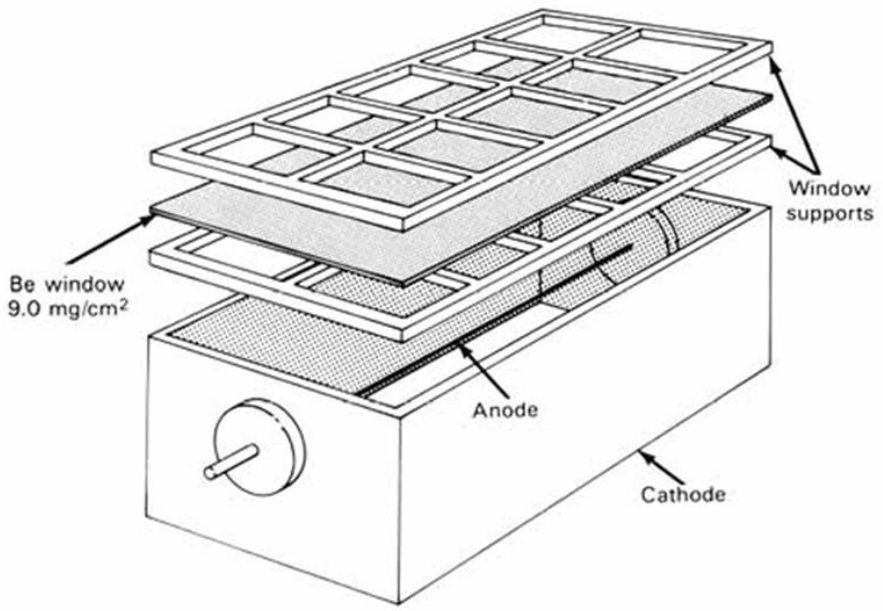
- Proportional counters are very closely related to Geiger counters but they operate at less than the trigger voltage.
- By using a lower voltage, saturation of the pulse is avoided, and its strength is then proportional to the energy of the original interaction.
- The gain is reduced to about $10^4 - 10^5$

Proportional counters (2)

- X-rays are detected through their interaction with inert gas (e.g Argon) in a windowed chamber.
- Primary photoelectron is emitted by the photoelectric effect.
- Secondary "Auger" electrons are emitted in a localized cloud.
- Mean number of electrons released is **N = E/w**, where E is the energy of the X-ray photon, and w is the ionization energy of the gas (w = 26.2 eV for Argon; 21.5 eV for Xenon).
- **Disadvantage:** other forms of ionizing radiation (electrons, positrons, protons, etc.) create false signals, and so do even lower energy photons in the FUV or EUV.

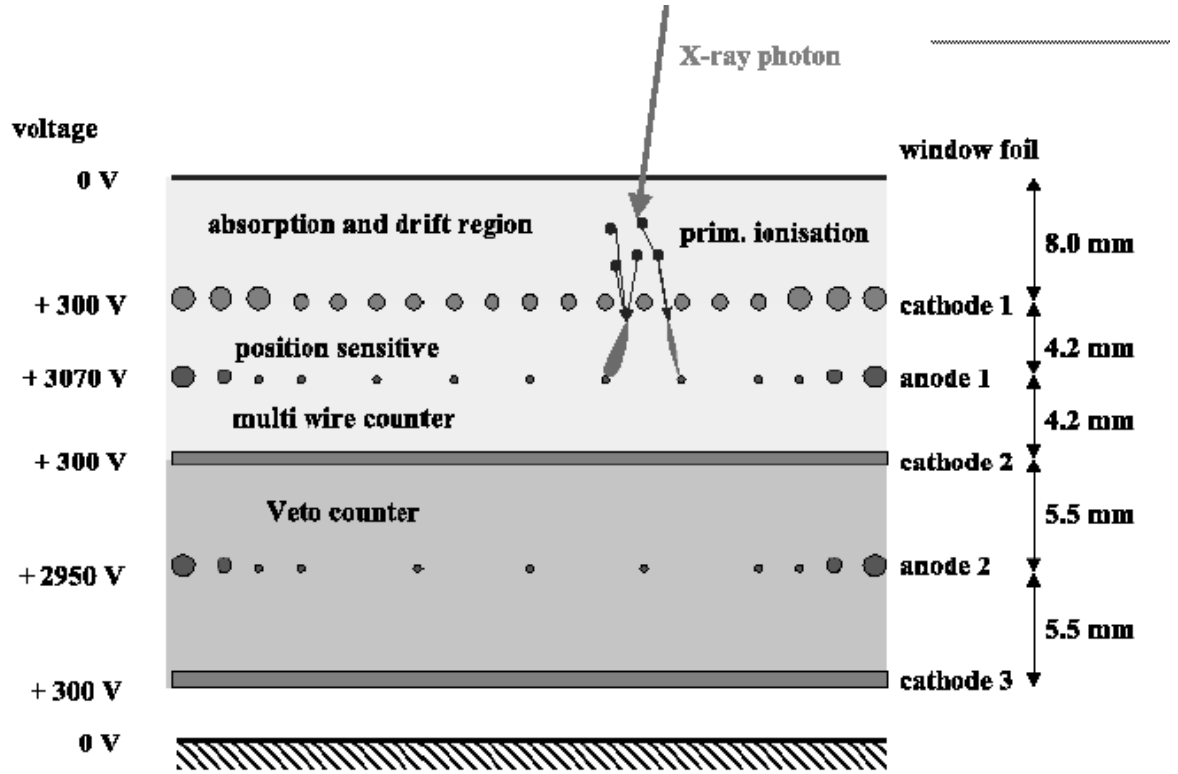
Position Sensitive Proportional Counter (1)

- Position information is obtained using a position sensitive (resistive) anode. The pulse is extracted from both ends of the anode, and from the size of the pulse at each end the location is determined.
- A grid of anodes provides two-dimensional imaging.
- Alternatively, the position can be measured using a pair of crossed cathode grids above and below the anode.



Large Area Proportional Counters

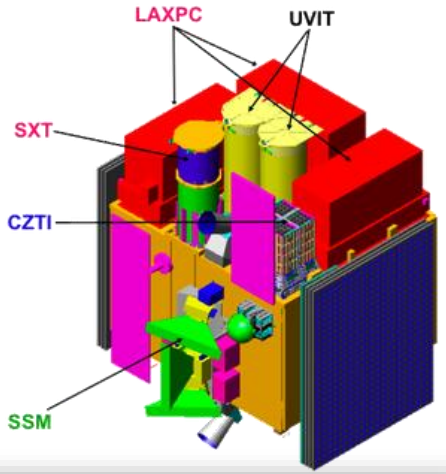
- The observation of the weak X-ray fluxes with non-imaging instruments (no telescopes, only mechanical collimators!) requires large area detectors with high background rejection capability.

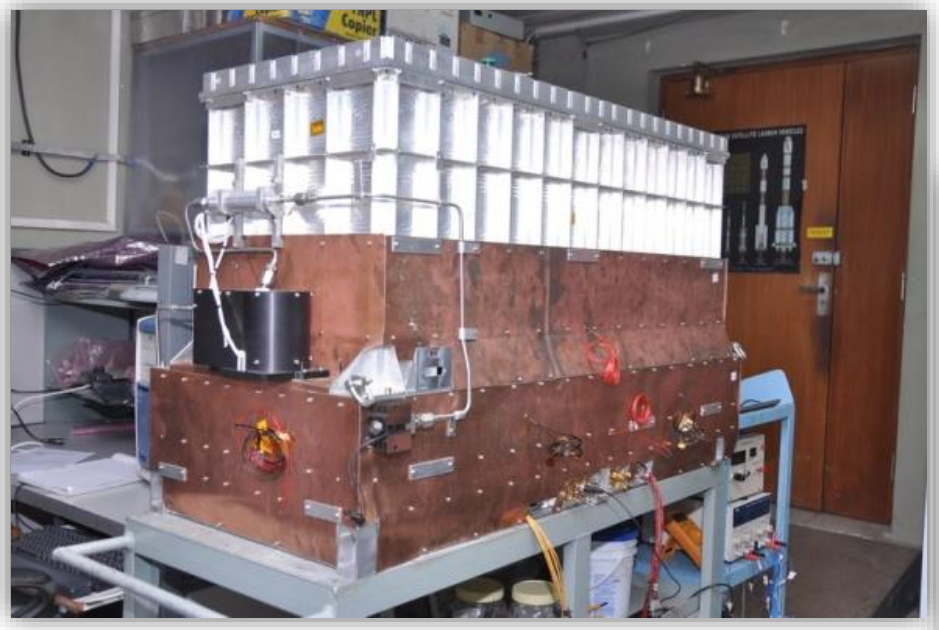


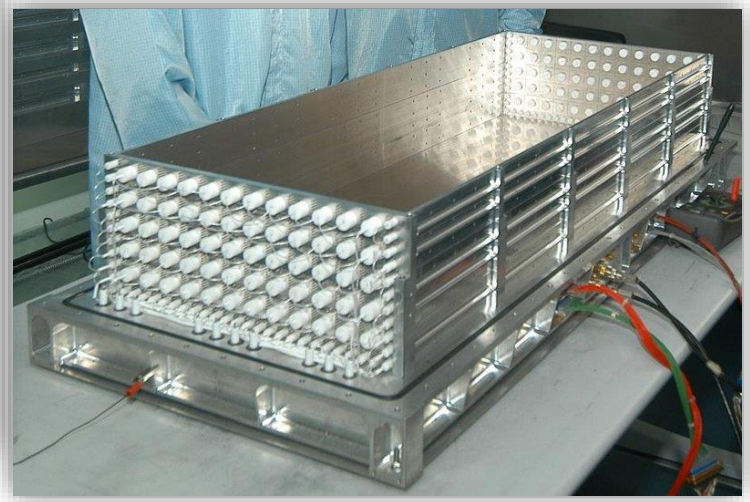
The bottom chamber, the "Veto counter" is used to detect/eliminate cosmic ray events.

Large Area X-ray Proportional Counter (LAXPC) instrument onboard ASTROSAT

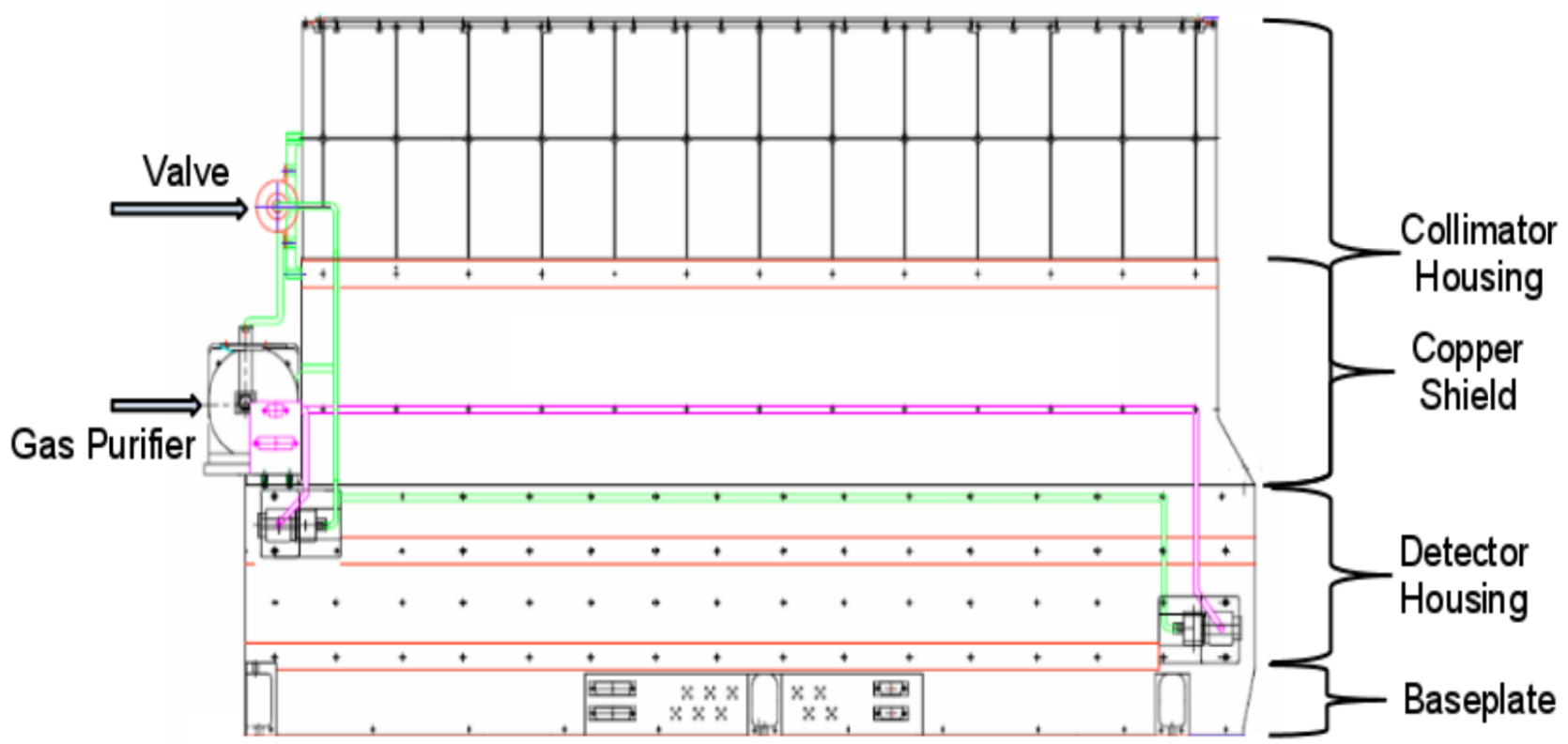
- Astrosat is India's dedicated multi-wavelength space telescope.
  It was launched on 28 September 2015.
  The scientific payload contains six instruments.
- LAXPC - a cluster of 3 co-aligned identical
  Large Area X-ray Proportional Counters (3–80 keV).







LAXPC



A Detector Size 120cm x 50cm x 70cm
X-ray detection volume 100cm x 36cm x 15cm

Proportional counters (3)

- The variation of the number N of electron-ion pairs created by the ionizing event is less than that estimated from Poisson statistics, because the collisions of the ionization process are **not statistically independent**.
- The Fano factor ***F*** is an empirical constant to adapt the experimental observed variance to the predicted one:

$$\sigma_N^2 = F\,N$$

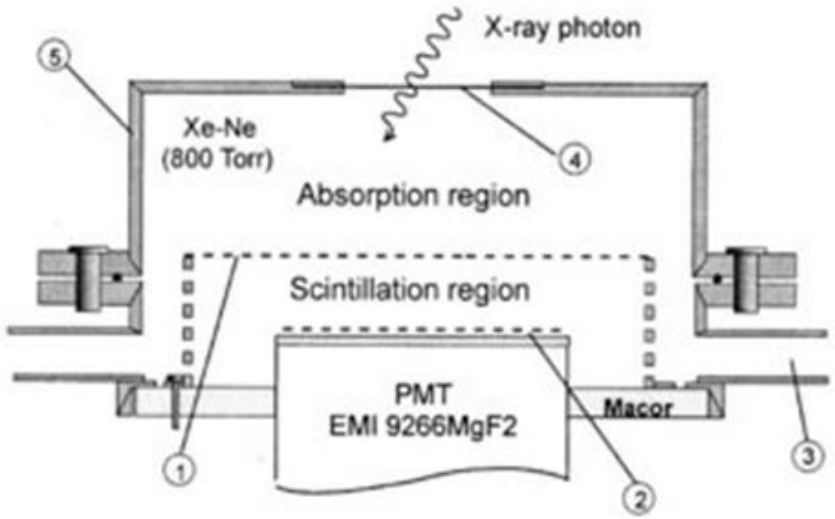
- F is a property of the gas, for Argon and Xenon $F \sim 0.17$.
- Energy resolution: $\frac{\sigma_E}{E} = \sqrt{\frac{w(F+b)}{E}}$,

  where w is the ionisation energy of the gas (~20 eV),
  b= 0.5 – 0.6 (related to the amplification of a single electron)

Proportional counters (4)

- Time resolution:
  - due to the low flux rate of most X-ray sources, each photon event can be time-tagged to high precision.
  - There is, however, a limit to your temporal resolution called "dead-time", a period when the signal from any incoming X-ray photons is essentially lost.
  - There is some time interval required for the system to return to its nominal state – such as the high voltage grid to return to normal voltage, and the ions to recombine, etc. For the typical PC this is a few of microseconds.
- **QE** approaches 100% for energies up to 50 keV

Gas Scintillation Proportional Counters (1)



The number of scintillation photons increases linearly with the number of exciting collisions of the electrons with the gas atoms.

- In conventional PCs, the charge generated by a high-energy photon is multiplied in a high electrical field.
- In gas scintillation PCs, the charge is not multiplied. However, these electrons are made to drift under an electric field and they acquire sufficient energy to excite the scintillation of the detector gas, but not to ionize it.
- Such a scintillation signal can be easily detected with a photosensor.

Gas Scintillation Proportional Counters (2)

- GSPCs reach energy resolutions of 8% at 6 keV, outperforming standard proportional counters by a factor of about 2.
- For soft X-rays below about 1 or 2 keV, GPSCs outperform any other type of large area detector, either cooled or room temperature.

X-ray CCDs (1)

- CCDs become very inefficient in the Ultraviolet, primarily because the electrodes they use become opaque to radiation.
- However, at soft X-ray wavelengths they become efficient again, especially thinned (but not so thin as optical) CCDs.
- CCDs now produce **many** electron-hole pairs per incident photon, and the number they produce depends upon the energy of the photon.
- Process is photoelectric, and is exactly the same as in a gas proportional counter.

X-ray CCDs (2)

- X-ray causes the release of a high energy electron, which in turn releases secondary electron-hole pairs by a solid state analogue of the Auger effect.
- Number of electrons released $\mathbf{N = E / w}$
- For silicon $w = 3.65$ eV
- The chip is read almost continuously, allowing reasonable time resolution (seconds).
- As some energy is transferred to the crystal lattice, there is a statistical variation in the number of electron-hole pairs produced.

- Thus, there is a solid state Fano factor, so again:

$$\sigma_N^2 = F N$$

For silicon $F \sim 0.1$

X-ray CCDs (3)

- For silicon, w is lower than for the inert gases (3.65 eV as opposed to ~20) and *F* is slightly lower (0.1 as opposed to 0.17) so CCDs have much better intrinsic energy resolution than gas proportional counters.
- Energy resolution is given by:

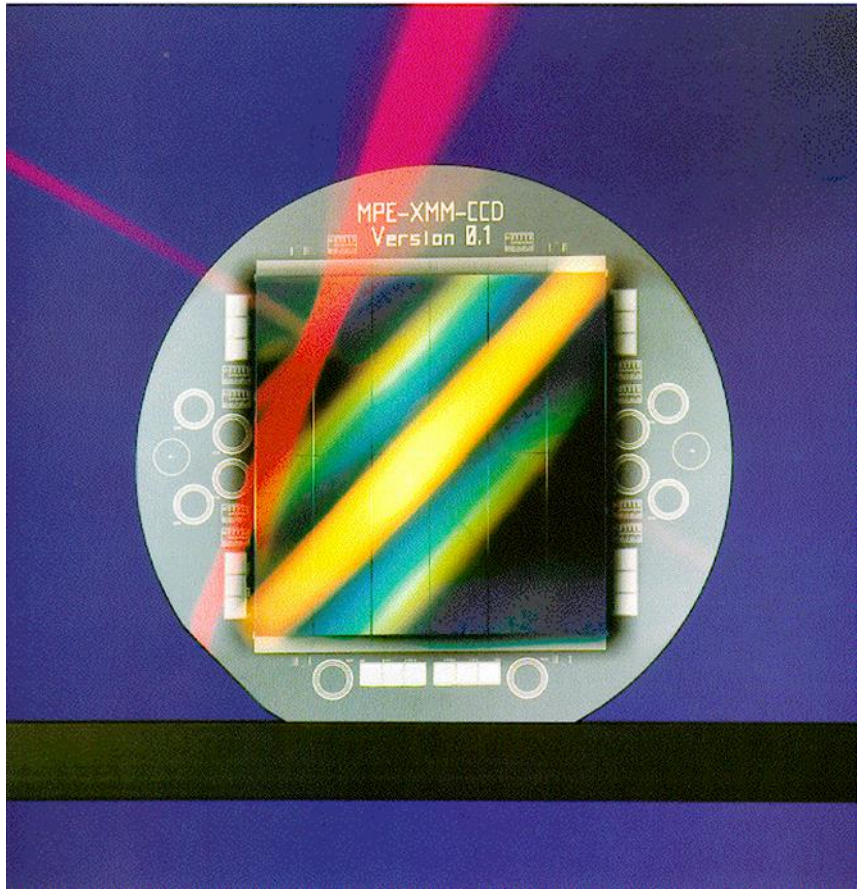
$$\frac{\sigma_E}{E} = \sqrt{\frac{wF}{E}} \quad \text{or} \quad \Delta\mathrm{E} = 2.35\sqrt{wFE}$$

factor 2.35 is because this is expressed as a full width half maximum rather than a standard deviation (FWHM $\approx 2.35\sigma$).

- This is why CCDs outperform gas proportional counters.

X-ray CCDs (4)

CCD array for the XMM-Newton satellite

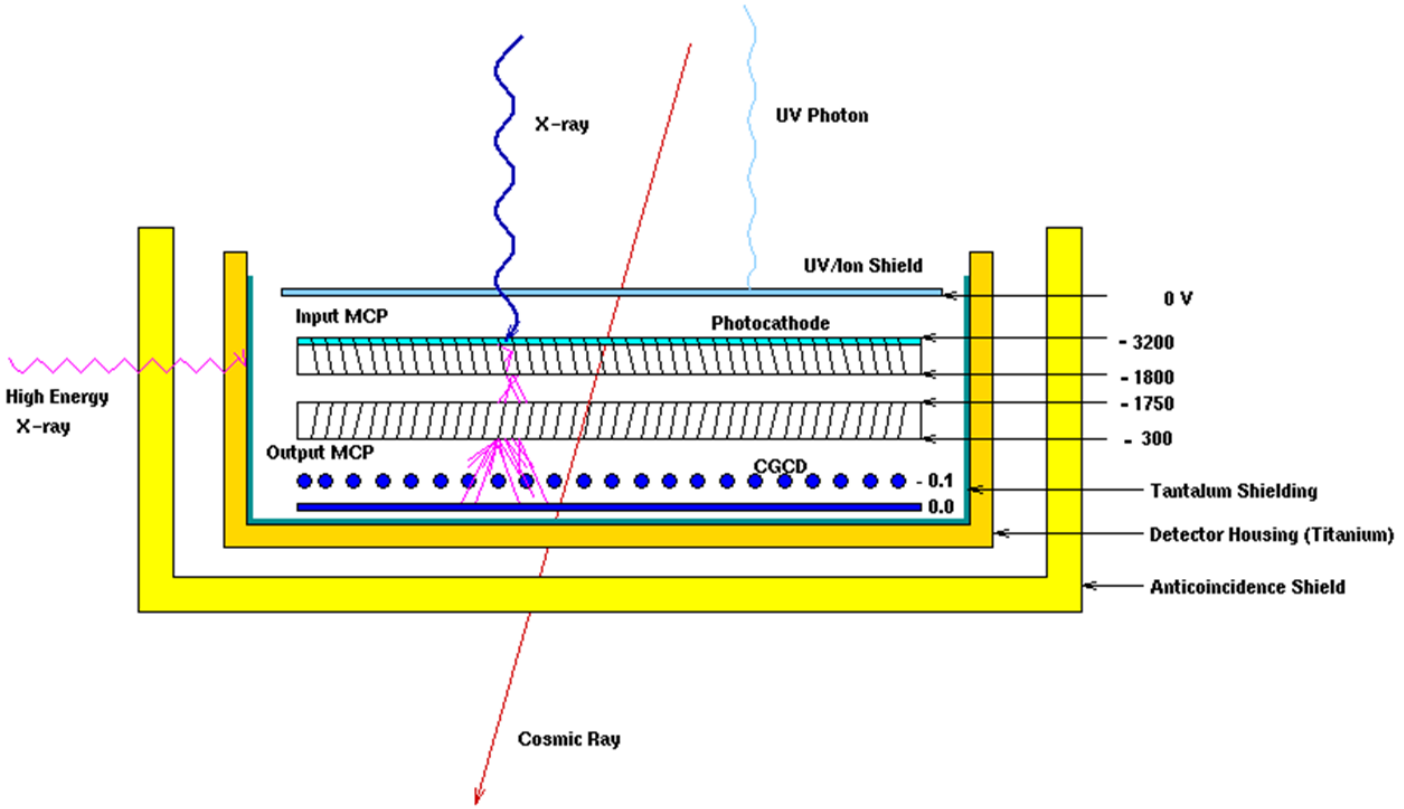


Microchannel plates (1)

- Microchannel plates work at X-ray wavelengths:
  - X-rays release photons in the lead glass channel walls directly.
  - Walls are coated with a material of high photoelectric yield, especially at energies below 1 keV.
  - At energies higher than 5 keV, X-rays can penetrate the channel walls to release photoelectrons in neighbouring channels. This degrades resolution.

Microchannel plates (2)

X-ray MCP detector flown on CHANDRA



Microchannel plates (3)

- The spatial resolution is limited to how closely the channels can be bunched together (channels can be of 12.5 microns in diameter, and are spaced at 15 micron intervals. The final resolution is limited by the channel spacing, your telescope optics, and the readout device. For Chandra , the resolution is about 0.5".
- Because these devices are operated in, or near saturation, they have insignificant energy resolution.
- The timing resolution of the devices is excellent: about 50 picoseconds (the travel time).

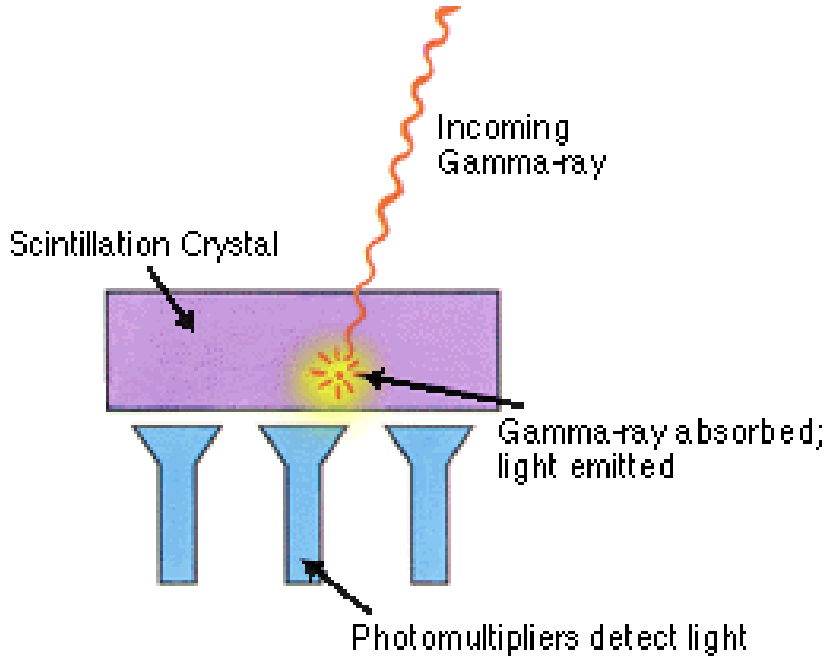
Gamma-ray Detectors

Gamma Ray Detectors

- We can't make a reflecting surface at gamma ray wavelengths, thus the distinction between telescope and detector is not so clear:
  - Scintillation crystals
  - Solid state Cadmium Zinc Telluride detectors
  - Compton Scattering detectors
  - Pair production detectors
  - Air Čerenkov detectors

Scintillation crystals

- Crystal converts a gamma ray to a shower of lower energy photons, which are detected by a photomultiplier.



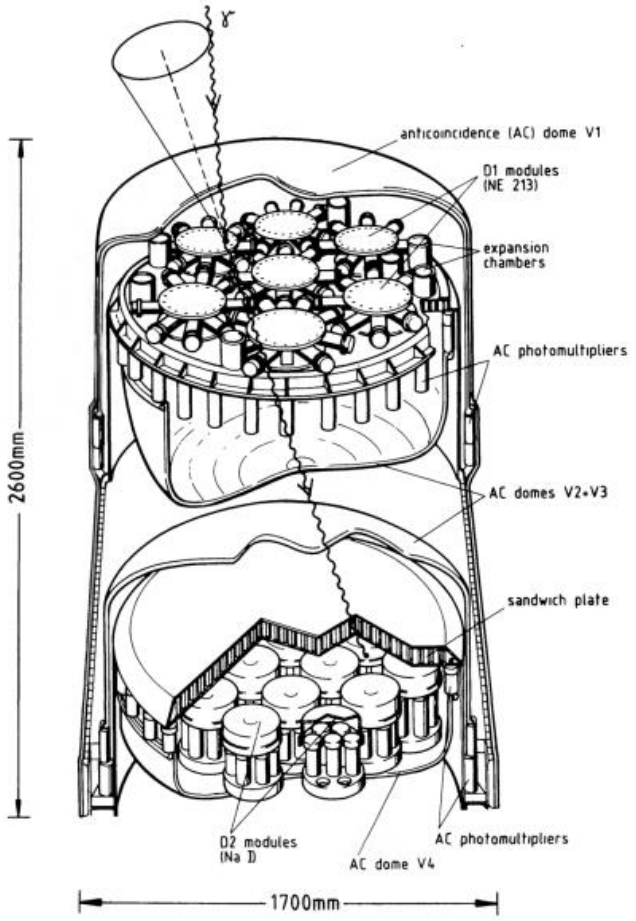
Scintillator produces very poor directional and energy resolution. **Not in use now.**

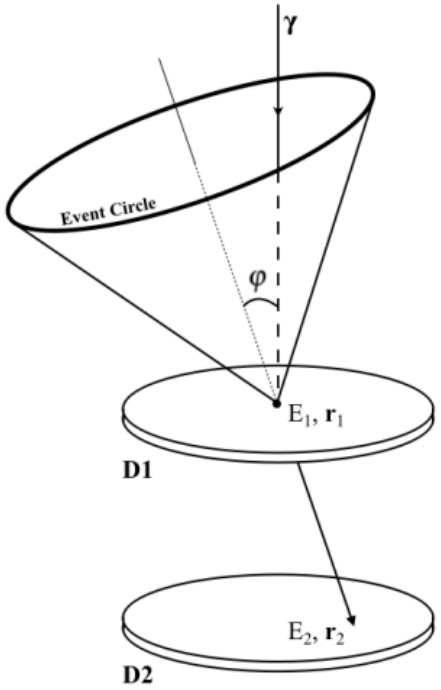
Compton Scattering Detectors (1)

We discussed them in Lecture 4 (Slides 167-168):

- Photon scatters off an electron, transferring energy to the electron.
- Detector consists of two levels. In the upper level the photon is scattered.
- Photon continues to the lower level, but electron emits a shower of photons which are measured by photomultipliers.
- Lower level is a scintillation detector, photon is absorbed again emitting a shower of low energy photons, which are measured by photomultipliers.
- From the location and energy of the two showers of photons, incoming gamma ray direction and energy can be calculated.

Compton Scattering Detectors (2)

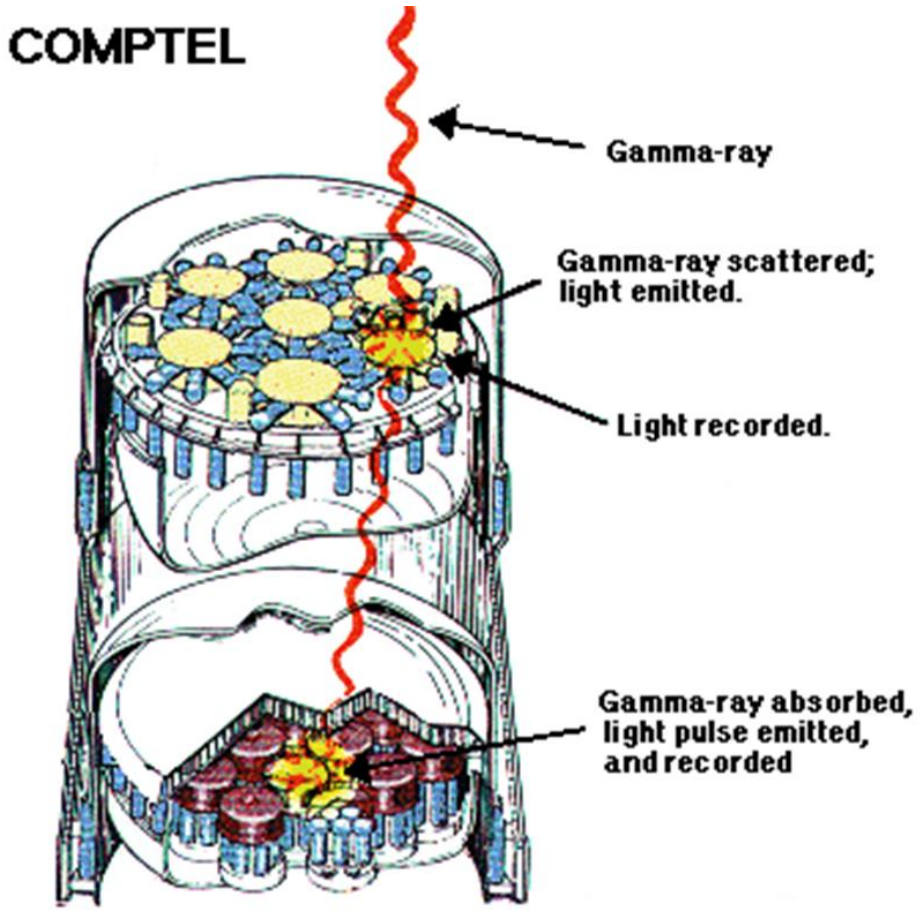




- A classic Compton telescope uses a scattering plane (D1) and an absorption plane (D2) to measure the position and energy of two interactions in a Compton event.
- From these measurements, the original photon direction can be constrained to a circle on the sky called the "event circle."

Compton Scattering Detectors (3)

The COMPTEL detector on the CGRO



Pair production detectors

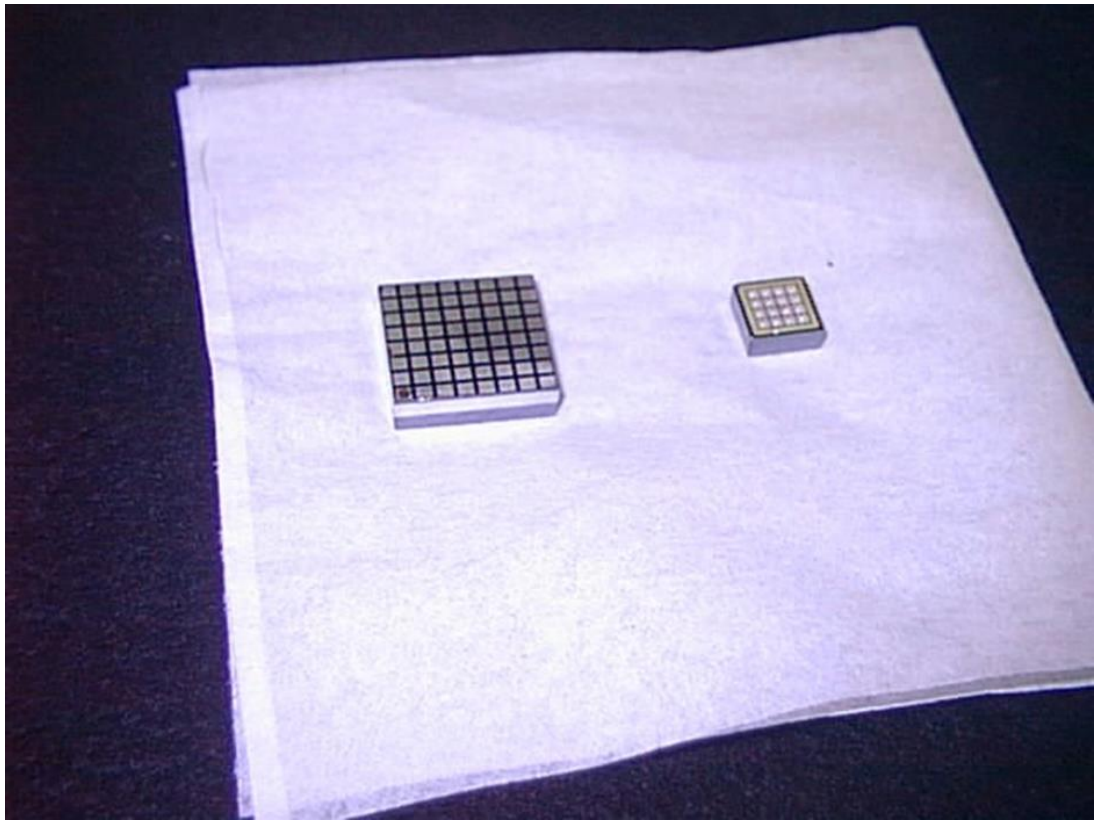
- Gamma ray with energy > 30 MeV interacts with a material to produce an electron-positron pair.
- Electron and positron are tracked by particle physics techniques (spark chamber layers).
- Energy and direction of the electron and positron give the energy and direction of the incoming Gamma ray by conservation of energy and momentum.

- A pair production detector on the EGRET satellite:



Cadmium Zinc Telluride (CZT) detectors

- Cadmium Zinc Telluride is a high band gap semiconductor.
- By adding pixel electrodes on one face of a wafer of CZT, a two-dimensional photoelectric solid-state detector can be made.
- These have small number of pixels, so pixels tend to be read out individually (rather than charge shuffled as in a CCD).

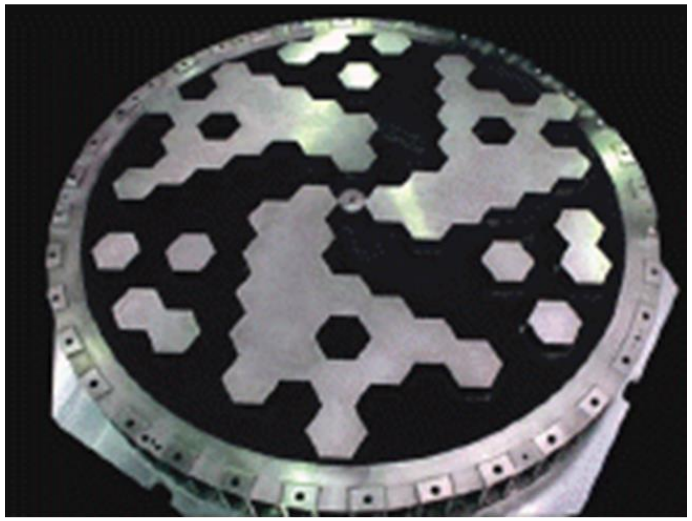


CZT crystals have a sensitivity range of 30 keV to a few MeV with a ~2% FWHM energy resolution at 662 keV.

Coded mask imaging with the CZT array

- CZT arrays are often used with coded mask imaging systems (see also Lecture 4).
- A coded mask is an array of transparent and opaque (usually lead) elements in an optimised pattern.
- It operates as an array of pinhole cameras.
- From the pattern on the CZT array you can determine the direction of the gamma ray source (there will only be one in your field of view!)

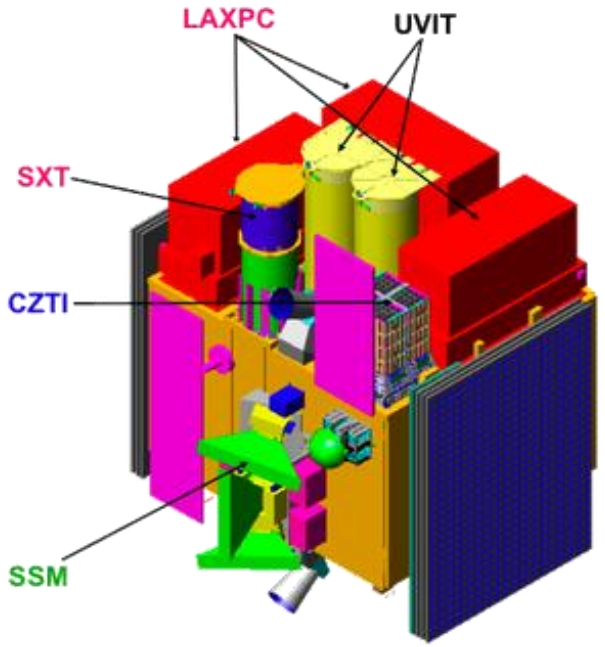


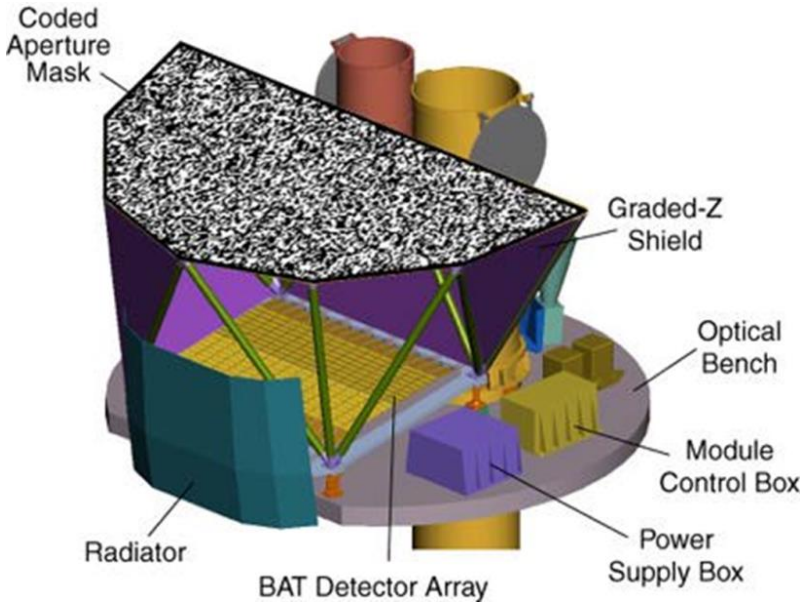


Coded mask/CZT array

Coded mask array and CZT array detector for the **SWIFT** project:
Effective area 5240 cm^2
Energy Range 15-150 keV (energy resolution ~7 keV)

AstroSat: The Cadmium Zinc Telluride Imager (CZTI) of 500 cm^2 effective area and the energy range from 10 to 150 kev.



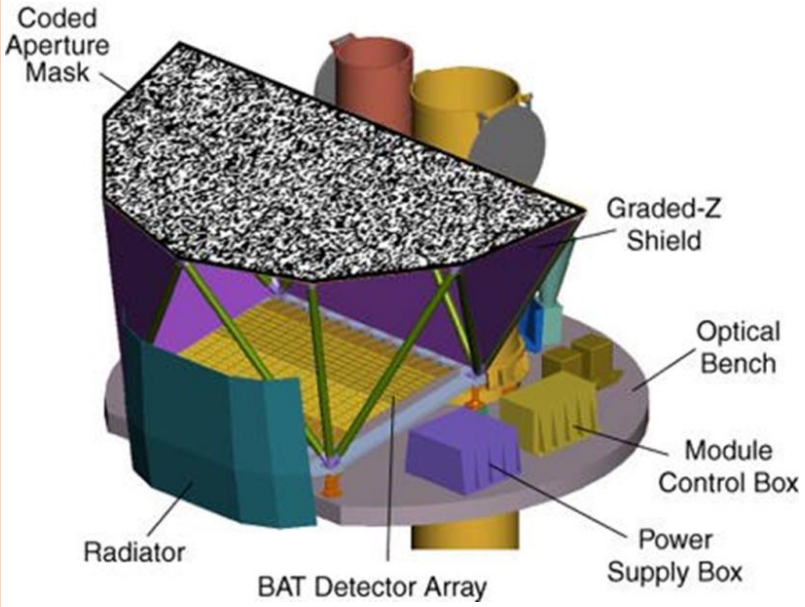


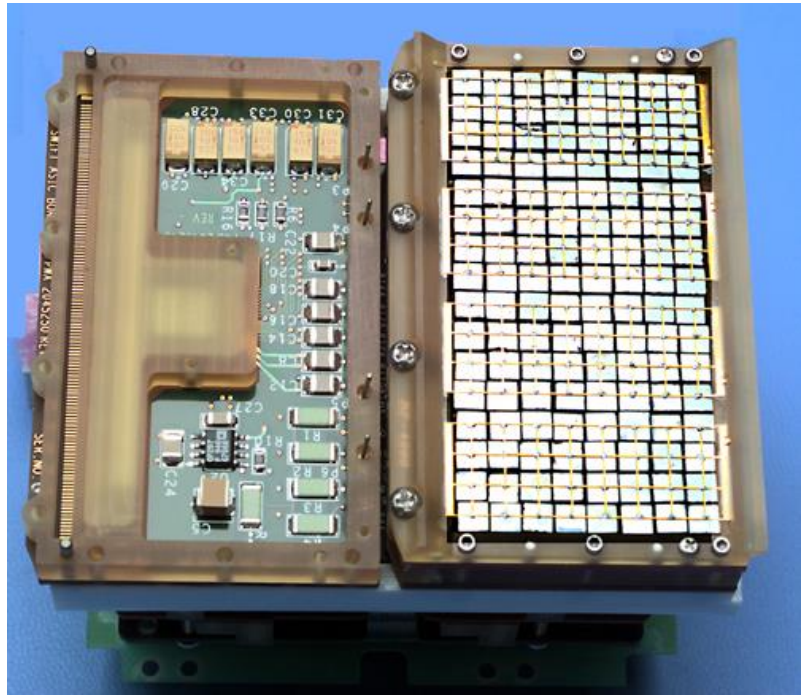
Swift's Burst Alert Telescope (BAT)

Coded mask array and CZT array detector for the **SWIFT** project: Effective area 5240 cm^2

The BAT's 32,768 pieces of 4 x 4 x 2 mm CdZnTe (CZT) form a 1.2 x 0.6 m sensitive area in the detector plane.

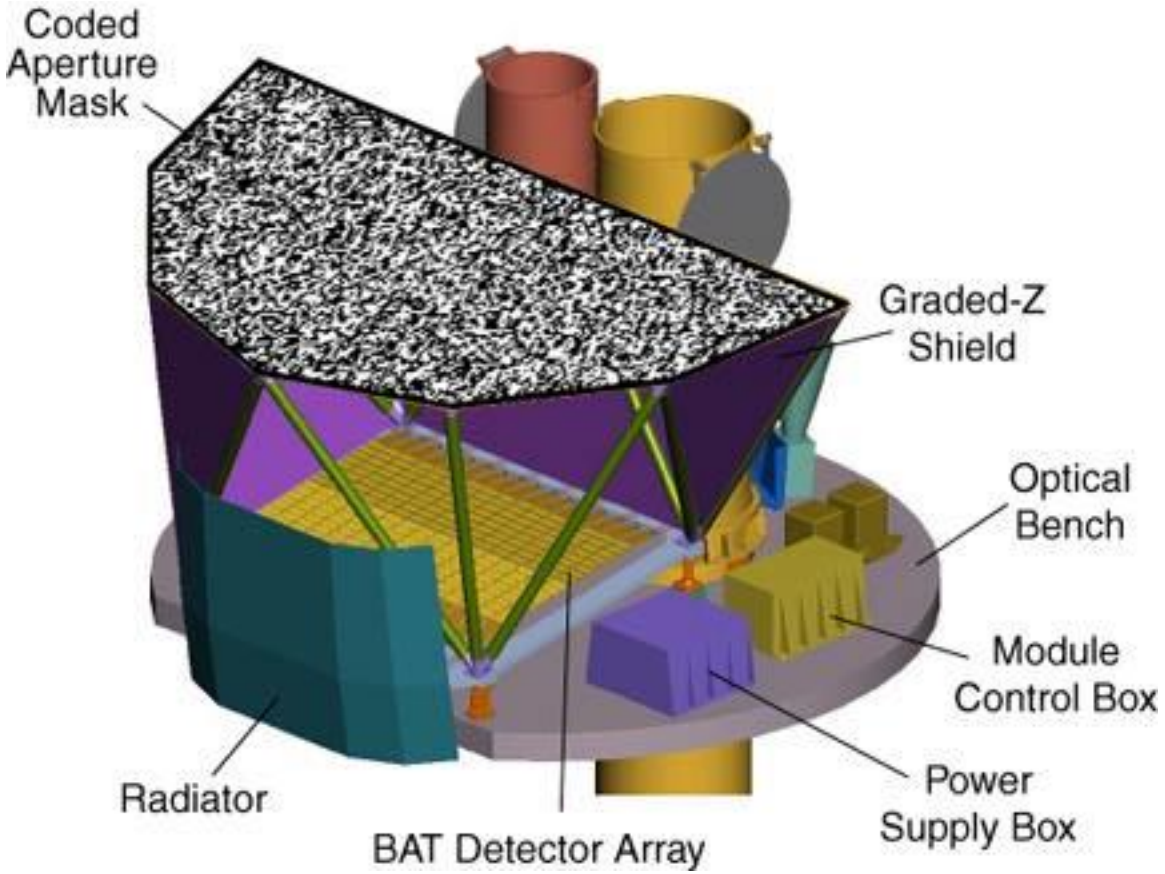
Groups of 128 detector elements are assembled into 8 x 16 arrays, each connected to 128-channel readout Application Specific Integrated Circuits (ASICs).





Coded mask/CZT array

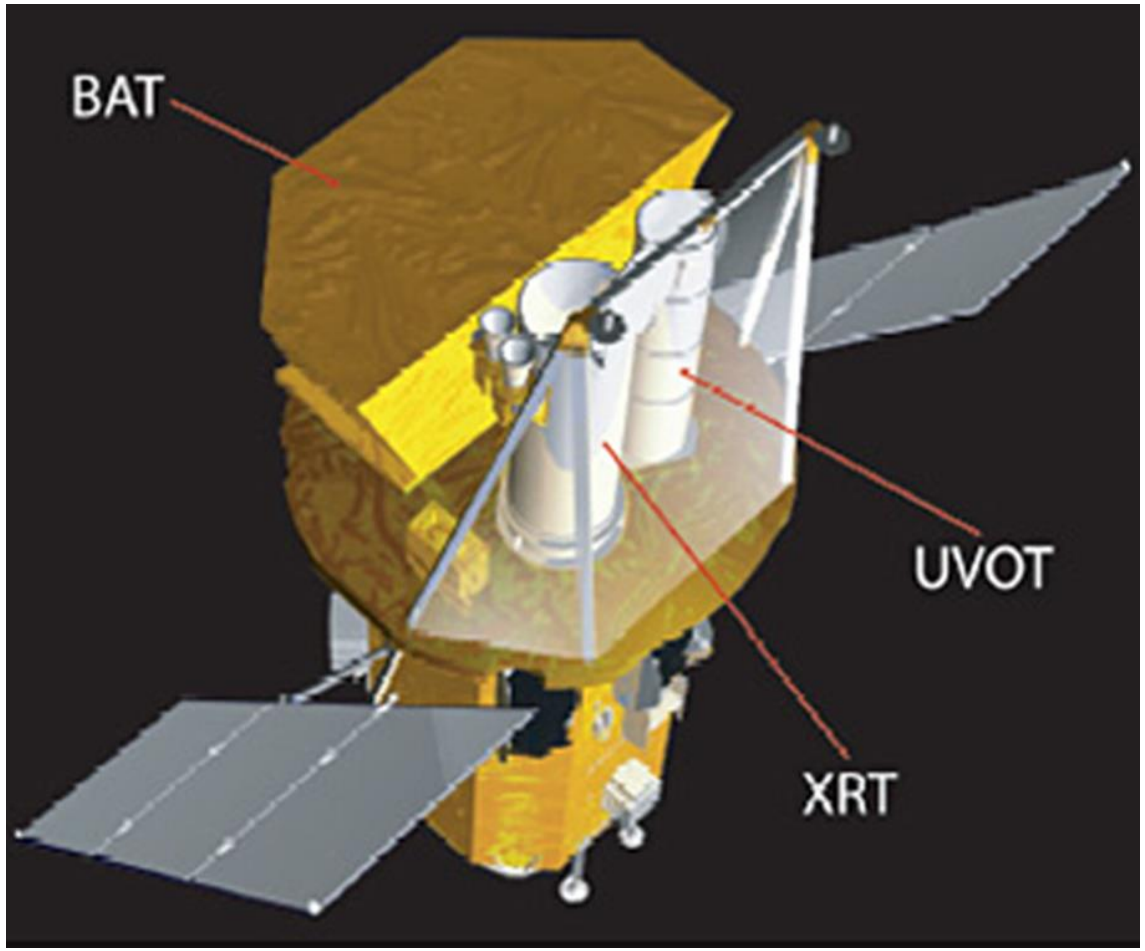
Coded mask array for SWIFT



SWIFT

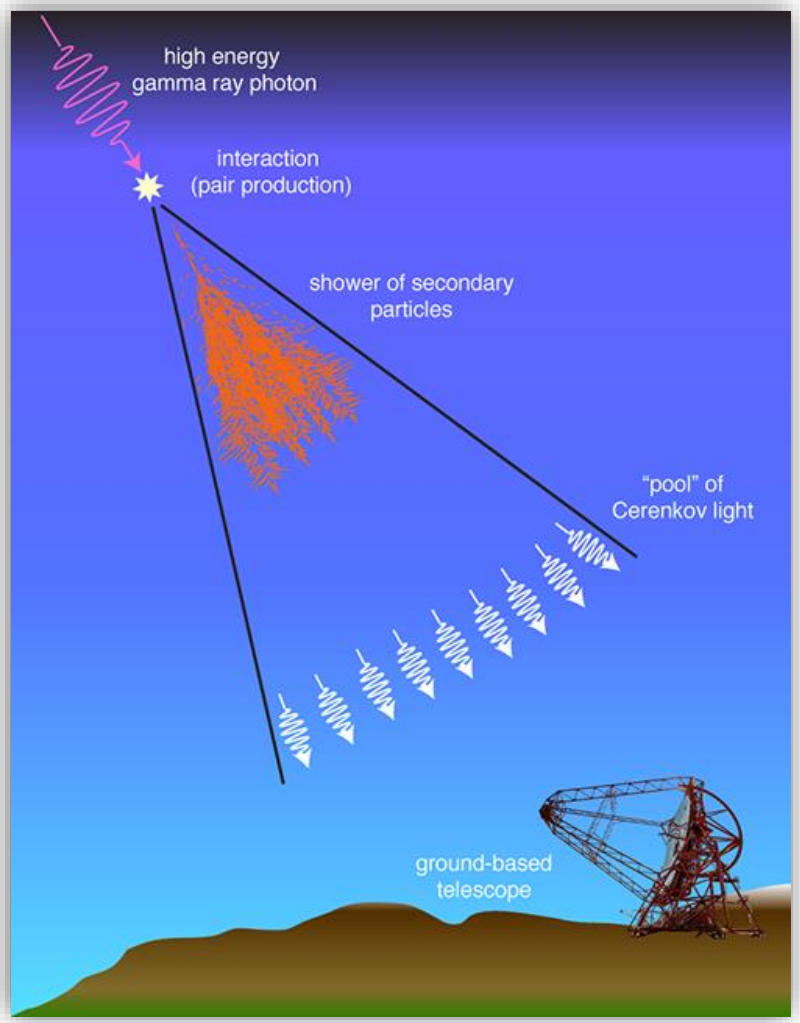
- Launched in November 2004 to detect Gamma Ray bursts.
- Separate Gamma Ray (coded mask), X-ray and UV/Optical telescopes.
- Rapid response and feedback is important.

SWIFT



Čerenkov air shower telescope

- Here the atmosphere acts as a detector for very high energy gamma rays (see Lecture 3).
- Gamma rays interact in the upper atmosphere to produce secondary particles, these particles are moving at speeds greater than the speed of light in the local medium.
- Čerenkov radiation results when charged particles move at greater than the local speed of light, and because of the rapidly changing electric field cause the local medium to radiate.
- Čerenkov radiation is concentrated in a cone.





Čerenkov telescopes are large collection area telescopes with fairly crude optics. Because the shower and the consquent Čerenkov flash is spread over a wide spatial area, arrays of such telescopes are built to improve the directionality.

The Cerenkov light reflected from the mirror is detected in the focal plane by one or many photomultipliers.



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 7

Vitaly Neustroev

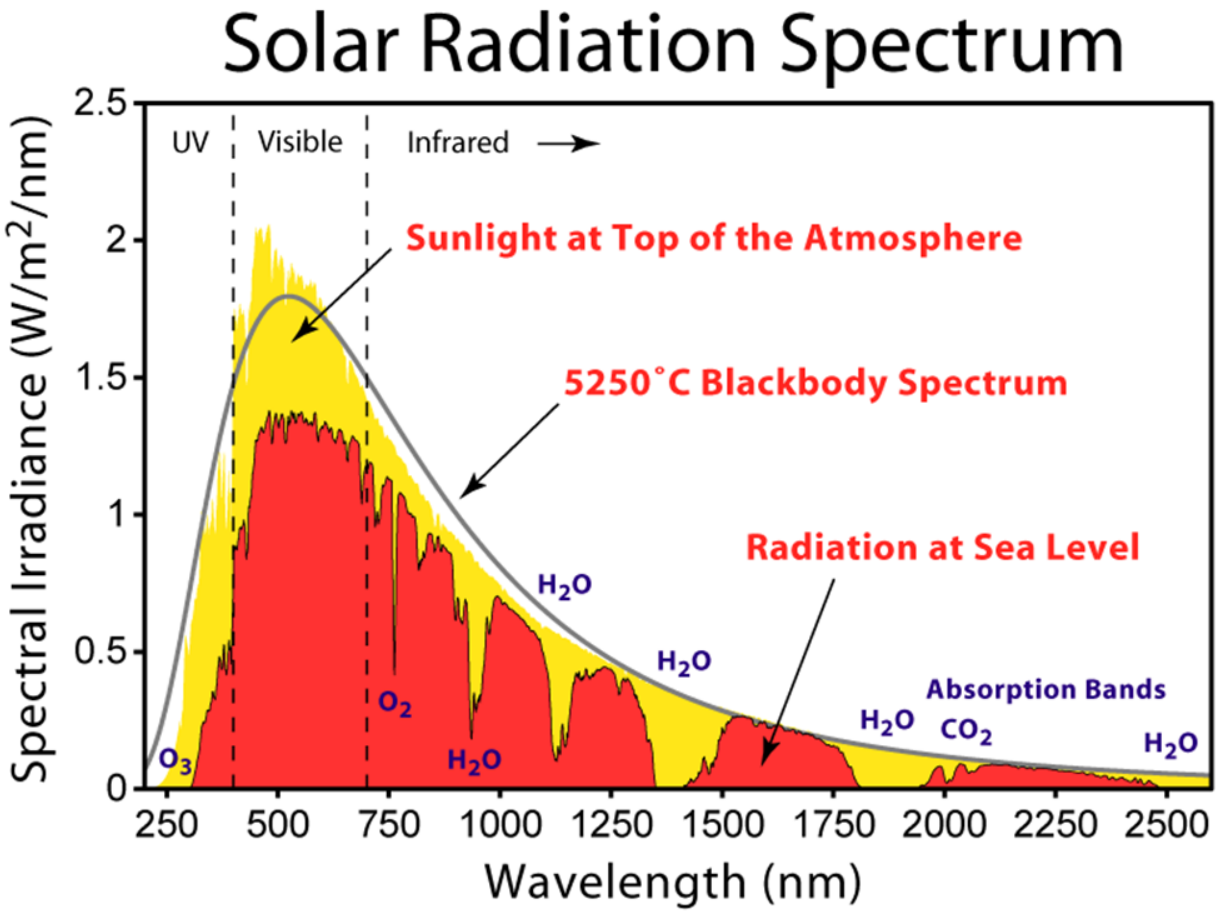
Infrared Detectors

Photoconductors

Bolometers

Infrared Passbands

Infrared passbands are determined by the windows of transparency of the atmosphere.



Infrared ranges and IR detectors

- The infrared region is divided into three wavelength ranges:
  - the near-infrared (NIR): 0.7-5 μm
  - the mid-infrared (MIR): 5-30 μm
  - the far infrared (FIR): 30-1000 μm

- Many of the detectors considered above have some infrared sensitivity, especially up to 1 μm.
- At longer wavelengths other types of detectors are needed.

Infrared Array detectors (1)

λ [Å] = 12.40/E [keV]

- Infrared semiconductor detectors cannot be made from silicon any more as the band gap is too large: the band gap of silicon is 1.1 eV and the critical wavelength is 1.1 microns. Radiation of longer wavelength cannot excite electrons from the valence band to the conduction band in silicon. Materials with smaller band gaps and longer critical wavelengths should be used.
- Germanium can be used at the shortest wavelengths, but Indium Antimonide and Mercury Cadmium Telluride are more widely used.
- People have tried making CCDs out of Indium Antimonide, but with no significant success.

Infrared Array detectors (2)

- There are two main types of IR detectors:
  - For the NIR and MIR and somewhat into the FIR — **the photoconductor.** In such detectors **individual photons create free charge carriers.** Typical IR detectors are classified as photovoltaic or photoconductive, according to whether they register photons by generating a current of electrons or merely by changing their resistance.
  - For the FIR — **the bolometers,** in which photons are absorbed and thermalized.

Photovoltaic detectors

- Single pixel infra-red detectors have long used the photovoltaic effect.
- Detectors are made of semiconductors and rely on the excitation of electrons from an energy band in which they are immobile (the valence band) to the conduction band, where they are free to move about.
- p-n junction generates an internal electric field to separate the photon generated electron-hole pairs.
- Migration of holes and electrons changes the electric field, hence there is a voltage change across the junction which can be measured.

Photoconductive Cells

- Exhibit a change in conductivity with the intensity of their illumination.
- The mechanism is the absorption of the radiation by the electrons in the valence band of a semiconductor and their consequent elevation to the conduction band.
- The conductivity increases with increasing illumination, and is monitored by a small bias current.
- There is a cut-off point determined the minimum energy required to excite a valence electron over the band gap (also correct for photovoltaic detectors).

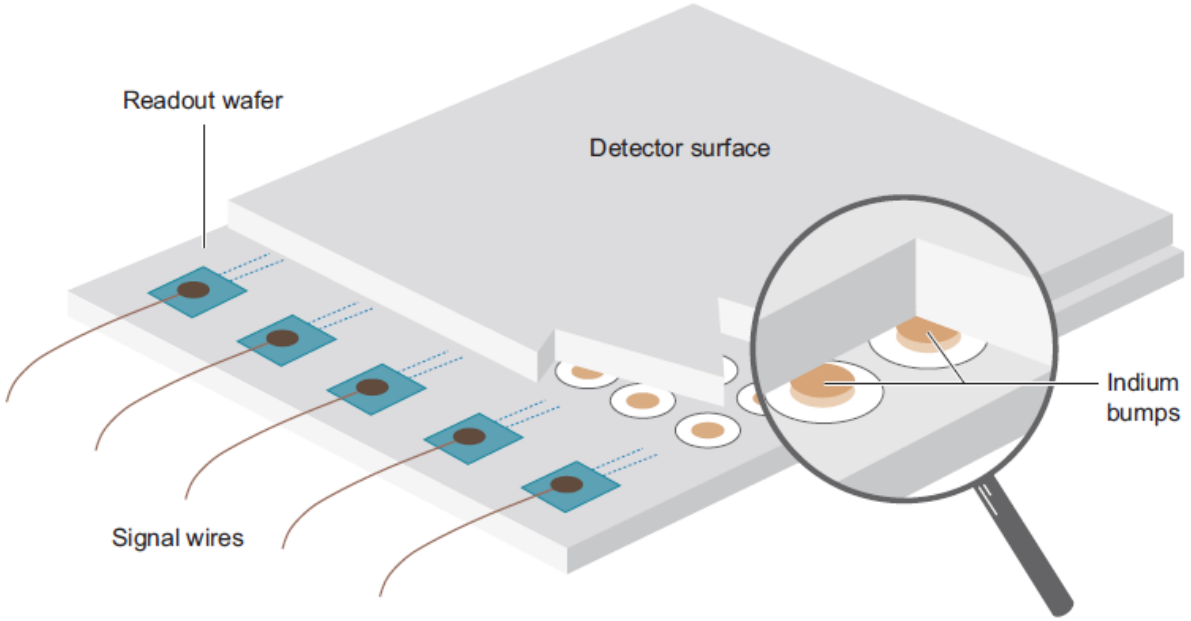
Infrared Detectors: limitations

- The smaller the bandgap, the more likely electrons are to enter the conduction band unwanted, by thermal excitation. It is therefore necessary to cool these longer-wavelength detectors more than the shorter-wavelength ones.
- The longer wavelength the colder the required temperature:
  - in the NIR – 77 K (liquid nitrogen);
  - in the MIR – 4 K (liquid helium);
  - in the FIR – down to 100 mK are used.

Hybrid Arrays (1)

- Modern IR arrays are hybrid arrays, formed of a sandwich of three layers.
- Top layer (assuming radiation is coming down) is an Indium Antimonide or Mercury Cadmiun Telluride, doped to act as a photovoltaic detector.
- Bottom layer is a silicon multiplexer (amplifier).

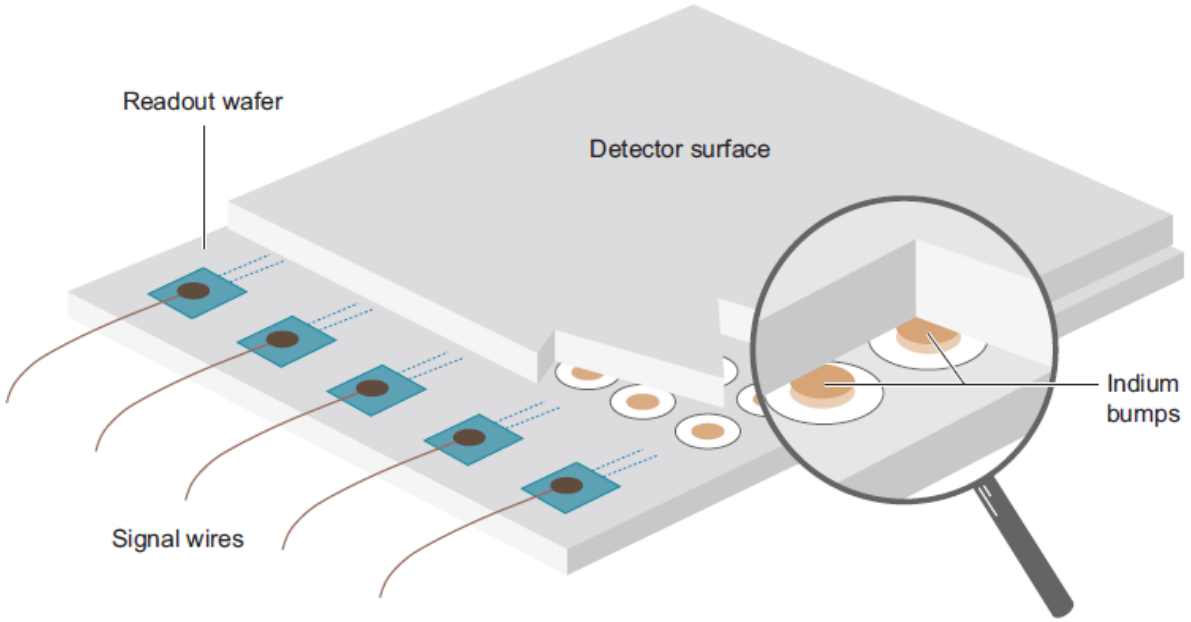
  It can be constructed using conventional Si-based techniques, but more often an array of tiny MOSFET (Metal Oxide Semiconductor Field effect Transistor) amplifiers used.



- The two parts are joined electrically with one connection for each pixel.

Hybrid Arrays (2)

- In between two layers are Indium bump bonds providing an electrical connection between locations on the IR detector and the elements of the silicon multiplexer.
  Indium is a good conductor which is soft even at low temperature, so the device does not crack when cooled.
- Gaps between the indium bonds are filled with epoxy resin to provide mechanical stability.



Infrared arrays (1)

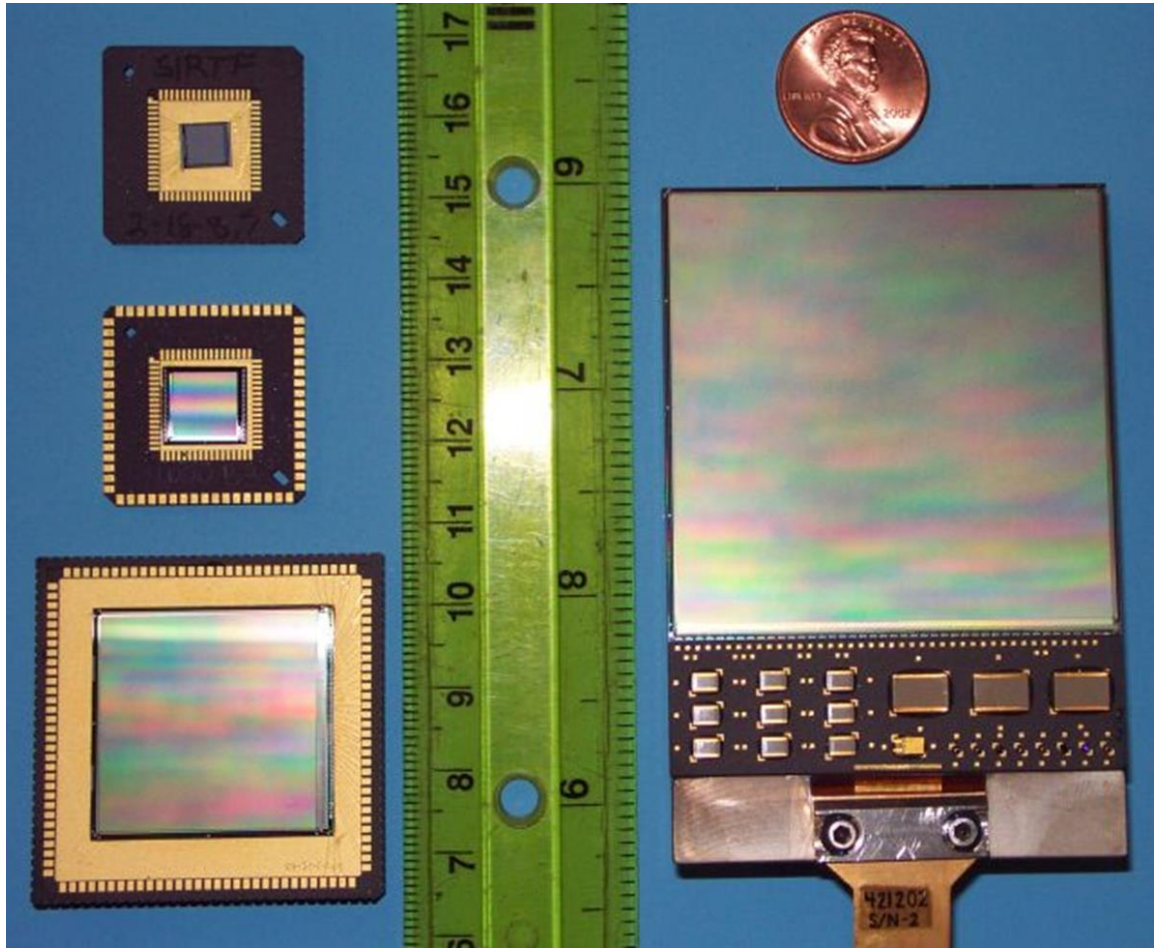
- Indium Antimonide and Mercury Cadmium Telluride arrays are available in 1024 square formats, maybe 2048 square.
- Quantum efficiency 60% - 80%
- Noise 40 electrons/pixel
- Cost ~ €300,000 per array!

Infrared arrays (2)

Original
64×64

256×256

1024×1024



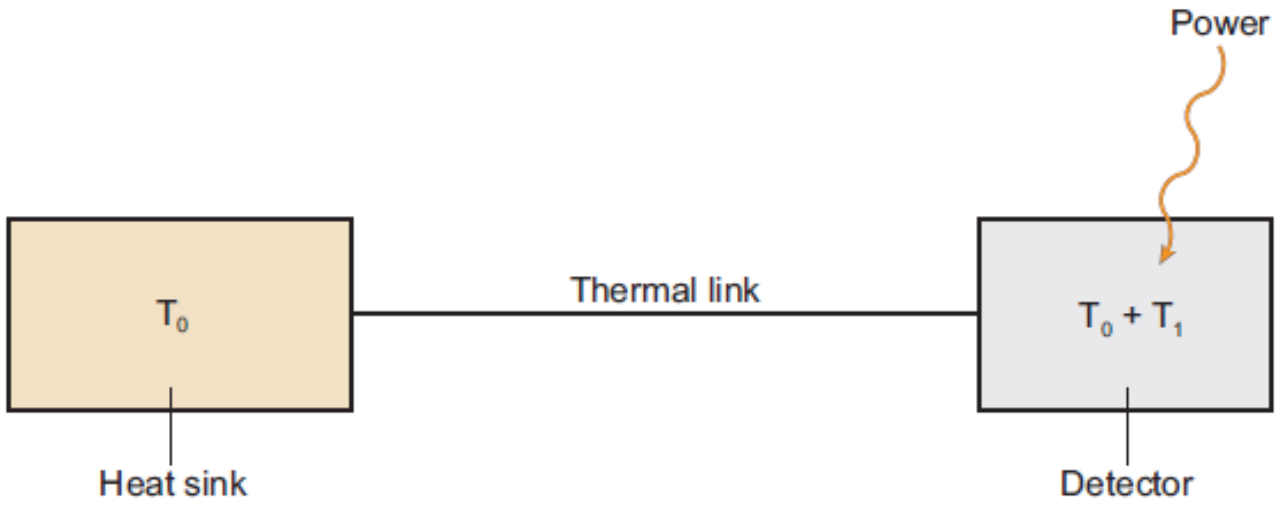
2048×2048

Bolometers (1)

- Bolometers operate on a different principle from the detectors discussed before.
- Rather than individual photons creating free charge carriers, photons are absorbed and thermalized, and the resulting energy is sensed.
- This approach yields very high performance detectors for the submillimeter and millimeter spectral regions.

Bolometers (2)

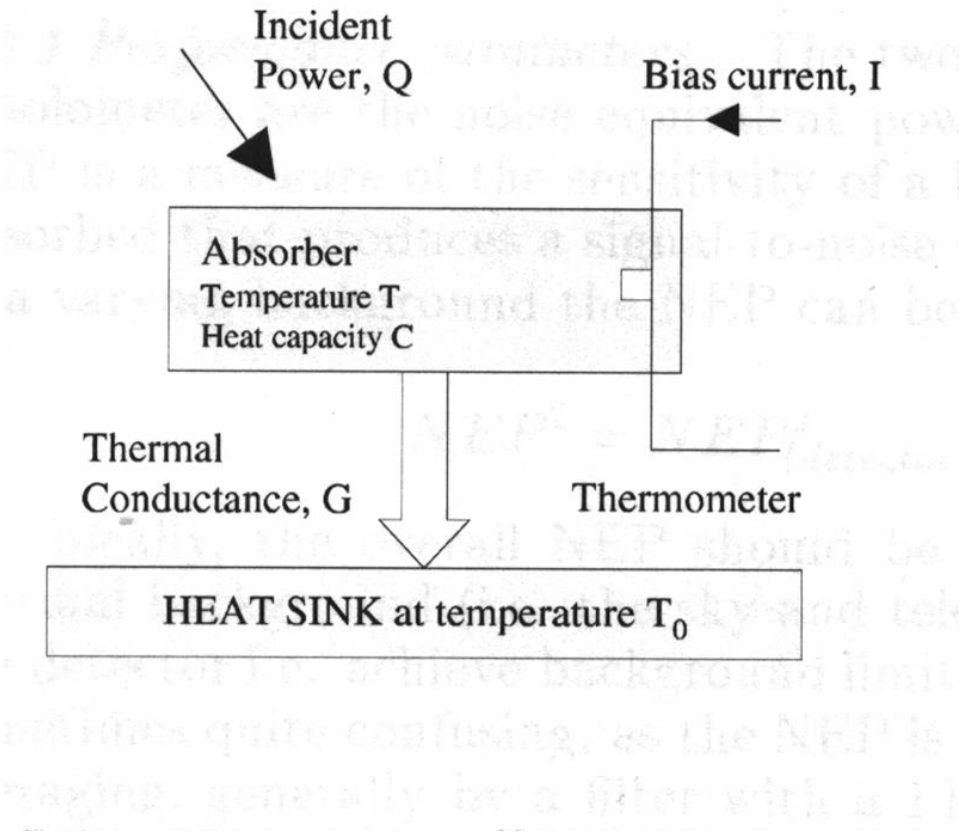
- Photons incident on the absorber raise its temperature, causing a sensitive thermometer attached to it to change resistance, producing a signal that can be amplified to achieve a detection.
- Bolometers must be operated at low temperatures to suppress thermal noise that arises as a result of thermodynamic fluctuations in the flow of energy across the thermal link.



Components of a Bolometer

Self reading material

- Absorber with heat capacity C
- Thermometer with resistance R, which is usually some device whose resistance changes as a function of temperature
- Heat sink held at fixed temperature T_0
- "Weak Link" – small thermal conductance G between absorber and heat sink.
- Load resistor R_L
- Constant current power supply. generating bias current I
- Device to measure voltage changes



Noise Equivalent Power (NEP)

Self reading material

- The Noise Equivalent Power (NEP) is a measure of the sensitivity of the bolometer, and is defined as the power absorbed which produces a signal-to-noise of unity at the output.
- Units are Watts/$Hz^{0.5}$

Components of the Bolometer NEP

- **Johnson noise** – due to the random motions of the electrons in the thermometer:

$$NEP_J^2 = 4\,k\,T\,R\,/\,S^2$$

where S is the responsively of the device, and gives the output power per unit input voltage, in Volts per Watt.
k is Boltzmann's constant

- **Phonon noise** – due to quantization of the phonons which transport energy between the absorber and the heat sink along the thermal conductance G.

$$NEP_P^2 = 4\,k\,T^2\,G$$

Components of the Bolometer NEP

- **Johnson noise:**

$$NEP_J^2 = 4\,k\,T\,R\,/\,S^2$$

- **Phonon noise:**

$$NEP_P^2 = 4\,k\,T^2\,G$$

So to minimize the detector noise we need to minimize **R, T, G** and maximize **S.**

Components of the Bolometer NEP

Self reading material

- However, the time constant of the system is:

$$\tau = C / G$$

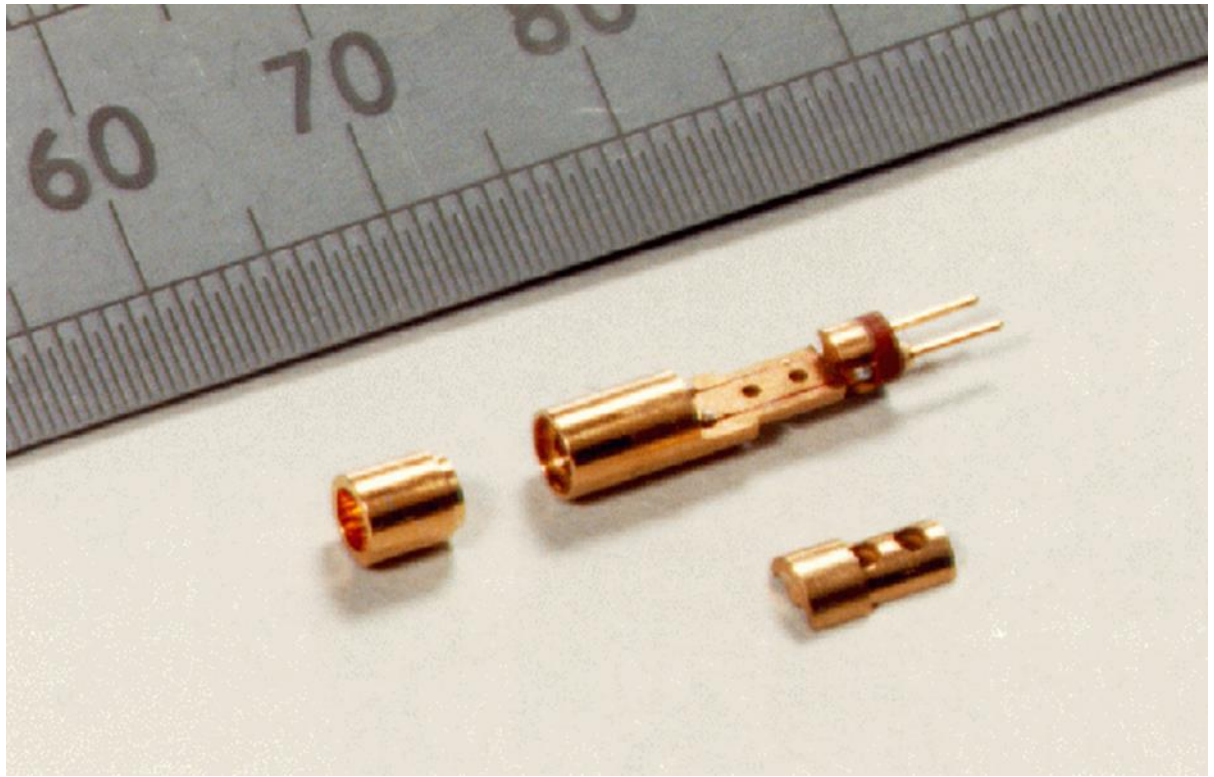
- If the time constant is long then time variability of the background dominates over the noise sources, so the choice of G is a tradeoff between lowering the NEP and the response time.
- Temperature coefficient of resistance is given by:

$$\alpha = (T/R)(dR/dT)$$

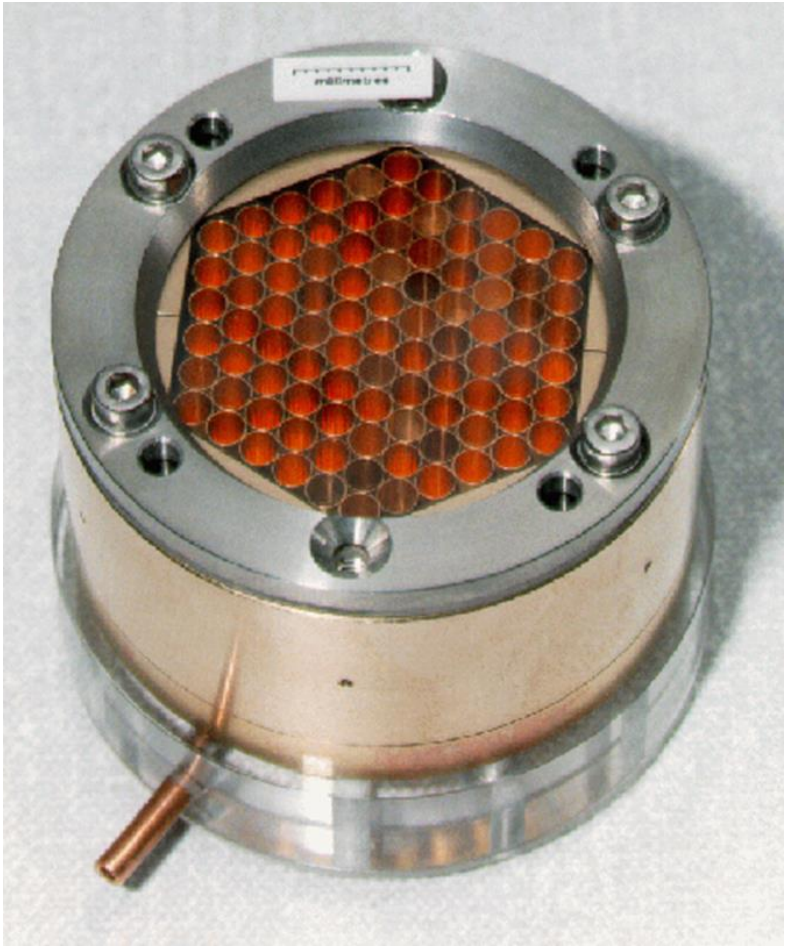
The responsively of the device depends upon α. For a typical semiconductor bolometer $\alpha \approx 5 - 10$.

Semiconductor bolometers

- Most bolometers incorporate a semiconductor resistance thermometer, and a metal coated dielectric as the absorber.



Bolometer Arrays





OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 8

Vitaly Neustroev

Methods of Observations

- Imaging
- Photometry
- Spectroscopy
- Polarimetry
- Interferometry

The Primary Tools of Astronomy

Light is the only thing we can work with. What can we do with it?

- Imaging – we can take "pictures" of the things we see. But pictures alone tend to lack the "quantitative" aspect that is needed for most serious scientific studies.
- Photometry – the technique that measures the relative *amounts* of light in different wavelength ranges.
  But these ranges are too wide to provide detailed information on the light's spectral distribution.
- Spectroscopy (spectrophotometry) – the most informative technique of light analysis, that measures how much light an object produces at various wavelengths of light.

Imaging

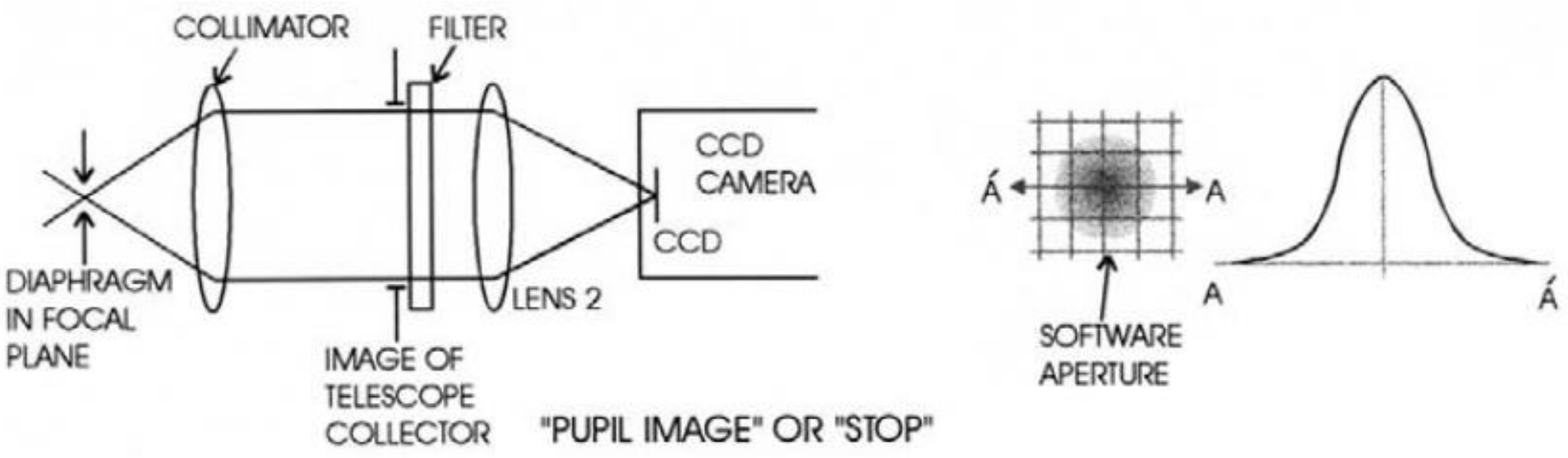
- Mapping the distribution of celestial sources on the sky in order to locate the position of source precisely – astrometry.
- Getting information on the source's form and that of its local environment.

Imaging: Pixel sampling

Pixel sampling and matching to the plate scale.

Two issues:

1. Maximizing observing efficiency → more light onto a pixel → less integration time
2. No compromise to the ability to obtain accurate brightness measurements



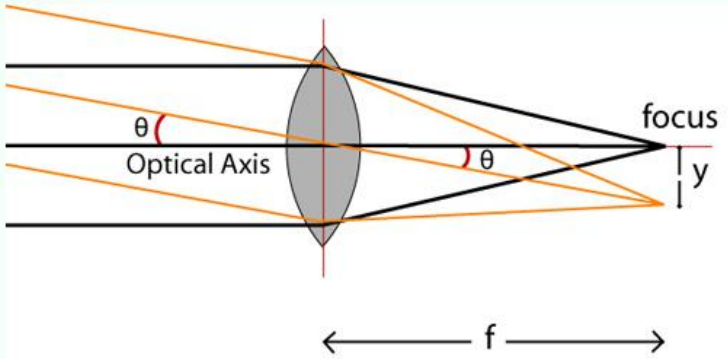
Imaging: specifications of a telescope

Two concepts used by astronomers when describing the specifications of a telescope are the plate scale and the focal ratio:

- **Plate scale**, *ps*, relates the angular size of an object on the sky, θ, to the size of its image in the focal plane, y. The term comes from the time when photographic plates were used to record images.

$$\tan\theta \approx \theta = \frac{y}{f} \qquad \rightarrow \qquad ps = \frac{\theta}{y} = \frac{1}{f}$$

where θ is measured in radians



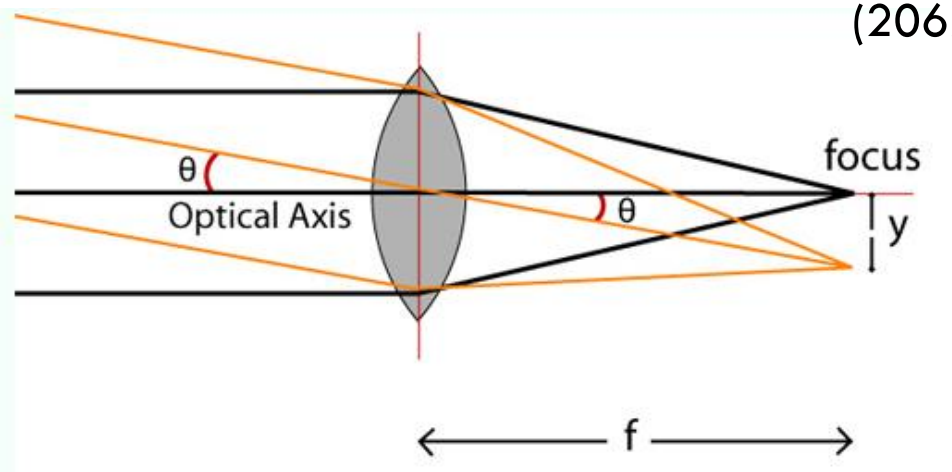
- If an astronomer wishes to study an object of angular size in **high** spatial detail, *ps* needs to be small, and therefore a **long** focal-length telescope is required.

Imaging: Plate scale

- Astronomers usually refer to the plate scale in units of arcseconds per mm. Then, the plate scale of the telescope:

$$ps = \frac{206265}{f} [“/\text{mm}]$$

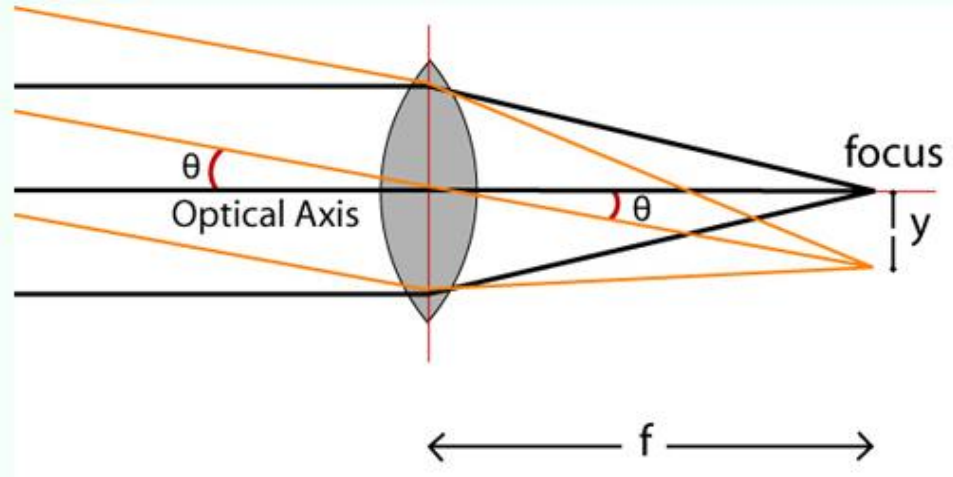
(206265 is ~ the number of arcseconds in 1 radian)



- For direct imaging, the angle on the sky subtended by the detector pixel is $\theta = ps\ d_{pix}$, where d_{pix} is the physical pixel size in mm.

Imaging: Focal ratio

- **Focal ratio** is defined as the ratio of the focal length of the telescope to its diameter, i.e. f/D.



- The term is often used in photography, where it refers to the “speed” of the camera. The larger the focal ratio, the "slower" the camera, as the amount of light falling on a given area of the focal plane is smaller. For larger f at fixed D, the plate scale ps is smaller and hence the light is spread out over a larger area. For a "faster" camera, the reverse is true.

Imaging

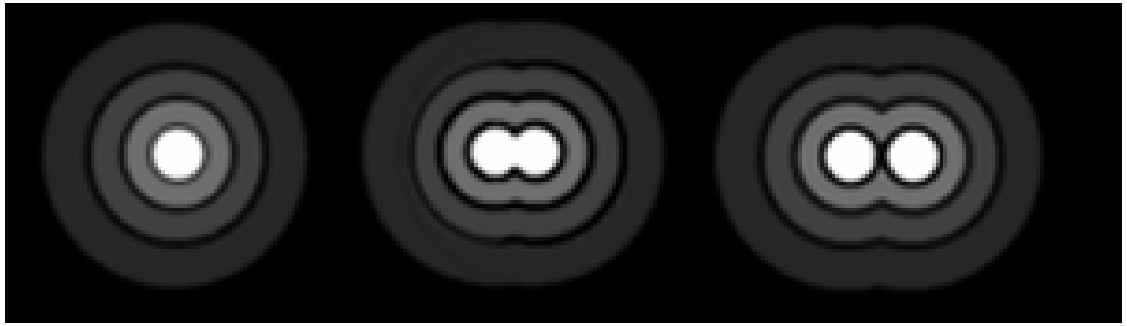
The spatial resolution element may be determined by seeing conditions or by optical constrains.

- In general, the image is critically sampled if there are about 2 pixels (the Nyquist limit) across the resolution element
- Or the image is oversampled if there are a few pixels (~4-5) across the resolution element

The resolution element: diffraction limit and seeing!

Effect of diffraction on image resolution

- Two distant point sources through a circular aperture:

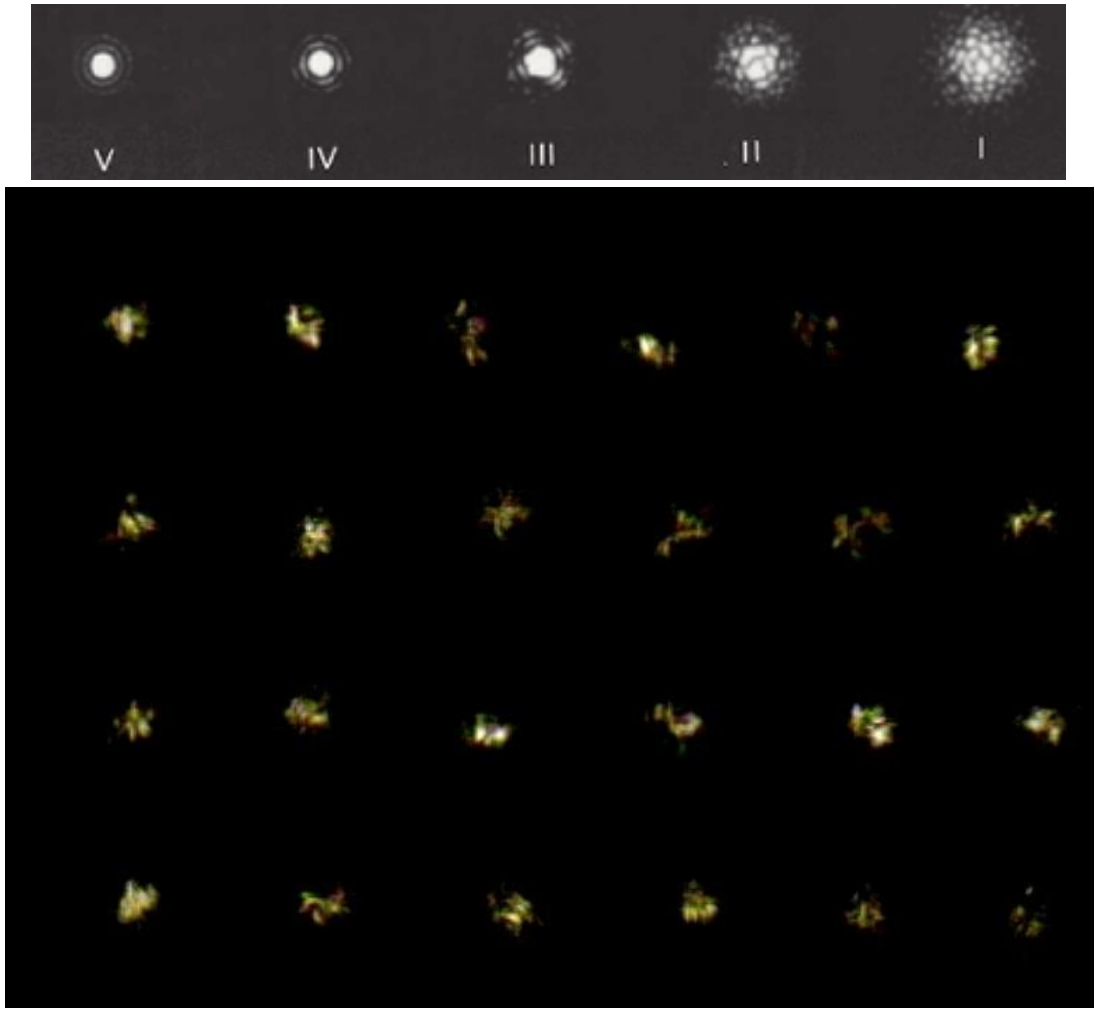


- **The Rayleigh criterion:** two images are just resolved if the centre of the first Airy pattern is superimposed on the 1st dark ring of the 2nd pattern.
- According to Rayleigh criterion, the minimum resolvable angular separation or angular limit of resolution is given by:

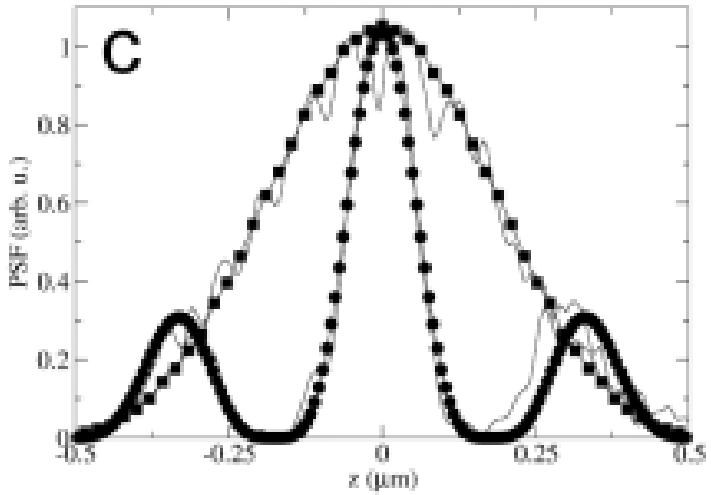
$$\vartheta \cong 1.22\ \lambda / D \text{ (radians)}$$

λ is wavelength and D is the diameter of the aperture

Imaging and Seeing



Ground-based telescope resolution is determined by the atmosphere, rather than aperture size. The central maxima of a Point Spread Function (PSF) expanded by turbulence is called *seeing disc*, and its FWHM is *seeing*.

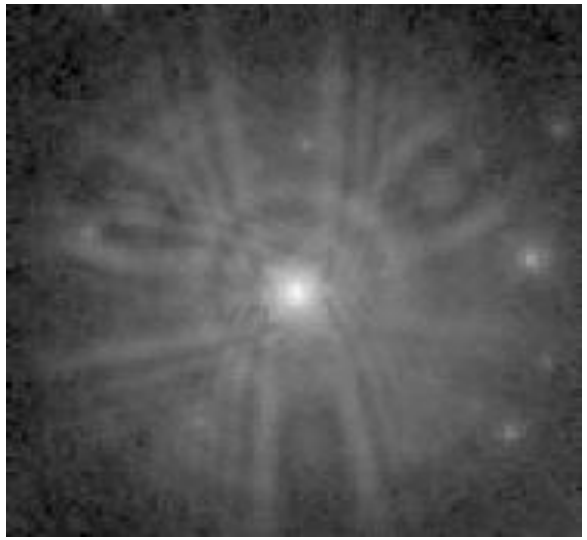


PSF, convolution & deconvolution (1)

- The **Point Spread Function** (PSF) is the main brick that builds up the whole acquired image.

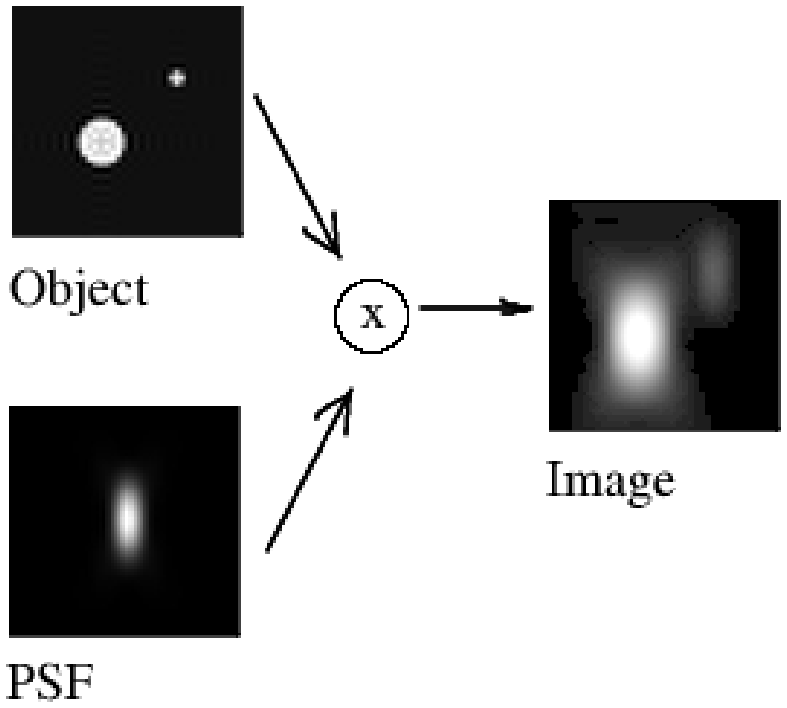
- The PSF is the image of a single point object (rescaled to make its integral all over the space equal 1). The degree of spreading (blurring) in the image of this point object is a measure for the quality of an image.

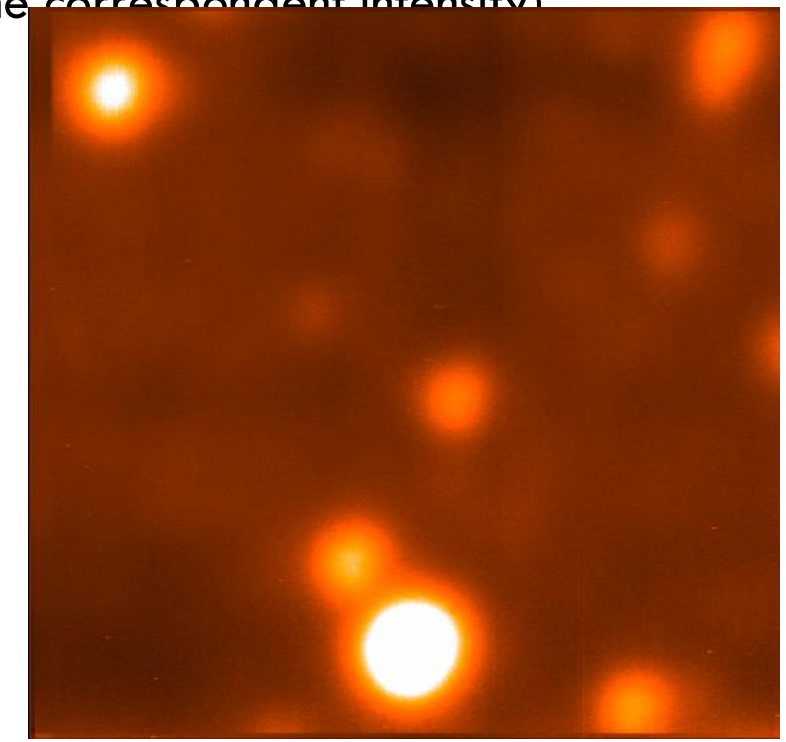
- The point spread function of Hubble Space Telescope's WFPC camera before corrections were applied to its optical system.



PSF, convolution & deconvolution (2)

- **Convolution:** The process is mathematically described by a convolution equation of the form $g = h * f$, where the image *g* arises from the convolution of the real light sources *f* (the object) and the PSF *h*.
- You can imagine that an image is formed in your telescope by *replacing* every original Sub Resolution light sc[illegible]ied by the correspondent intensity)





PSF, convolution & deconvolution (3)

- **Deconvolution (Image restoration)** is the recovery of images from raw data. Deconvolution is an "inverse" problem: the object of deconvolution is to find the solution of a convolution equation of the form

$$h * f + \varepsilon = g,$$

where h is PSF, g is some recorded image, ε is noise that has entered our recorded signal, and f is the real light source that should be recovered.

- The algorithms and methods are:
  - **Richardson – Lucy** deconvolution
  - **Maximum Entropy Method**





Lucky Imaging

- Lucky Imaging is an effective technique for delivering near-diffraction-limited imaging on ground-based telescopes.
- The basic principle is that the atmospheric turbulence that normally limits the resolution of ground-based observations is a statistical process. If images are taken fast enough to freeze the motion caused by the turbulence we find that a significant number of frames are very sharp indeed where the statistical fluctuations are minimal.
- By combining these sharp images we can produce a much better one than is normally possible from the ground.

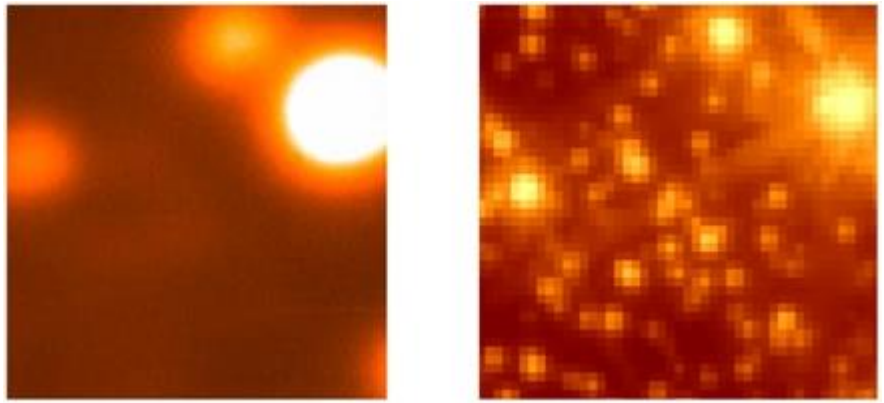


Image processing (1)

- A CCD image is built up from three signal sources:
  - object: photons imaged onto each pixel of the CCD
  - dark current: thermal electrons collected in each pixel during the exposure
  - bias: a low level electrical signal added to each pixel during image readout

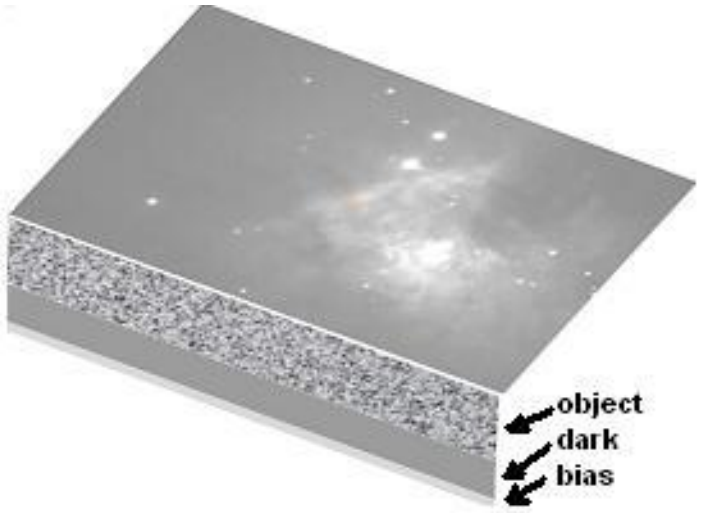
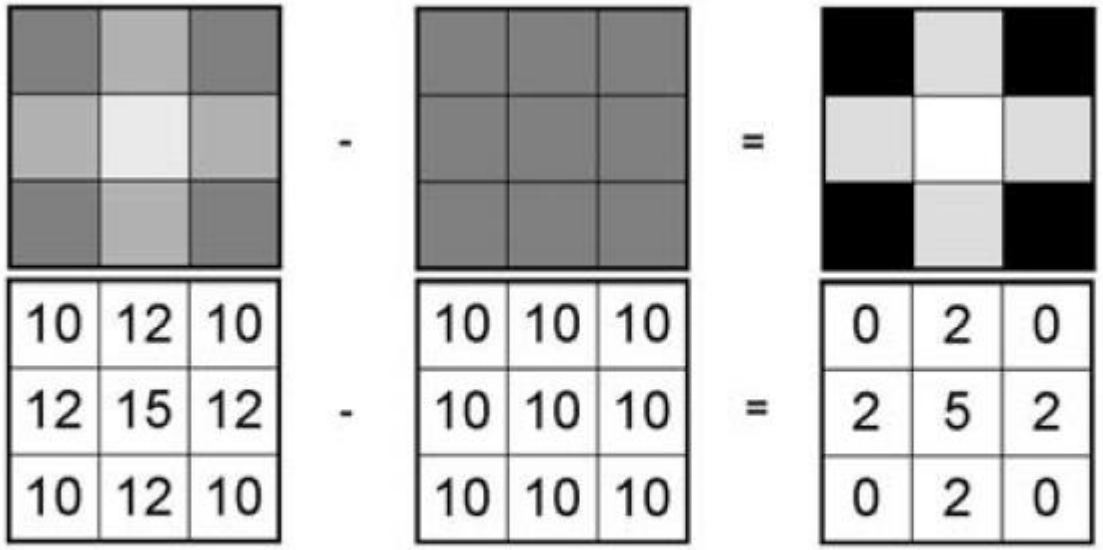


Image processing (2)

- Most processing steps must be carried out pixel by pixel



A simple operation on digital image data to subtract one frame from another.

Image processing (3)

- **Darks:**
  - In order to remove the accumulated background due to thermal dark current, it is usual to take several "darks" which are images taken with the CCD camera shutter closed. These darks should be of the same exposure time and camera temperature as the object exposure to be dark subtracted.
  - It is best to take a set of dark images and then to combine them to get a "master" dark based on the average of the dark set.
    [Note: it is usual to "median combine"]

Image processing (4)

- **Bias:**
  - In addition to thermal noise, each pixel charge will carry with it a fixed offset voltage value called the bias. Thus even if the output coming from the CCD were exactly zero electrons for every pixel, there would still be a signal that would vary from pixel to pixel in a repeatable fashion. An bias frame is one taken to determine this bias pattern.
  - Again, it is best to take a set of bias images and then to combine them to get a "master" bias based on the average of the bias set.

Image processing (5)

- **Flat Fielding:**
  each pixel on the CCD may have a different sensitivity to incoming photons due to small variations in individual pixel dimensions and quantum efficiency. For precision photometry it is necessary to calibrate such pixel-to-pixel variations and this is the function of "flat fielding". There are several different approaches but the two that are applicable to most observatories are
  - Dome Flats: A uniformly illuminated target is installed in the dome or attached to the front of the telescope
  - Sky Flats: Images are taken of the sky which is assumed to be uniform in brightness over the (usually small) field of view of the CCD

Image processing (6)

- **Flat Fielding:**
  - Take an exposure of a source which will uniformly illuminate each pixel of the CCD;
  - Pixel sensitivity (and perhaps other effects) will be a function of wavelength so separate flat fields are needed for each filter;
  - Use a long enough exposure time to fill pixels to more than 50% of their full well capacity.

Image processing (7)

- **Flat Field** types:
  - Twilight Flat: Exposures of the twilight sky (well away from horizon). Can correct for all of the types of sensitivity variations but the twilight sky is typically much bluer than the typical program object.
  - Dome flat: Exposure of a special target, usually mounted on the dome (such a target is completely out of focus and thus effectively uniform). Does well on Pixel to Pixel variations, poorly on vignetting.
  - Sky Flat: Median combination (to remove stars) of many exposures of the night sky
    Good correction of vignetting, poor correction of Pixel to Pixel variation.
    Hard to get enough photons on CCD – exposures may be very long.

Image processing (8)

- **Flats:**
  - Take a number of dome or long exposure (5 to 10 seconds) twilight flats
  - Do Dark and Bias corrections for each image
  - Average or median combine to get high S/N image
  - Make a copy and smooth with a large area filter (25 x 25 pixels) to remove pixel to pixel variations
  - Divide high S/N image by smoothed image
  - Normalize to 1.00 at the center
  - Save as a 32- bit real image
- The result is your "pixel flat"

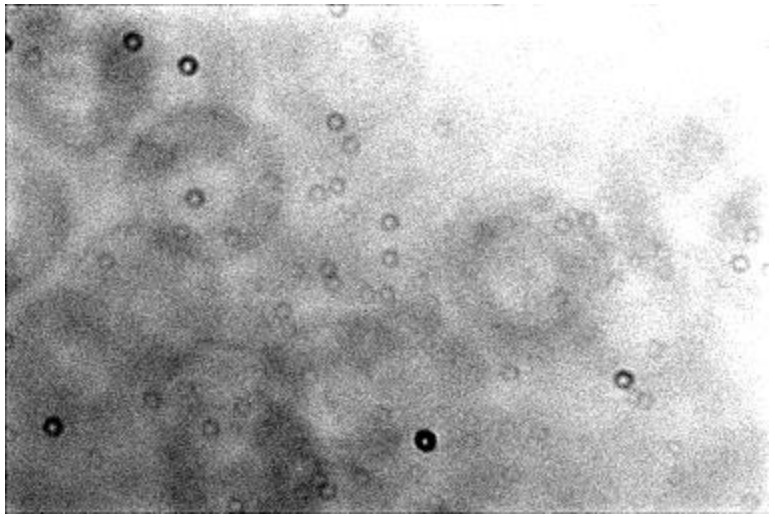


Image processing (9)

- Cleaned image after dark, bias, and flat-field corrections

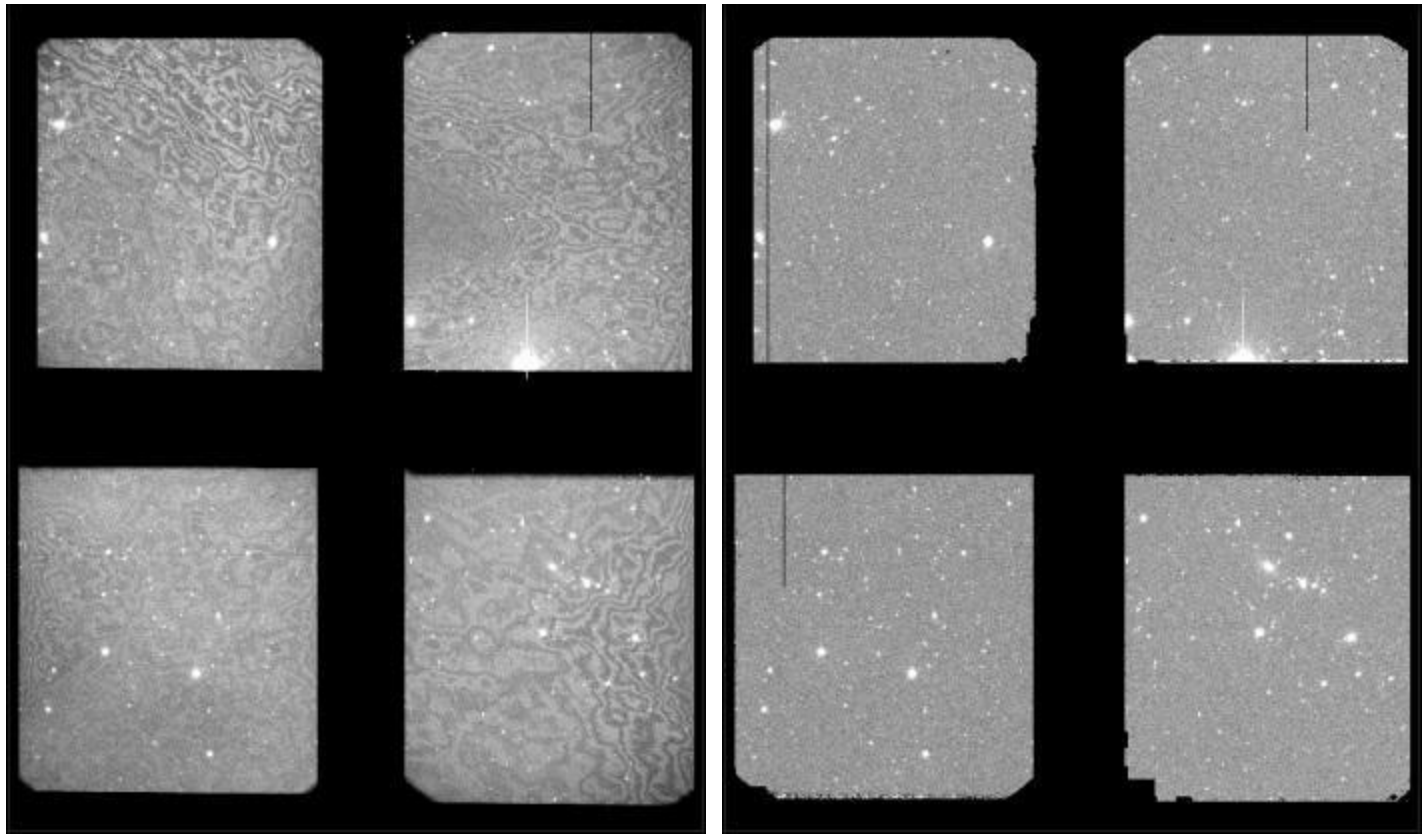


Image processing (10)

- **Summary:**
  - Start with raw image
  - Subtract dark and bias images
    - use high s/n “master bias” and “master dark” frames for best results
  - Divide by flat field
    - use high s/n “master flats” for best results
    - flats must have dark and bias removed



Image processing (11)

Do not forget! Cosmic Ray Events:



Image processing (12)

- There are several major computer packages:
  - IRAF
  - MIDAS
  - IDL (commercial)
  - Starlink

- Other Imaging Techniques are used in High-Energy Astronomy (we will not discuss the data reduction procedures adopted for high energy data):
  - HeaSoft



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 9

Vitaly Neustroev

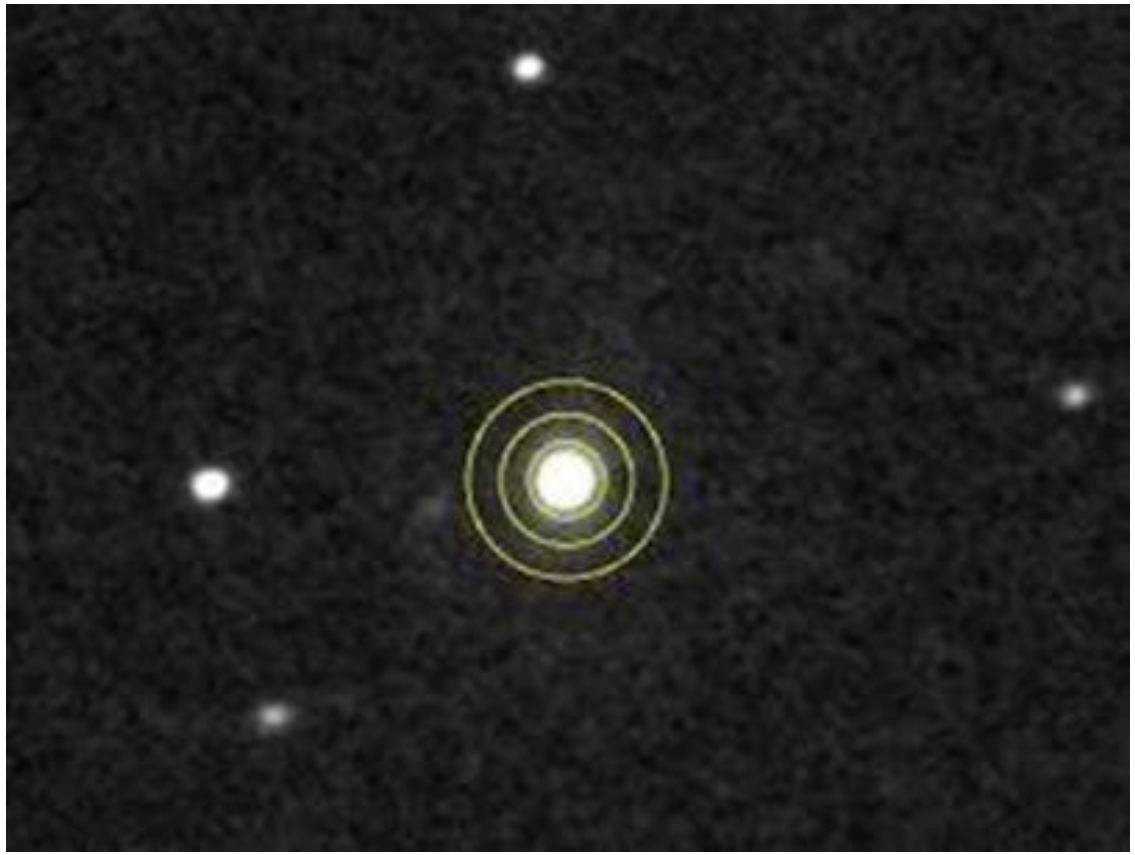
350

Photometry

The technique that measures the relative amounts of light in different wavelength ranges.

Simple aperture photometry

- We observe a star on a CCD detector, and process the data (i.e. correct for bias, dark, and flat field).
- An area centred on the star is defined to be the object area and is large enough to contain all the photons from that star.
- An equal area some distance away, which is found to be free of stars, is defined as the sky background area (sky aperture), and the sky background is measured from that.
- We then count all the photons in the object and sky background areas.
- Now substract the second from the first and obtain thus an estimate of the object flux in counts.



Photon flux can then be converted to physical units, e.g. ergs / (sec cm^2 Å).

Stellar magnitudes

- **Magnitudes:**
  - Apparent magnitudes $m_1 - m_2 = -2.5 \log \frac{E_1}{E_2}$

Precision of photometric measurements

Physical limitations on the precision of photometric measurements (1)

- What we get out of our detectors?
- Have we taken enough data?
- How much longer should we observe?

The important quantity that compares the level of a desired signal to the level of background noise is

the Output Signal-to-Noise Ratio (S/N)

Physical limitations on the precision of photometric measurements (2)

- Let's consider a stationary source of light (a star) with an average photon flux at the detector of N_* photons per second.
- The intensity of a source will produce the **average** number of photons, but the actual number collected will be more than, equal to, or less than the average, and their distribution about that average will be **a Poisson distribution**.

- The "counts" accumulated in a CCD pixel (or similar detectors) have **a Poisson distribution**.

Physical limitations on the precision of photometric measurements (3)

- The standard deviation of the photon noise is equal to the square root of the average number of photons (Poisson statistics). The ***Input*** **Signal-to-Noise Ratio** is then

$$S/N = \frac{N_*}{\sqrt{N_*}} = \sqrt{N_*}$$

where N_* is the average number of photons collected.

- When N_* is very large, the signal-to-noise ratio is very large as well. It can be seen that photon noise becomes more important when the number of photons collected is **small**.

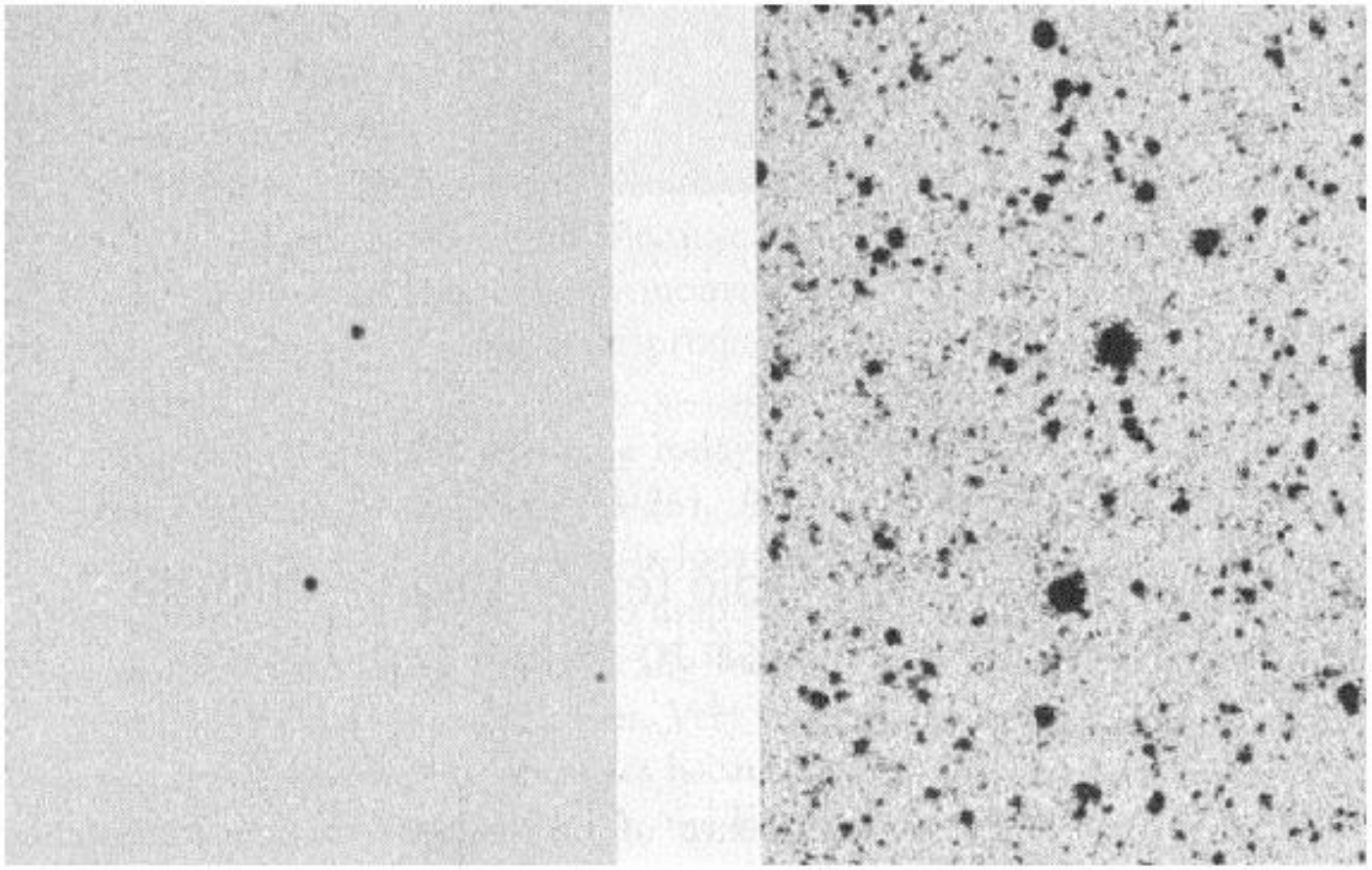
Physical limitations on the precision of photometric measurements (4)

- In the absence of sky background and readout noise it would be simple to calculate S/N, we would integrate for some time, detect some 10000 photons, and have a signal to noise ratio of 100 or so.

- So, if we ignore the sky background it's a simple calculation, **but we can't**.

Physical limitations on the precision of photometric measurements (5)

Sky and readout noise can be significant.



Physical limitations on the precision of photometric measurements (6)

- In dark sky at a dark site (no moon, no reflected street light), the magnitude of a 1 arcsecond patch of sky in the *V* band is approximately V_{sky}=21.5 mag.

Thus, every square arcsecond of sky gives:

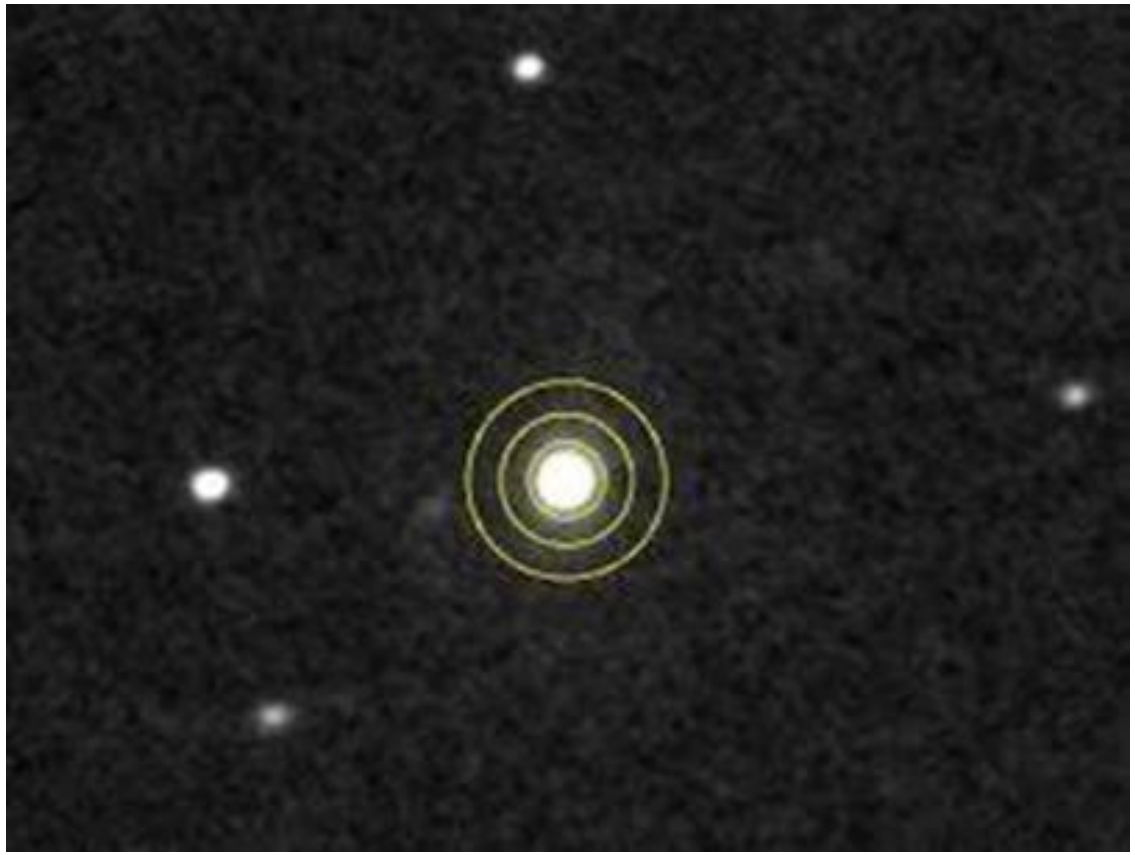
$S_{sky} = 2.5 \times 10^{-3}$ photons / (cm^2 $arcsec^2$ second)

Physical limitations on the precision of photometric measurements (7)

- To calculate the ***Output* Signal-To-Noise Ratio** of an observation we need to know the signal, and all sources of noise. These are:
  - Photon noise (shot noise) from the signal;
  - Photon noise from the sky background under the signal;
  - Photon noise from the sky background measurement to be subtracted off;
  - Readout noise from all sources;
  - Fixed pattern noise;
  - Bias noise;
  - Dark current noise.

A case study of simple aperture photometry (1)

- We observe a star on a CCD detector, and process the data in the simplest way possible.
- An area centred on the star is defined to be the object area and is large enough to contain all the photons from that star.
- An equal area some distance away, which is found to be free of stars, is defined as the sky background area (sky aperture), and the sky background is measured from that.



A case study of simple aperture photometry (2)

- We will make some assumptions:
  - We have eliminated fixed pattern noise by dividing the image by a normalised long exposure of a uniform light source, this is called *a flat field*.
  - Bias noise and dark current noise are negligible, as this is a cryogenically cooled, buried channel CCD.

Aperture photometry

- There are a number of parameters we need to take into account to calculate the signal which reaches the detector:
  - t – exposure time
  - β – angular size of a source (defined by the seeing)
  - D – diameter of the telescope
  - S_{sky} [photons / (cm^2 arcsec^2 second)] – brightness of the sky
  - η – quantum efficiency of a detector (QE)
  - f_* [photons / (cm^2 second)] – the source flux to be measured

Signal calculation (1)

- We start from the number of photons incident from this star, from the sky, and from the star + the sky:
  - $A \sim D^2$ is the telescope collecting area [cm^2]
  - $B \sim \beta^2$ is the source area on the sky [arcsec^2]

- $n_* \approx \eta\, D^2\, t\, f_*$ – an average number of photons from the source
- $n_{sky} \approx \eta\, D^2\, t\, \beta^2\, S$ – an average number of photons from the sky
- $n_{*+sky} \approx \eta\, D^2\, t\, (f_* + \beta^2\, S)$ – an average number of photons from the source and the sky

Signal calculation (2)

- That is without the Readout noise and other detector noises. If we want to take them into account – we must add $N_d = n_d\, t$ to the right side of equations.

- There is also a geometric efficiency factor as part of the aperture of the telescope is blocked by the secondary mirror.

For a while, we will not take these factors into account.

Noise on the measurements

- Noise on the measurements is given by the square root of the number of photons:

$$\sigma_{*+sky} = \sqrt{n_{*+sky}}$$

$$n_* \approx n_{*+sky} - n_{sky}$$

$$\sigma_* = \sqrt{n_{*+sky} + n_{sky}} = \sqrt{n_* + 2n_{sky}}$$

(If x and y have independent random errors σ_x and σ_y , then the error in $z = x \pm y$ is $\sigma_z^2 = \sigma_x^2 + \sigma_y^2$)

Signal to Noise ratio (1)

$$n_* \approx \eta\, D^2\, t\, f_*$$

$$S/N = \frac{n_*}{\sigma_*} = \frac{n_*}{\sqrt{n_* + 2 n_{sky}}} = \frac{\eta\, D\, t\, f_*}{\sqrt{\eta\, t\, (f_* + 2\, \beta^2\, S)}}$$

$$n_{sky} \approx \eta\, D^2\, t\, \beta^2\, S$$

Let's now consider two special cases:

- If the Source dominates over the Sky: $n_* \gg n_{sky}$
- If the Sky noise dominates: $n_{sky} \gg n_*$

Signal to Noise ratio (2)

$$S/N = \frac{n_*}{\sqrt{n_* + 2n_{sky}}} = \frac{\eta D t f_*}{\sqrt{\eta t (f_* + 2\beta^2 S)}}$$

If the Source dominates over the Sky: $n_* \gg n_{sky}$

$$S/N \cong \frac{n_*}{\sqrt{n_*}} = \sqrt{n_*} = D\sqrt{\eta t f_*}$$

$$f_{min} \sim 1 / (D^2 t)$$ for the given S/N

the telescope aperture is most important!

Signal to Noise ratio (3)

$$S/N = \frac{n_*}{\sqrt{n_* + 2n_{sky}}} = \frac{\eta D t f_*}{\sqrt{\eta t (f_* + 2\beta^2 S)}}$$

If the Sky noise dominates: $n_{sky} \gg n_*$

$$S/N \cong \frac{n_*}{\sqrt{2n_{sky}}} = \frac{\eta D t f_*}{\sqrt{2 \eta t \beta^2 S}} = \frac{D f_*}{\beta} \sqrt{\frac{\eta t}{2 S}}$$

$$f_{min} \simeq \frac{\beta}{D} \sqrt{\frac{S}{t}}$$ for the given S/N

the seeing (angular size of a source) is most important!

Signal to Noise ratio (4)

Source dominates over the Sky

$$S/N \cong D\sqrt{\eta\, t\, f_*}$$

$$f_{min} \sim \frac{1}{D^2 t}$$

for the given S/N

most important is

the telescope aperture

Sky noise dominates

$$S/N \cong \frac{D\, f_*}{\beta}\sqrt{\frac{\eta\, t}{2\, S}}$$

$$f_{min} \sim \frac{\beta}{D}\sqrt{\frac{S}{t}}$$

for the given S/N

the seeing

Stellar magnitudes

- **Magnitudes:**
  - Apparent magnitudes $m_1 - m_2 = -2.5 \log \frac{E_1}{E_2}$
  - Absolute magnitudes

$$M - m = -2.5 \log \left(\frac{D}{10}\right)^2$$

D is the object's distance in parsecs

$$M = m + 5 - 5 \log D$$

$$M = m + 5 - 5 \log D \ - \mathrm{A} \cdot \mathrm{D}$$

A is the interstellar absorption in magnitudes per parsec. Within the galactic plane A is ~0.002 mag pc^{-1}.

- Sometimes M may be estimated by some independent method. Then:

$$D = 10^{[(m-M+5)/5]} \text{ pc}$$

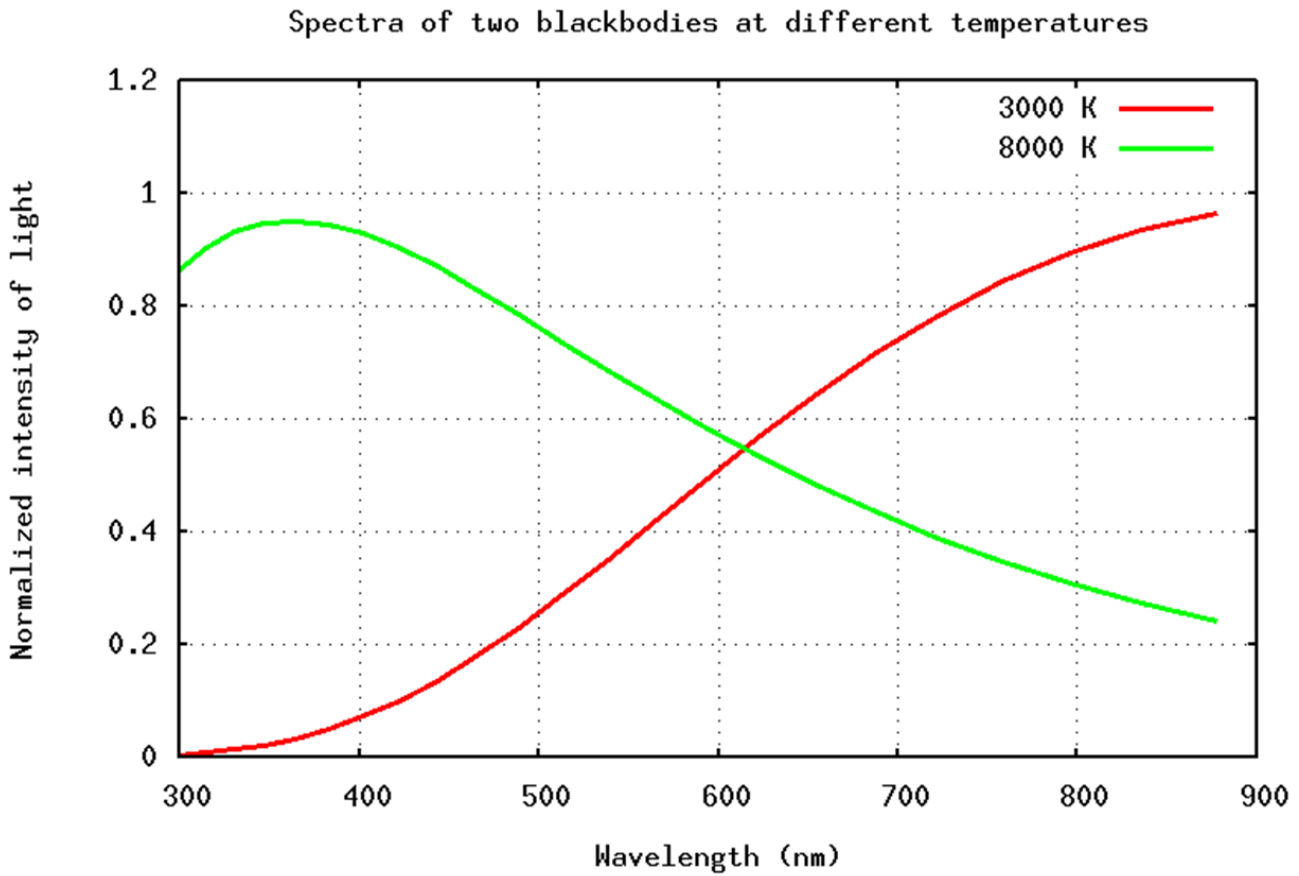
Filters and photometric systems

- Filter (photometric) systems:
  - Filters are used to restrict the wavelengths of electromagnetic radiation that hit the detector.
- Why may we want to do that?
  - Because stars have different colours that means they have different temperatures.

Observing through filters (1)

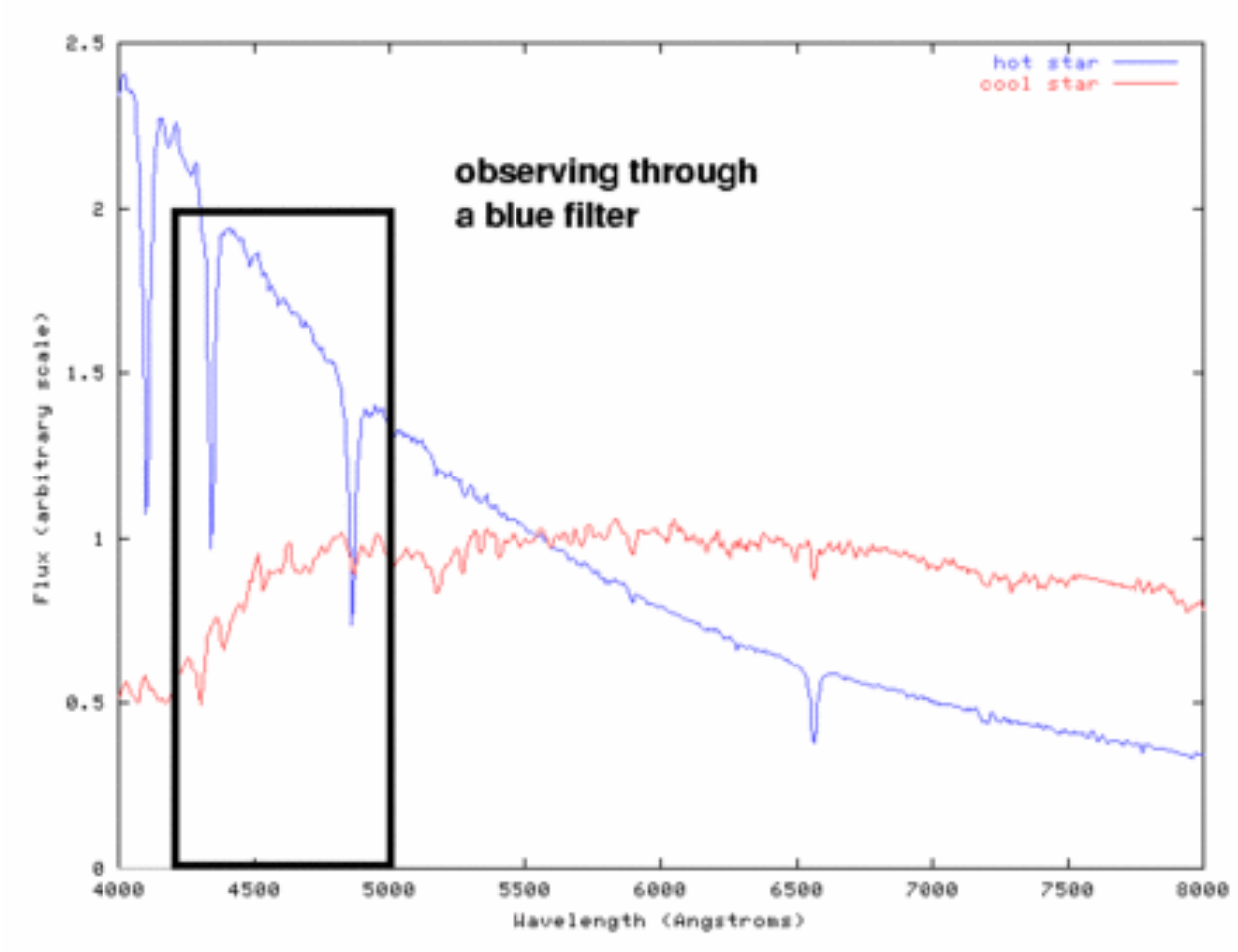
Hot objects emit most of their light at short wavelengths

Cool objects emit most of their light at long wavelengths



Observing through filters (2)

Observing through filters allows us to estimate temperatures.



Observing through filters (3)

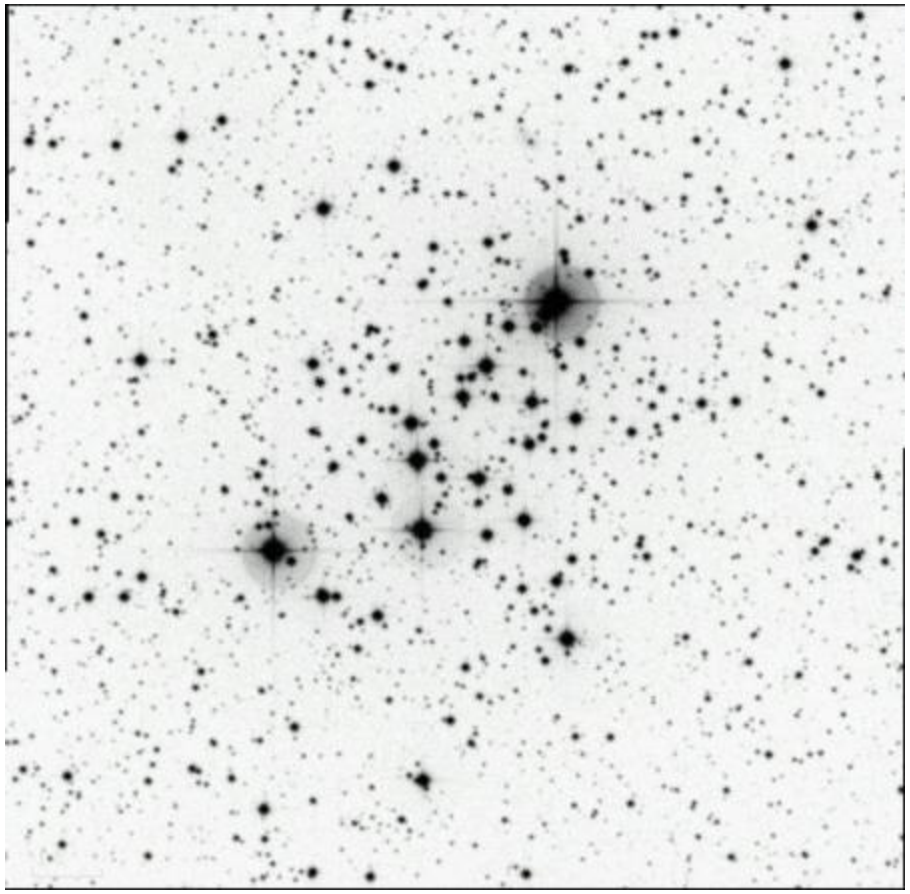
Which are the three brightest stars?



Observing through filters (4)

Which are the brightest stars?

It depends on
the bandpass through which
one observes them.



Photometric systems

- There is a number of different **photometric systems**, each one based on a particular passband (i.e. a particular combination of filter and detector and telescope).
- They may be grouped into wide, intermediate, and narrowband systems according to the bandwidth of their transmission curves. In the visible region:
  - Wide (broadband) filters have bandwidths of ~1000 Å
  - Intermediate: 100-500 Å
  - Narrowband filters range from 0.5 to 100 Å.
- One should always remember to specify the system when quoting the magnitude of a star.

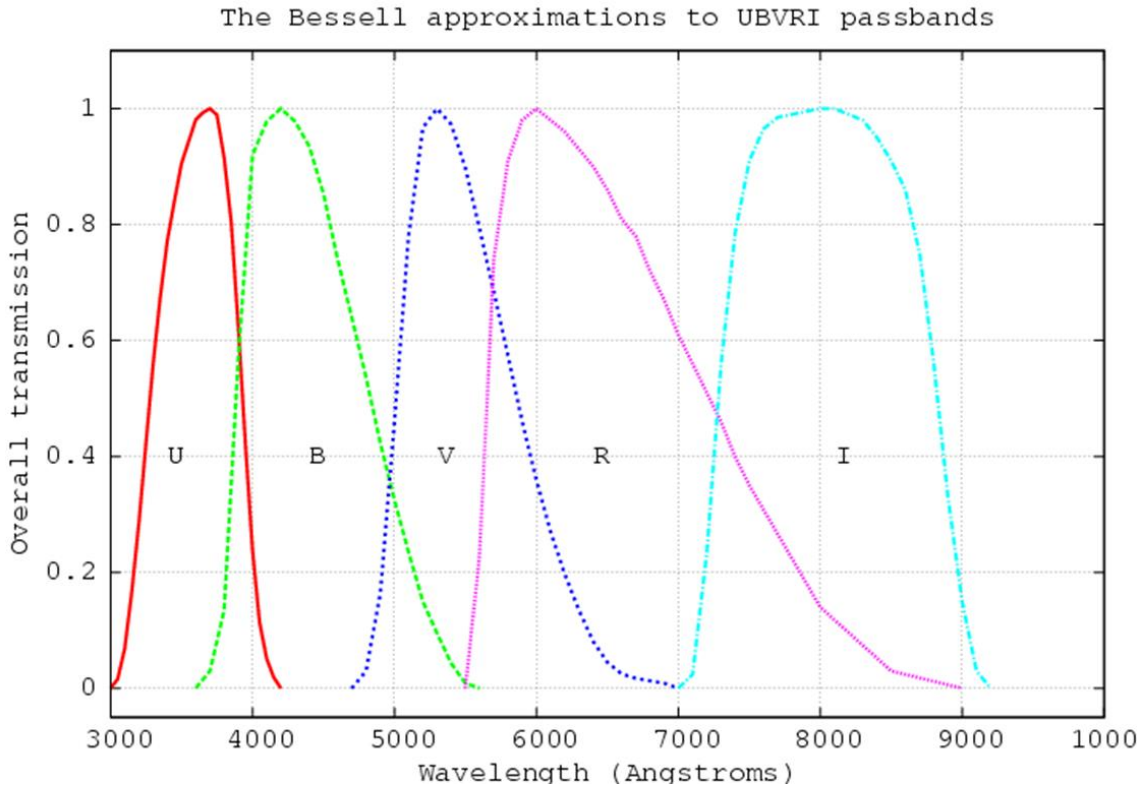
Johnson-Cousins photometric system (UBVRI)

- Most astronomers working in the optical use the Johnson-Cousins **UBVRI** photometric systems:
  - Johnson and Morgan defines the **UBV** system with stars visible in the northern hemisphere
  - Cousins defines the redder **R** and **I** passbands.

- The systems are defined by particular combinations of glass filters and photomultiplier tubes (they were created many years ago before CCDs existed). Since photomultipliers and CCDs have very different spectral sensitivities, it is difficult to make the effective passband of a CCD-based instrument match that of a photomultiplier-based instrument.
- In 1990, Michael Bessell came up with a recipe for making filters out of common colored glasses which would reproduce pretty closely the official Johnson-Cousins **UBVRI** passbands → Bessell filters.

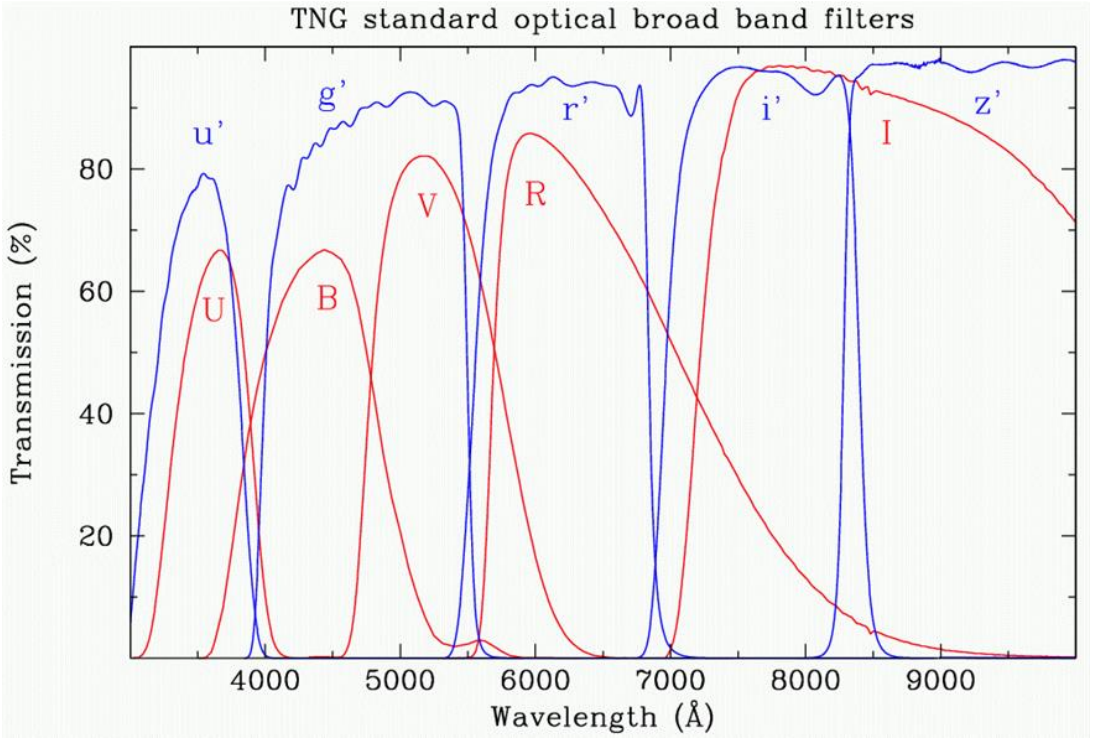
Johnson-Cousins photometric system

- The spectral resolution of the broadband UBVRI passbands is small: $R = \lambda/\Delta\lambda \approx 5$



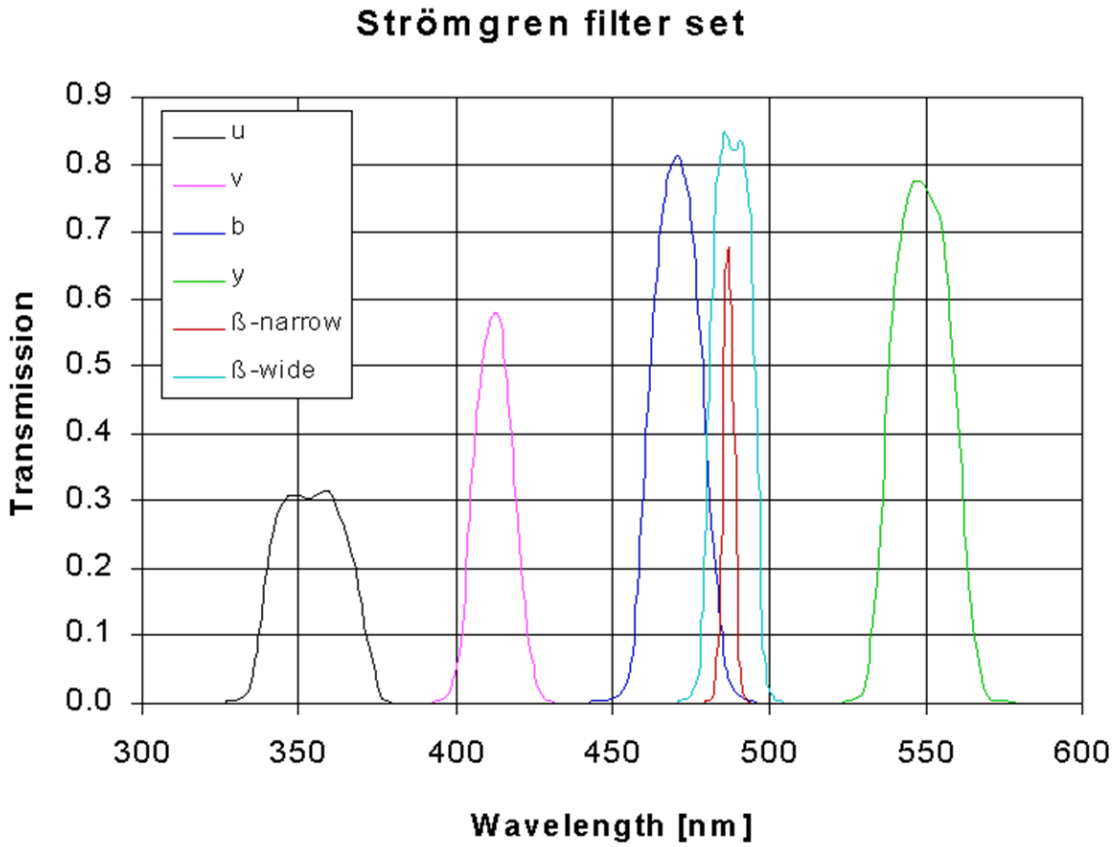
SDSS (ugriz) photometric system

- Although **UBVRI** is the best known optical system, there are a number of others. Some were specifically designed to solve a particular astrophysical problem, others to mesh with particular detectors.
  One important system is the **u'g'r'i'z'** that is being used by the Sloan Digital Sky Survey (**SDSS**) and the Panoramic Survey Telescope and Rapid Response System (Pan-STARRS). It has become very popular recently.



Strömgren photometric system

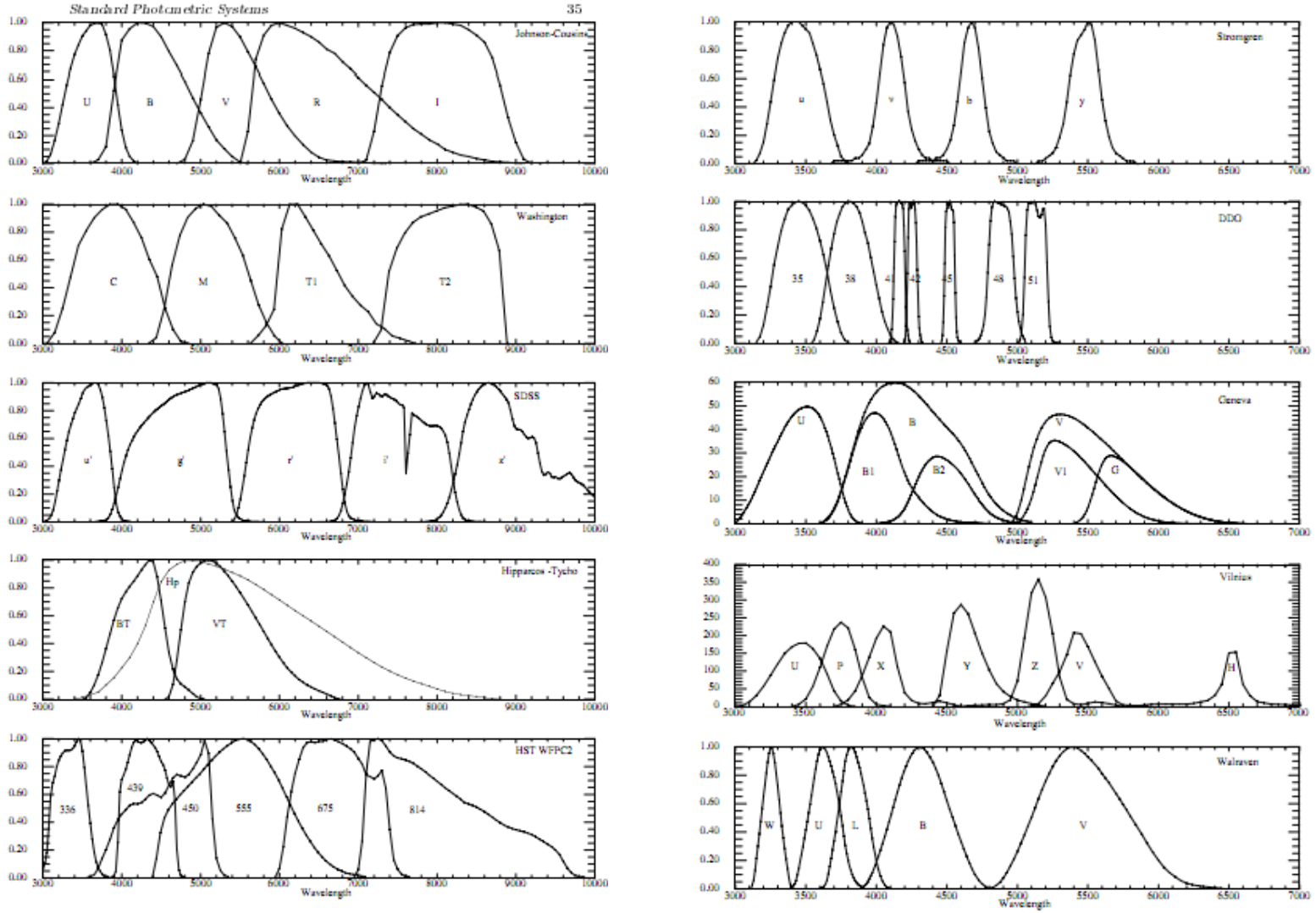
- **Strömgren photometric system** (uvby) is four-colour intermediate-band photometric system (plus Hβ filters) for stellar classification. It was pioneered by the Danish astronomer Bengt Strömgren in 1956.



Narrowband photometric systems

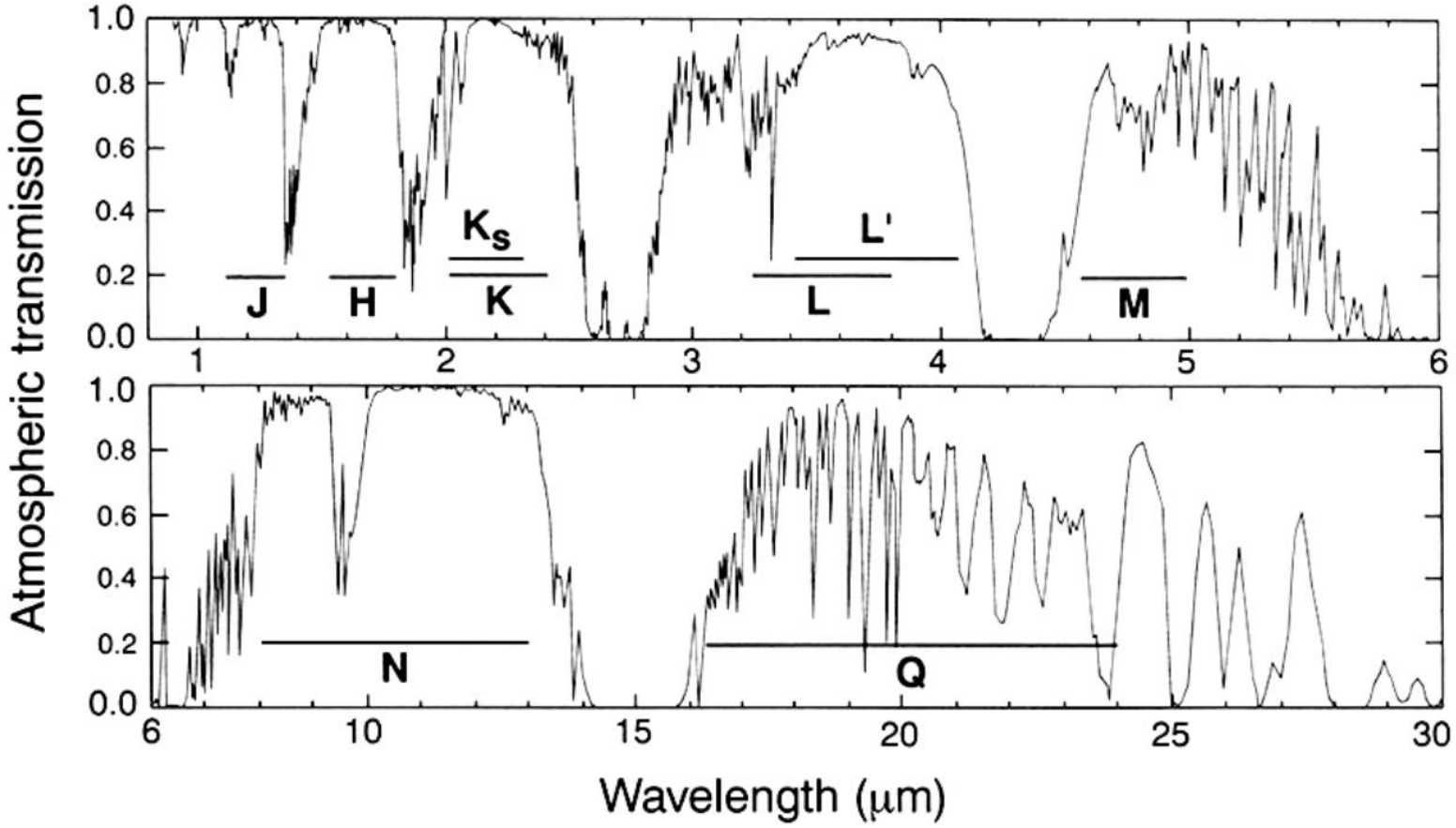
- For some applications, astronomers use narrowband filters; a common filter used to measure light emitted by hydrogen atoms is centered at 6563 Angstroms and roughly 20 Angstroms wide: $R = \lambda/\Delta\lambda \approx 330$
- A narrowband filter like this requires much longer exposure times to build up the same signal as a broadband filter. Since telescope time is so precious, astronomers tend to use broadband systems.
  That's one reason for the popularity of the UBVRI or SDSS systems.

Photometric systems (optical)



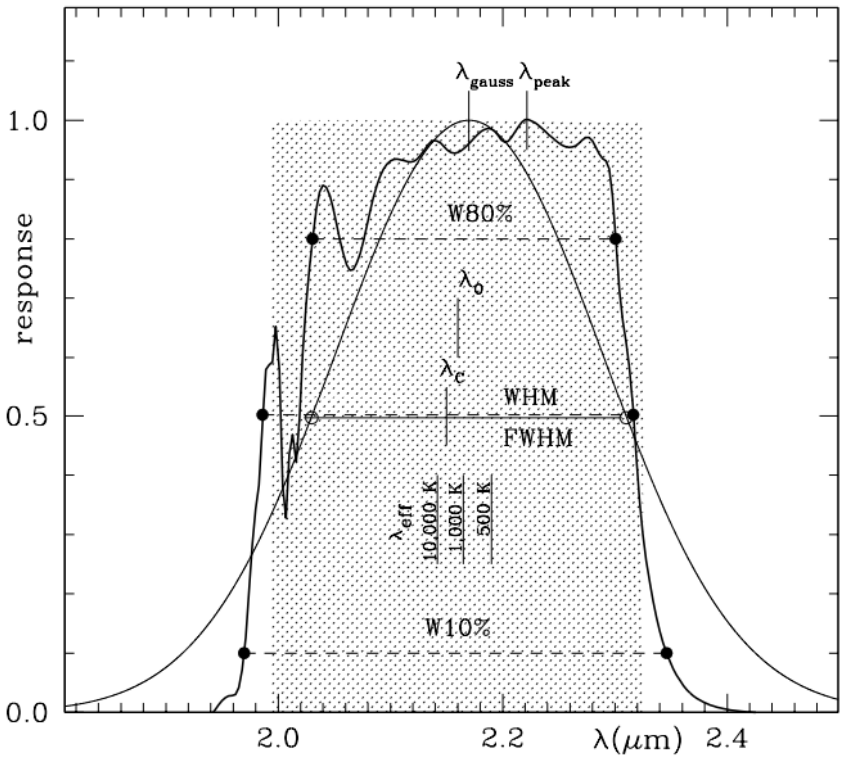
The Infrared Photometric Bands: JHK+others

… where the atmospheric transmission windows are



Filter transmission curves (1)

- Typical broad-band transmission curves are **not** rectangular, and even **not** symmetric.
- Different quantities can be used to describe a filter, e.g.:
  - λ_c is the wavelength halfway between the points, where the band transmission profile reaches half of the maximum value.
  - WHM is the the full wavelength span between the points, where the band transmission profile reaches half of the maximum value.
  - λ_{peak} is the wavelength at which the band transmission profile reaches its maximum.



From Fiorucci and Munari, 2003, A&A, 401, 781

Filter transmission curves (2)

- Some important parameters depend on the source spectrum. For example,
  - λ_0 is the mean wavelength of the band, the property of just a band:

$$\lambda_o = \frac{\int \lambda F(\lambda)\, d\lambda}{\int F(\lambda)\, d\lambda}.$$

  - whereas the effective wavelength λ_{eff} is

$$\lambda_{eff} = \frac{\int \lambda F(\lambda) S(\lambda)\, d\lambda}{\int F(\lambda) S(\lambda)\, d\lambda}.$$

where

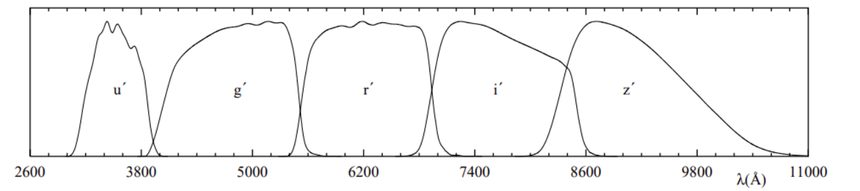
$F(\lambda)$ is the transmission profile of the band, and

$S(\lambda)$ the energy distribution of a source spectrum.

Filter transmission curves (3)

Good sources of info:

- *The Asiago Database on Photometric Systems* (218 systems; checked on 2025-09-18)
  - http://ulisse.pd.astro.it/Astro/ADPS
  - Fiorucci and Munari, 2003, A&A, 401, 781
- *Filter Profile Service* (11013 filters available on 2025-09-18)
  - http://svo2.cab.inta-csic.es/theory/fps/

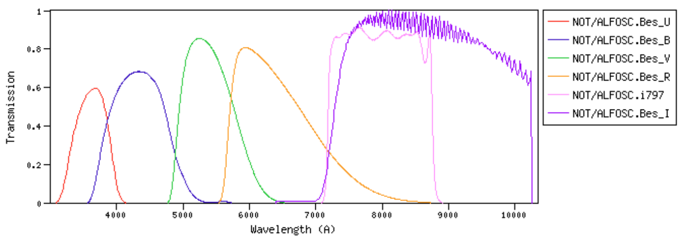


u′									
$\lambda_c = 3530$, WHM = 642, $W_o = 590$, $\mu = 201$, $I_{asym} = 0.01$, $I_{kurt} = -0.88$	$\lambda_o = 3521$, W10% = 831	$\lambda_{peak} = 3431$, W80% = 437	$\lambda_{gauss} = 3519$, FWHM = 555	B3 3504 [599]	Vega 3551 [602]	Sun 3538 [565]	K2 3593 [517]	M2 3636 [448]	Carbon 3525 [498]
				WN 3489	WC 3500	PN_{Ne} 3642	PN^{Ne} 3533	Nova 3517	WDA 3481
	$\lambda_{eff} = 3521.5 + 41.9 \times E(B-V)$ r=1.00			$W_{eff} = 600.1 - 38.8 \times E(B-V)$ r=−0.98					
	$\lambda_{eff}(T) = 3472 + 145 \times \theta + 77 \times \theta^2 - 55 \times \theta^3$			$W_{eff}(T) = 556 + 263 \times \theta - 562 \times \theta^2 + 193 \times \theta^3$					
g′									
$\lambda_c = 4788$, WHM = 1411, $W_o = 1325$, $\mu = 419$, $I_{asym} = -0.12$, $I_{kurt} = -1.04$	$\lambda_o = 4803$, W10% = 1641	$\lambda_{peak} = 5173$, W80% = 1111	$\lambda_{gauss} = 4820$, FWHM = 1245	B3 4683 [1238]	Vega 4708 [1271]	Sun 4817 [1318]	K2 4903 [1230]	M2 5015 [1057]	Carbon 5100 [1004]
				WN 4696	WC 4706	PN_{Ne} 4943	PN^{Ne} 4767	Nova 4923	WDA 4716
	$\lambda_{eff} = 4807.4 + 142.3 \times E(B-V)$ r=1.00			$W_{eff} = 1371.8 - 208.0 \times E(B-V)$ r=−0.99					
	$\lambda_{eff}(T) = 4647 + 312 \times \theta + 241 \times \theta^2 - 173 \times \theta^3$			$W_{eff}(T) = 1156 + 909 \times \theta - 1424 \times \theta^2 + 387 \times \theta^3$					
r′									
$\lambda_c = 6242$, WHM = 1387, $W_o = 1343$, $\mu = 407$, $I_{asym} = -0.01$, $I_{kurt} = -1.09$	$\lambda_o = 6253$, W10% = 1565	$\lambda_{peak} = 6191$, W80% = 1248	$\lambda_{gauss} = 6247$, FWHM = 1262	B3 6160 [1282]	Vega 6168 [1294]	Sun 6220 [1335]	K2 6256 [1341]	M2 6307 [1315]	Carbon 6365 [1251]
				WN 6220	WC 6124	PN_{Ne} 6531	PN^{Ne} 6432	Nova 6444	WDA 6156
	$\lambda_{eff} = 6253.4 + 91.0 \times E(B-V)$ r=1.00			$W_{eff} = 1370.2 - 106.1 \times E(B-V)$ r=−0.98					
	$\lambda_{eff}(T) = 6145 + 139 \times \theta + 156 \times \theta^2 - 80 \times \theta^3$			$W_{eff}(T) = 1255 + 289 \times \theta - 183 \times \theta^2 - 109 \times \theta^3$					
i′									
$\lambda_c = 7704$, WHM = 1532, $W_o = 1374$, $\mu = 453$, $I_{asym} = 0.14$, $I_{kurt} = -1.05$	$\lambda_o = 7667$, W10% = 1756	$\lambda_{peak} = 7242$, W80% = 1005	$\lambda_{gauss} = 7635$, FWHM = 1291	B3 7573 [1322]	Vega 7584 [1335]	Sun 7620 [1359]	K2 7649 [1369]	M2 7732 [1340]	Carbon 7653 [1453]
				WN 7582	WC 7625	PN_{Ne} 7495	PN^{Ne} 7608	Nova 7623	WDA 7582
	$\lambda_{eff} = 7666.9 + 78.9 \times E(B-V)$ r=1.00			$W_{eff} = 1391.2 - 65.7 \times E(B-V)$ r=−0.97					
	$\lambda_{eff}(T) = 7562 + 101 \times \theta + 123 \times \theta^2 - 52 \times \theta^3$			$W_{eff}(T) = 1310 + 144 \times \theta - 9 \times \theta^2 - 104 \times \theta^3$					
z′									
$\lambda_c = 9038$, WHM = 1408, $W_o = 1411$, $\mu = 536$, $I_{asym} = 0.48$, $I_{kurt} = -0.39$	$\lambda_o = 9115$, W10% = 2212	$\lambda_{peak} = 8717$, W80% = 845	$\lambda_{gauss} = 9018$, FWHM = 1326	B3 9022 [1379]	Vega 9046 [1390]	Sun 9057 [1400]	K2 9086 [1412]	M2 9136 [1410]	Carbon 9076 [1562]
				WN 9027	WC 9016	PN_{Ne} 8894	PN^{Ne} 9022	Nova 9068	WDA 8993
	$\lambda_{eff} = 9113.4 + 73.5 \times E(B-V)$ r=1.00			$W_{eff} = 1422.0 - 42.8 \times E(B-V)$ r=−0.97					
	$\lambda_{eff}(T) = 8997 + 88 \times \theta + 105 \times \theta^2 - 36 \times \theta^3$			$W_{eff}(T) = 1357 + 91 \times \theta + 24 \times \theta^2 - 76 \times \theta^3$					

NOT filters: ALFOSC | NOT | NOTcam | StanCam

Filter ID	λ_{ref}	λ_{mean}	λ_{eff}	λ_{min}	λ_{max}	W_{eff}	ZP_ν	ZP_λ	Obs. Facility	Instrument	Description
NOT/ALFOSC.Bes_U	3600.85	3617.41	3670.73	3102.79	4129.62	580.28	1758.31	4.07e-9	NOT	ALFOSC	Bessell U
NOT/ALFOSC.Bes_B	4306.12	4346.66	4319.73	3579.81	5682.69	1004.43	3923.93	6.34e-9	NOT	ALFOSC	Bessell B
NOT/ALFOSC.Bes_V	5389.63	5417.18	5365.72	4786.00	6447.52	885.24	3670.94	3.79e-9	NOT	ALFOSC	Bessell V
NOT/ALFOSC.Bes_R	6396.64	6464.12	6329.59	5551.69	8522.76	1279.53	3085.76	2.26e-9	NOT	ALFOSC	Bessell R
NOT/ALFOSC.i797	7927.59	7966.41	7886.27	7103.45	8872.45	1499.39	2435.41	1.16e-9	NOT	ALFOSC	interference i
NOT/ALFOSC.Bes_I	8559.60	8682.26	8466.07	6392.60	10246.30	2578.97	2338.38	9.57e-10	NOT	ALFOSC	Bessell I.

Filter Plots



Magnitudes & Photometric systems

- When writing the magnitude of a star, astronomers use an abbreviation to denote the photometric system of the measurement:
  - V = 1.03 (or 1.03V) means "magnitude of this star in the V filter is 1.03"
  - B = 0.46 (or 0.46B) means "magnitude of this star in the B filter is 0.46"
  - g'=15.21 means "magnitude of this star in the SDSS g' filter is 15.21"

But a magnitude system can be different!



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 10

Vitaly Neustroev

Magnitude systems

$$m_1 - m_2 = -2.5\, log \frac{F_1}{F_2} \quad \text{or} \quad m = -2.5\, log \frac{F}{F_0}$$

where the flux F_0 defines the reference or **zeropoint** of the magnitude scale.

The choice is arbitrary!

- Standardizing magnitudes (magnitude systems):
  - Vega system
  - AB system
  - ST Magnitudes

A magnitude system is **not** a photometric (filter) system
(you can use a filter in any system)

Photometry: Vega system

- Astronomers have chosen to use the bright star Vega (α Lyr) as their starting point.
- In the UBVRI systems, the star Vega is **defined** to have a magnitude of zero in all bands (actually, this is not quite true):

$$U = 0.0;\ B = 0.0;\ V = 0.0;\ R = 0.0;\ I = 0.0$$

- This means also that all the colours of Vega are zero.
- The zero-point of this system depends on the flux of Vega (outside the atmosphere) and is different in different bands.

Photometry: AB system

- In the **AB system**, which is not based on Vega, it is assumed that the flux constant F_0 is the **same** for all wavelengths and passbands.
- That constant is per definition such that in the V filter: $m_V^{Vega} = m_V^{AB} = 0$ (or more accurately: $F_\lambda \, d\nu \equiv F_\lambda \, d\lambda$ when averaged over the V filter, or at the effective wavelength of the V filter, $\lambda_{eff} = 5480$ Å. Based on the work of Oke (1974), then

$$m_\nu = -2.5 \, log \, F_\nu - (48.585 \pm 0.005)$$

where $F_\nu(\lambda)$ is the spectral flux density per unit frequency of a source at the top of the Earth's atmosphere in units of erg $s^{-1}cm^{-2}$ Hz^{-1}.

Photometry: AB system

- Note that the AB magnitude system is expressed in F_ν rather than F_λ!

  The flux density in F_ν is related to the flux density in F_λ by:

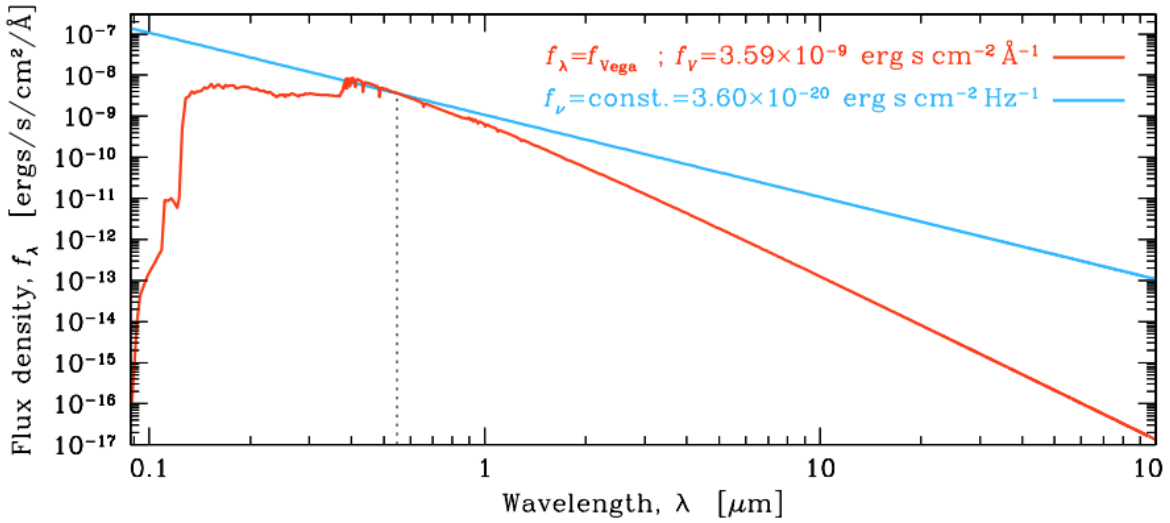
$$F_\nu\,[\text{ergs s}^{-1}\ \text{cm}^{-2}\ \text{Hz}^{-1}\ ] = 10^{-8}\,\frac{\lambda[\text{Å}]^2}{c[\text{cm s}^{-1}]}\,F_\lambda\,[\text{ergs s}^{-1}\ \text{cm}^{-2}\ \text{Å}^{-1}]$$

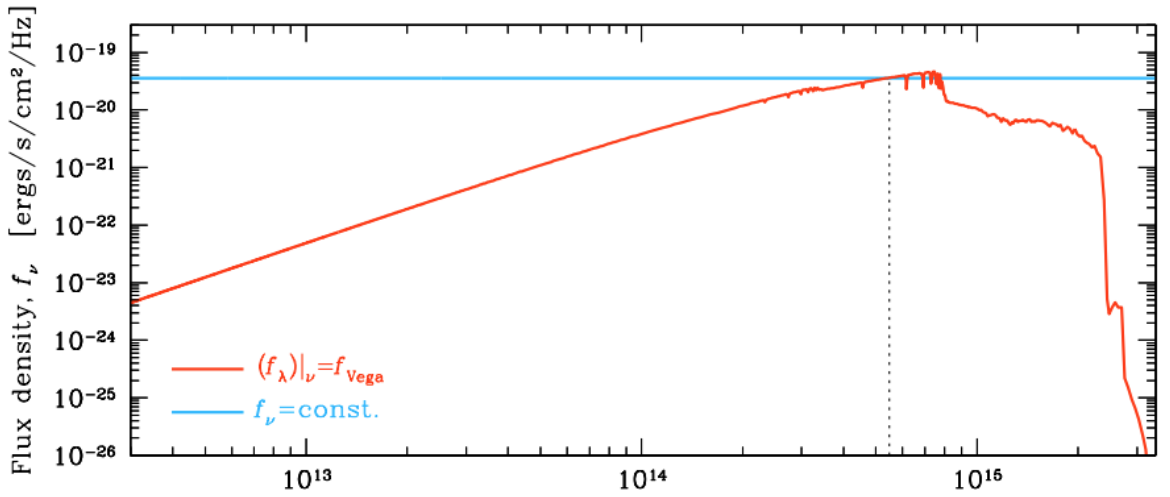
- One can easily convert between AB magnitudes and Janskys:
  In AB magnitudes, mag 0 has a flux of 3631 Jy.

AB and VEGA systems compared

- The difference between AB and Vega magnitudes becomes very large at redder wavelengths!

- The spectrum of Vega is very complicated at IR wavelengths and often model atmospheres are used adding to uncertainties





Photometry: ST Magnitudes

The ST magnitude system is defined such that an object with constant flux $F_\lambda = 3.63 \times 10^{-9}$ ergs s^{-1} cm^{-2} $Å^{-1}$ will have magnitude ST = 0 in every filter. In general,

$$ST_{mag} = -2.5\, log\, F_\lambda - 21.1$$

We will not discuss this system anymore.

Bolometric magnitudes

- **Bolometric magnitudes:** this gives a magnitude corresponding to the total flux integrated over all wavelengths
- The calculations are expressed as the difference between the bolometric magnitude and observed magnitude. The difference is then known as the bolometric correction: $BC = m_{bol} - V$
- The XXIXth IAU General Assembly in Honolulu recommended zero points for the absolute and apparent bolometric magnitude scales:
  - Resolution B2 defines the zero point of the absolute bolometric magnitude scale such that a radiation source with $M_{Bol}=0$ has luminosity $L_0=3.0128\times10^{28}$ W.
  - The zero point of the apparent bolometric magnitude scale ($m_{bol}=0$) corresponds to irradiance $F_{Bol} = 2.518 \times 10^{-8}$ W m^{-2}. The zero points were chosen so that the nominal solar luminosity (3.828×10^{26} W) corresponds to $M_{Bol}(\text{Sun}) = 4.74$.
  - The nominal total solar irradiance (1361 W m^{-2}) corresponds approximately to apparent bolometric magnitude $m_{bol}(\text{Sun}) = -26.832$.

Standard Stars for Photometry (1)

- The primary standards for the UBV system are a set of 10 bright, naked eye stars of magnitude 2 to 5, known as the North Pole sequence – comprise stars within 2° of the North pole star. The magnitudes of these stars define the UBV colour system.
- Instead of using the primary standards directly, we use a series of secondary standard stars, or just standard stars, whose magnitudes have been carefully measured relative to the primary stars.
- For broadband optical work (UBVRI filter system) the standard stars used most frequently today are from the work of the astronomer Arlo Landolt. Landolt has devoted many years to measuring a set of standard star magnitudes.

Standard Stars for Photometry (2)

- What makes a good standard star?
  - A standard star must not be variable!
  - Standard stars must be of a brightness that will not overwhelm the detector and telescope in use, but must be bright enough to give a good S/N in a short exposure. For very large telescopes, many of the Landolt stars are too bright.
  - Ideally, a set of stars very close together in the sky will cover a wide range of colours.
  - Standard stars should be located across the sky so that they span a wide range of airmass.

Colour indices (1)

- **Colour indices**: this is the difference between magnitudes at two separate wavelengths:

$$C_{BV} = B - V;\ C_{VR} = V - R,\ \text{and so on.}$$

- International colour index (outdated, but can be found in the literature) based upon photographic and photovisual magnitudes:

$$m_p - m_{pv} = C = B - V - 0.11$$

Colour indices (2)

- The B – V colour index is closely related to the spectral type with an almost linear relationship for main sequence stars.
- For most stars, the B and V regions are located on the long wavelength side of the maximum spectral intensity.
- If we assume that the effective wavelengths of the B and V filters are 4400 and 5500 Å, then using the Planck equation:

$$L_\lambda(T) = \frac{2\,h\,c_0^2}{\lambda^5}\left[\exp\left(\frac{h\,c_0}{\lambda\,k_B T}\right) - 1\right]^{-1}$$

we obtain:

$$B - V \approx -2.5\log\left[3.05\,\frac{\exp(2.617\times 10^4/T)}{\exp(3.27\times 10^4/T)}\right]$$

Colour indices (3)

- For T < 10000 K this is approximately

$$B - V \approx -2.5 \log\left[3.05 \frac{\exp(2.617 \times 10^4/T)}{\exp(3.27 \times 10^4/T)}\right] = -1.21 + \frac{7090}{T}$$

The magnitude scale is an arbitrary one.
For T = 9600 K (Vega temperature), B–V = 0.0,
but we have obtained ~0.5. Using this correction, we get:

$$T = \frac{7090}{(B - V) + 0.74}\ K$$

Colour excess and Interstellar absorption

- More distant stars are affected by interstellar absorption, and since this is strongly inversely dependent upon wavelength.
- The colour excess measure the degree to which the spectrum is reddened:

$$E_{U-B}=(U-B)-(U-B)_0$$
$$E_{B-V}=(B-V)-(B-V)_0$$

where the subscript 0 denotes unreddened quantities – intrinsic colour indices.

- In the optical spectrum, interstellar absorption varies with both wavelength and the distance like this semi-empirical relationship:

$$A_\lambda=6.5\times10^{-10}/\lambda-2.0\times10^{-4}\ \text{mag pc}^{-1}$$

where λ is in nanometers

Photometry

- Simple UBV photometry for hot stars results in determinations of temperature, Balmer discontinuity, spectral type, and reddening. From the latter we can estimate distance.
- Thus, we have a very high return of information for a small amount of observational effort. This is why the relatively crude methods of wideband photometry is so popular.

Photometry

Effective wavelengths (for an A0 star like Vega), absolute fluxes (corresponding to zero magnitude) and zeropoint magnitudes for the UBVRIJHKL Johnson-Cousins system

Bessell et al.
(1998, A&A, 333, 231)

Band	λc (Å)	f_{v0}	$f_{\lambda 0}$	$zp(f_\lambda)$	$zp(f_v)$
U	3660	1.790	417.5	-0.152	0.770
B	4380	4.063	632.0	-0.602	-0.120
V	5450	3.636	363.1	0.000	0.000
R	6410	3.064	217.7	0.555	0.186
I	7980	2.416	112.6	1.271	0.444
J	12200	1.589	31.47	2.655	0.899
H	16300	1.021	11.38	3.760	1.379
K	21900	0.64	3.961	4.906	1.886
L	34500	0.285	0.708	6.775	2.765

f_λ (10^{-11} erg $s^{-1}cm^{-2}$ $Å^{-1}$)
$mag_\lambda = -2.5 \log (f_\lambda) - 21.100 - zp(f_\lambda)$

f_v (10^{-20} erg $s^{-1}cm^{-2}$ Hz^{-1} = 1000 Jy)
$mag_v = -2.5 \log (f_v) - 48.585 - zp(f_v)$

Photometry: Fun with Units (1)

- **Why do we continue to use magnitudes?**
  - Historical reasons: astronomers have built up a vast literature of catalogues and measurements in the magnitude system.
  - The magnitude system is logarithmic, which turns the huge range in brightness ratios into a much smaller range in magnitude differences: the difference between the Sun and the faintest star visible to the naked eye is only 32 magnitudes.
  - Simplicity: Astronomers have figured out how to use magnitudes in some practical ways which turn out to be easier to compute than the corresponding brightness ratios.
- However, in general converting between different magnitude and photometric systems is difficult: conversion factors depend on the spectrum of each object.

Photometry: Fun with Units (2)

- Astronomers who study objects outside the optical wavelengths do not have any historical measurements to incorporate into their work.
- In those regimes, measurements are almost always quoted in "more rational" systems: units which are linear with intensity (rather than logarithmic) and which become larger for brighter objects:

  - erg $s^{-1}cm^{-2}$ $Å^{-1}$
  - erg $s^{-1}cm^{-2}$ Hz^{-1}

  - 1 Jansky [Jy] = 10^{-26} W m^{-2} Hz^{-1} = 10^{-23} erg $s^{-1}cm^{-2}$ Hz^{-1}
    F_ν [Jy]=$3.34\times10^{4}\ \lambda^2\ F_\lambda$ [erg s^{-1} cm^{-2} $Å^{-1}$]
    F_λ [erg s^{-1} cm^{-2} $Å^{-1}$]= $3.00\times10^{-5}\ \lambda^{-2}\ F_\nu$ [Jy]

Photometry: Fun with Units (3)

- Fluxes for a V = 0 star of spectral type A0 V at 5450 Å:
  - $f^0_\lambda = 3.63 \times 10^{-9}$ erg $s^{-1}cm^{-2}$ $Å^{-1}$, or
  - $\varphi^0_\lambda = f^0_\lambda / h\nu = 996$ photons $s^{-1}cm^{-2}$ $Å^{-1}$

  - Useful:
    - 1 Jy = $1.51 \times 10^3 / \lambda$ photons $s^{-1}cm^{-2}$ $Å^{-1}$

    - $\Delta\lambda/\lambda$ = 0.15 (U), 0.22 (B), 0.16 (V), 0.23 (R), 0.19 (I)

Night Sky Brightnesses

Lunar Age (days)	U	B	V	R	I
0	22.0	22.7	21.8	20.9	19.9
3	21.5	22.4	21.7	20.8	19.9
7	19.9	21.6	21.4	20.6	19.7
10	18.5	20.7	20.7	20.3	19.5
14	17.0	19.5	20.0	19.9	19.2

Signal from the sky background is present in every pixel of the aperture. Because each instrument generally has a different pixel scale, the sky brightness is usually tabulated for a site in units of mag/arcsecond^2.



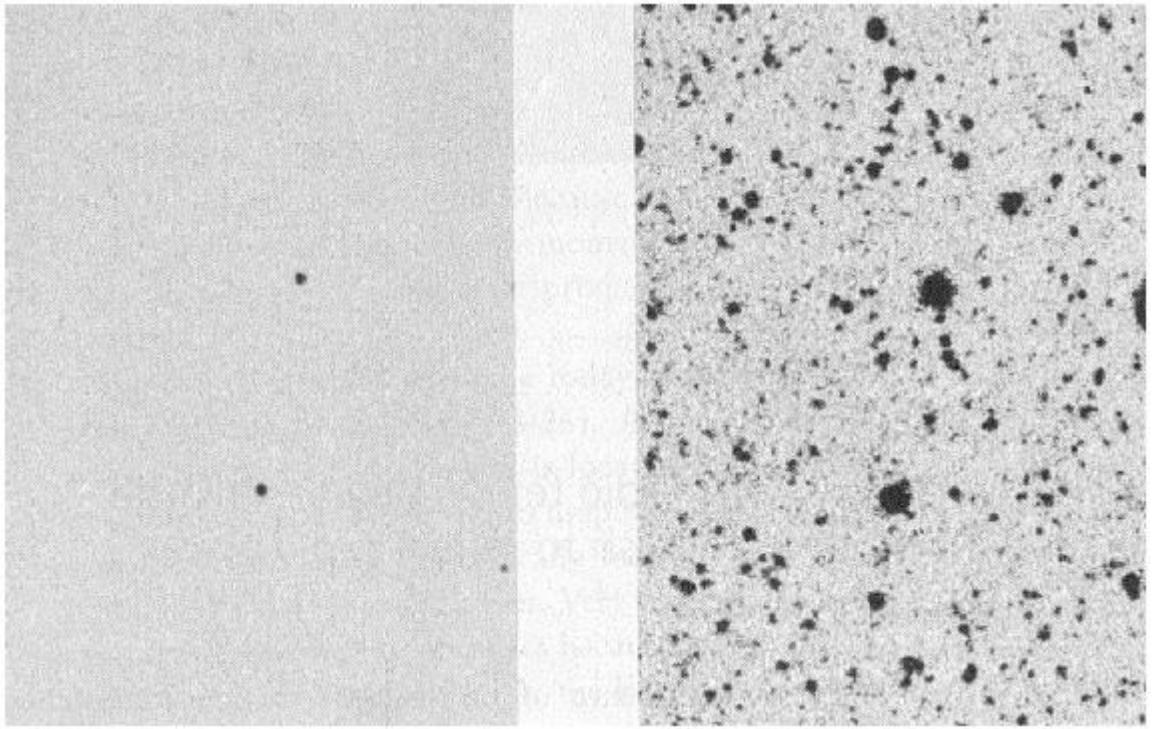
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Lecture 11

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Practical Photometry: S/N (2)



Physical limitations on the precision of photometric measurements (1)

From lecture 9 (slides 353+):

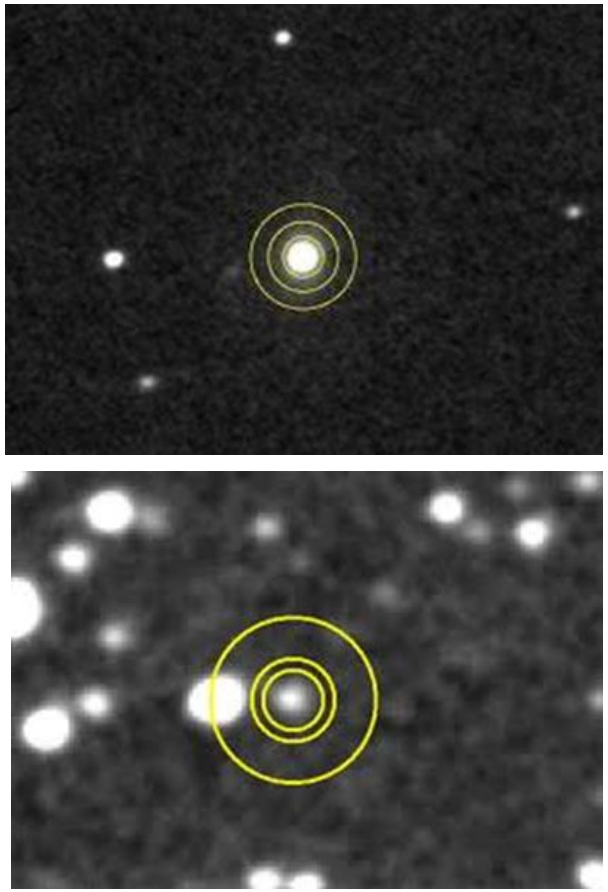
- To calculate the *Output* Signal-To-Noise Ratio of an observation we need to know the signal, and all sources of noise. These are:
  - Photon noise (shot noise) from the signal;
  - Photon noise from the sky background under the signal;
  - Photon noise from the sky background measurement to be subtracted off;
  - Readout noise from all sources;
  - Fixed pattern noise;
  - Bias noise;
  - Dark current noise.

Physical limitations on the precision of photometric measurements (2)

- Detective quantum efficiency= DQE $=\left[\frac{SNR_{out}}{SNR_{in}}\right]^2$
- We observe a star on a CCD detector, and process the data in the simplest way possible.

- $Reduced\ Frame = \frac{Object\ Frame - Bias\ Frame}{Flat\ Frame - Bias\ Frame}$

Physical limitations on the precision of photometric measurements (3)



The upper is good, the bottom is bad

- An area centred on the star is defined to be the object area, and is large enough to contain all of the photons from that star.
- An equal area some distance away, which is found to be free of stars, is defined as the sky background area, and the sky background is measured from that.

Aperture photometry (1)

- There are a number of parameters we need to take into account to calculate the signal which reaches the detector:
  - **t** – exposure time
  - **D** – diameter of the telescope
  - S_{sky} [photons / (cm^2 arcsec^2 second)] – brightness of the sky
  - **η** – quantum efficiency of a detector (QE)
  - φ_* [photons / (cm^2 second Å)] – the source flux to be measured
  - Star is observed in a circular aperture of area **β** square arcseconds which covers $\mathbf{n_{pix}}$ pixels
  - Sky background is determined from a circular aperture of the same size
  - Readout noise is σ_R electrons
  - We observe a star of magnitude V in the V filter



Signal calculation (1)

- We start from the number of photons incident upon the top of the atmosphere of the Earth from this star:

$$N_* = \phi_* \, \Delta\lambda \, A \text{ photons/second}$$

incident upon the top of the atmosphere in photometric (clear) conditions

$\Delta\lambda$ is the filter passband in Å

φ_* is the flux from a star in photons $s^{-1} cm^{-2} Å^{-1}$

A is the telescope collecting area in centimetres^2

Signal calculation (2)

- That's at the top of the atmosphere. There are a number of efficiency factors we need to multiply by to calculate the signal which reaches the detector:
  - Atmospheric transmission ε_{atm} (~0.88 for a star at the zenith, in the V filter).
  - Telescope reflection efficiency ε_{tel}
    (~0.92 per mirror = 0.846 for a Cassegrain telescope)
  - Filter transmission ε_{filt} (~0.85 for a broadband filter)
  - CCD Responsive Quantum Efficiency η (~0.75)
  - Cryostat entrance window efficiency ε_{win} (~0.95)

Signal calculation (3)

- There is also a geometric efficiency factor as part of the aperture of the telescope is blocked by the secondary mirror.

  For a ***D*** metre aperture telescope with a ***d*** metre secondary mirror:

$$\varepsilon_{\text{geom}} = \frac{\pi D^2 - \pi d^2}{\pi D^2} = \frac{D^2 - d^2}{D^2}$$

For example, if ***D***=2.0m and ***d***=0.6m, then $\varepsilon = 0.91$

Signal calculation (4)

- For a star the number of photons which is detected is given by:

$$N_{star} = \eta\, \varepsilon_{atm}\, \varepsilon_{tel}\, \varepsilon_{filt}\, \varepsilon_{win}\, \varepsilon_{geom}\, \phi_{*}\, \Delta\lambda\, A\, t$$

t is the exposure time in seconds,
$\Delta\lambda$ = 870 Å for the V band.

For example, for a star of magnitude V=23 on a 2 metre telescope with the efficiencies we have quoted:

$$N_{star} = 2.5\, t$$

Signal calculation (5)

- In the absence of sky background and readout noise it would be simple, we would integrate for 1000 seconds, detect 2500 photons, and have a signal to noise ratio of 50. But sky and readout noise are significant.
- Every square arcsecond of sky gives:

$$N_{sky} = \eta\, \varepsilon_{atm}\, \varepsilon_{tel}\, \varepsilon_{filt}\, \varepsilon_{win}\, \varepsilon_{geom}\, \phi_{sky}\, \Delta\lambda\, A\, t$$

φ_{sky} is the flux from the sky in photons $s^{-1} cm^{-2}$ $Å^{-1}$ $arcsec^{-1}$

≈10 t photons from the dark sky (V≈21.5)

Aperture photometry (2)

- Assume we have two apertures, one on the star and one on sky. Star aperture includes sky as well, and our estimate of the star intensity is the difference between the two.
- Signal in the sky aperture is:

$$n_{sky} = \beta N_{sky}$$

- Signal in the star aperture is:

$$n_{*+sky} = \beta N_{sky} + N_{star}$$

Noise on the measurements

- Noise on the measurements has two components, photon noise which is given by the square root of the number of photons, and readout noise, which is determined by the readout noise and by the number of pixels in the aperture. The noise components add in quadrature:

$$\sigma_{sky}^2 = n_{sky} + n_{pix}\,\sigma_R^2$$

$$\sigma_{*+sky}^2 = n_{*+sky} + n_{pix}\,\sigma_R^2$$

$$n_* \approx n_{*+sky} - n_{sky}$$

$$\sigma_*^2 = n_{*+sky} + n_{sky} + 2\,n_{pix}\,\sigma_R^2$$

$$\sigma_*^2 = 2\,\beta\,N_{sky} + N_{star} + 2\,n_{pix}\,\sigma_R^2$$

Signal to Noise ratio

$$S/N = n_* / \sigma_* = N_{star} / \sigma_*$$

$$\frac{S}{N} = \frac{N_{star}}{\sqrt{2\,\beta N_{sky} + N_{star} + 2\,n_{pix}\sigma_R{}^2}}$$

$N_{star} = \eta\, \varepsilon_{atm}\, \varepsilon_{tel}\, \varepsilon_{filt}\, \varepsilon_{win}\, \varepsilon_{geom}\, \phi_*\, \Delta\lambda\, A\, t$
$N_{sky} = \eta\, \varepsilon_{atm}\, \varepsilon_{tel}\, \varepsilon_{filt}\, \varepsilon_{win}\, \varepsilon_{geom}\, \phi_{sky}\, \Delta\lambda\, A\, t$

If exposure time is short then readout noise ($\sigma_R \sim 10$) will **dominate**, especially when seeing is bad.

Note: seeing comes in with n_{pix} term

What is ignored in this S/N eqn?

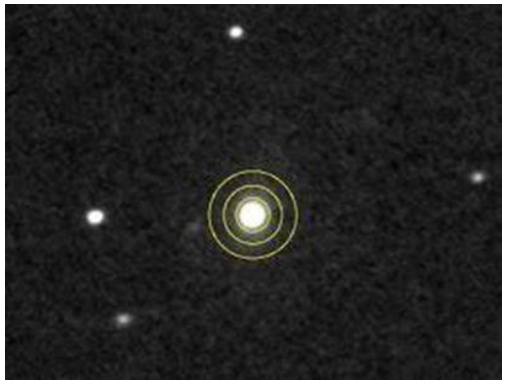
- Bias level/structure correction
- Flat-fielding errors
- Charge Transfer Efficiency (CTE)
  0.99999/pixel transfer
- Non-linearity when approaching full well
- Scale changes in focal plane
- A zillion other potential problems

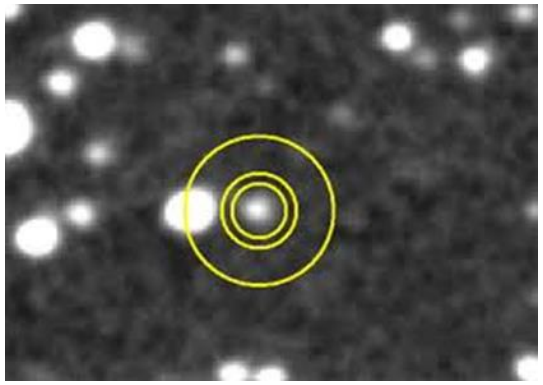
Improving the Signal to Noise (1)

- Larger Sky Aperture – Increasing the sky aperture and scaling it to the size of the object aperture, or using several sky apertures and averaging them, reduces the noise to:

$$\sigma_*^2 = \zeta \beta N_{sky} + N_{star} + \zeta n_{pix*} \sigma_R^2$$

where $\zeta = (1 + n_{pix*} / n_{pix\_sky})$, and n_{pix*} and n_{pix_sky} are the number of pixels in the star and sky apertures respectively. In practice, the sky aperture is often an annulus around the star aperture. Must be careful that stars do not get in the sky aperture!





Improving the Signal to Noise (2)

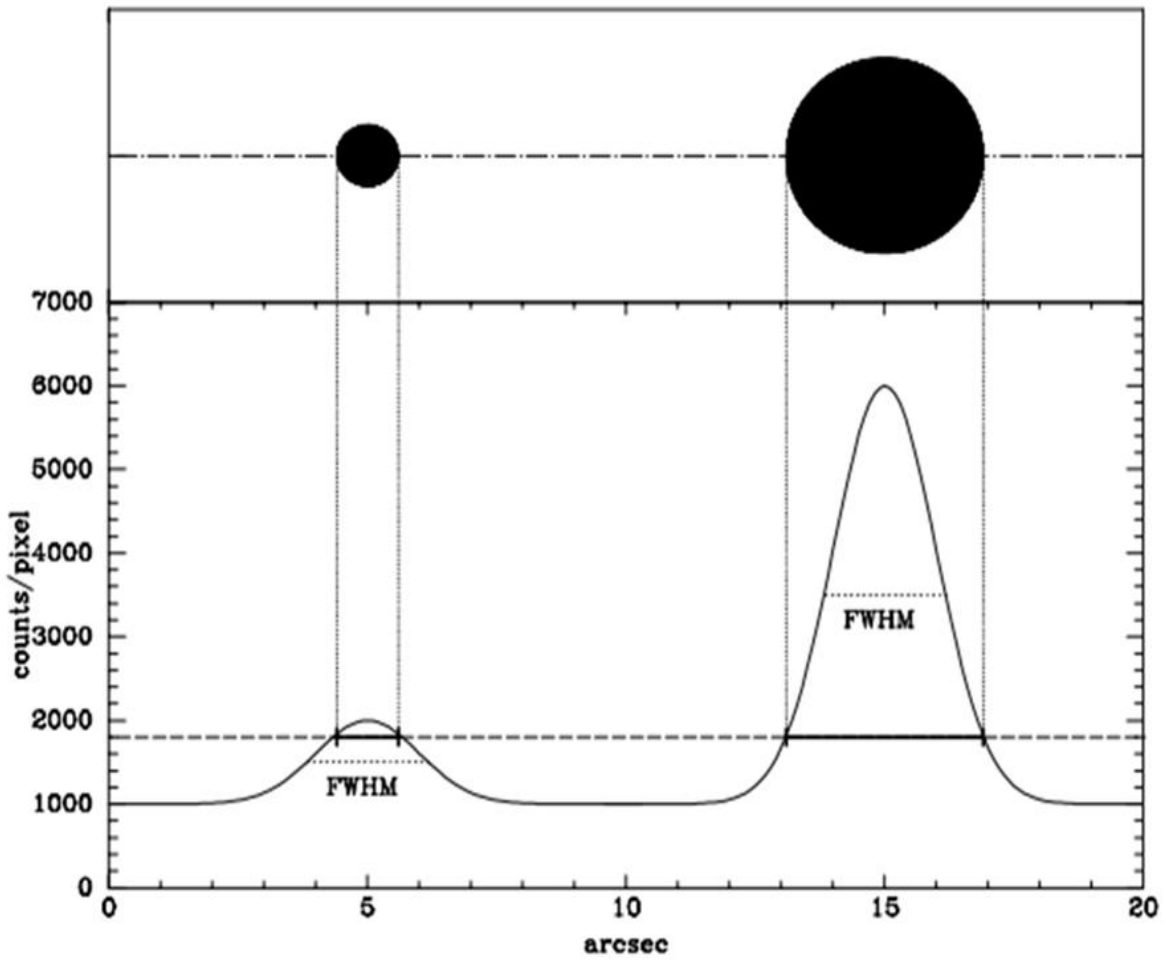
- Smaller object aperture – reducing the object aperture reduces both sky noise and readout noise. However, you lose signal. The problem is if you are comparing the signal in different images, and fluctuations in image size (seeing) cause the amount of signal you lose to vary, then this introduces systematic errors in the brightness measured (photometry).

Solution – Aperture Correction

Aperture Correction (1)

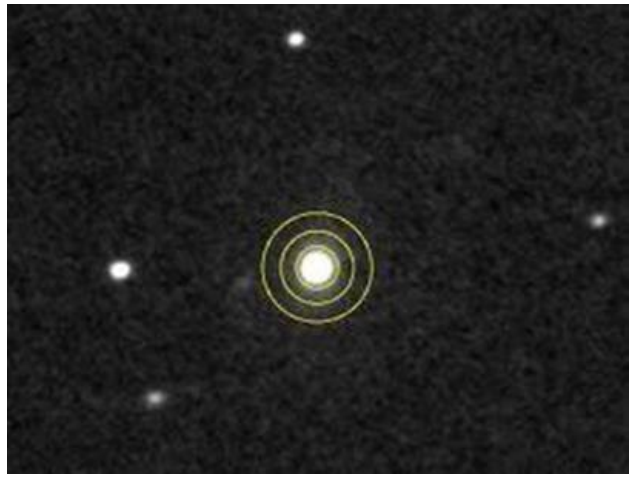
- The **point spread function** (PSF) is the shape of the CCD image of a point (unresolved) source of light.
- Since the PSF is the shape of a point of light on the CCD, and since all stars are points, then all stars have exactly the same shape and size on the CCD, if aberrations are not significant.
- The PSF does not have an edge. The intensity of the star fades smoothly to zero with increasing radius, but there is no place that we could call an "edge".

Aperture Correction (2)



Brighter stars may look bigger, but that is caused by the following effect:

the shape of the faint and bright star are exactly the same, we are simply looking at a larger diameter at a **given intensity** for a bright star than for a faint star.



Aperture Correction (3)

- If we want to measure all the light from a star, how far out in radius do we have to go?
  - One logical answer might be: as big as possible, to get "all" the light from the star. This is **not** a good answer.
  - Reducing the object aperture **reduces** both sky noise and readout noise.

- **But**, a small aperture will only encompass a **fraction** of the total light from the star! However, if the seeing were constant, any aperture would measure the same fraction of light for any star, and when comparing one star with another the effect would cancel out.

Aperture Correction (4)

- The problem is that seeing is not constant. A small aperture might measure 0.5 of the total light from a star on one CCD image, then, if the seeing worsens, the same size aperture might measure only 0.4 of the light from the star on the next CCD image.
- Seeing affects mostly the inner Gaussian core of the image. Using an aperture 4 to 10 times the diameter of the typical FWHM will get most of the light. In this size aperture, reasonable variations in the seeing will not result in measurable variations in measured counts.

Aperture Correction (5)

- However, for faint objects, an aperture 4 times the FWHM will contain a lot of sky signal. This will result in a low S/N ratio.
- Aperture Correction: If we measure the bright object in a small aperture (say radius = 1 FWHM) and also in a bigger aperture which gets "all" the light (say 4 FWHM) we can easily find the ratio of light in the small to large aperture (which we express as a magnitude difference).

Aperture Correction (6)

- The aperture correction is defined as:

$$\Delta = m_{inst}(4\ \text{FWHM}) - m_{inst}(1\ \text{FWHM})$$

(Δ is always a negative number)

- How do we use the aperture correction?

$$\text{total} = m_{inst}(1\ \text{FWHM}) + \Delta$$

- "Total" is our estimate of the total instrumental magnitude in the faint star, $m_{inst}(1\ \text{FWHM})$ is the measured magnitude in the small aperture for the faint star, and Δ is the aperture correction derived from a bright star in the same frame.

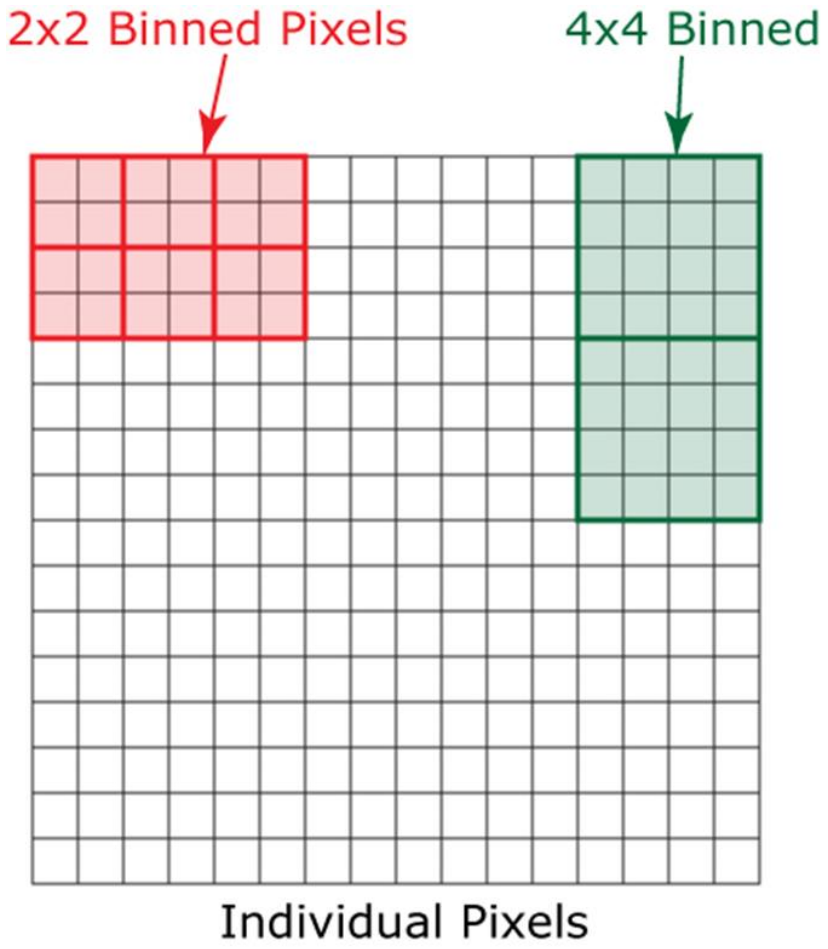
Aperture Correction (7)

- There must be an optimum aperture size that gives the maximum S/N.
- The optimum size of the small aperture has been studied by several authors.
- The optimum aperture seems to be achieved when the measurement aperture has a diameter about 1.4 × FWHM of the PSF. At this aperture, the aperture correction is about −0.3 mag.
- However, the S/N does not appear to be too sensitive to the exact small aperture size.

Improving the Signal to Noise (3)

- On-chip binning – Most CCDs have the option of *binning*: combining a set of adjacent pixels into a single pixel produced as output. For example, a square of 4 pixels on the CCD chip might be reported as one pixel containing their combined value. But you only have to read the output capacitor out once and you only get one lot of readout noise.

On-chip binning



- This way you reduce readout noise at the expense of resolution. Resolution should always be smaller than the characteristic size of the star images.

Profile Fitting (1)

- Profile fitting is used most commonly in crowded fields, where it is difficult or impossible to define a sky aperture free of stars (or galaxies).
- It does however offer an advantage in precision even in sparse fields, because it weights the data more correctly.

Profile Fitting (2)

- Basic assumption is that the intensity profile (which is in principle a 2 dimensional function) is the same for all stars in a particular CCD image.
- Intensity profile is determined by seeing or by diffraction, or occasionally by aberrations.
- If it is determined by aberrations you need to be very careful, because the assumption that the profile is the same at all positions on the CCD may **not** be correct.

Profile Fitting (3)

- From a set of isolated, comparatively bright (but not saturated) stars in the frame, determine the image profile, this is called the **Point Spread Function (PSF).**
- For ground based data an empirical approximation to the PSF is the Moffat function:

$$f(r) = C_i \ (1 + r^2/R_0^2)^{-\beta} + B_i \quad (r < r_{max})$$

$$f(r) = B_i \ (r > r_{max})$$

R_0 is the characteristic radius of the star image,
r is the distance from the centre of the image,
β describes the overall shape of the PSF,
B_i is the background in the region of star i, and
C_i is the relative brightness of star i.

Profile Fitting (4)

$$FWHM = 2\ R_0 \sqrt{2^{1/\beta} - 1}$$

- Fit this function for each of the stars in the image to the data, using a least squares or similar technique.
- For each star determine B_i and C_i.
  R_0 and β are constant within an image.

Profile Fitting (5)

- Then we have a set of scaling factors, which can be converted to a relative magnitude.
- We need aperture photometry of one star, either from this CCD frame or from another, this can be a bright isolated star with high S/N, this gives the magnitudes of all of the stars in the frame.
- The fit gives the correct weighting, rather than adding in lots of pixels with very little signal, S/N from profile fitting is usually at least a factor of 2 higher than from aperture photometry.
- Profile fitting can cope with fields in which stars are close or their images even overlap.

Profile Fitting (6)

- For ground-based data the PSF is determined by the seeing, and must be redetermined for each CCD image.
- For space based (e.g. Hubble Space Telescope) data the PSF is fixed, and is often available as part of the standard calibration data produced with the observations.
  It still depends upon the passband (filter).



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 12

Vitaly Neustroev

Procedures for photometry

Extracting Photometric Data

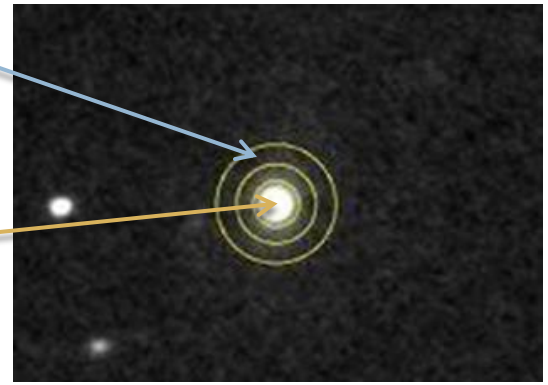
Differential (Relative) Photometry

Instrumental Magnitudes

Extinction

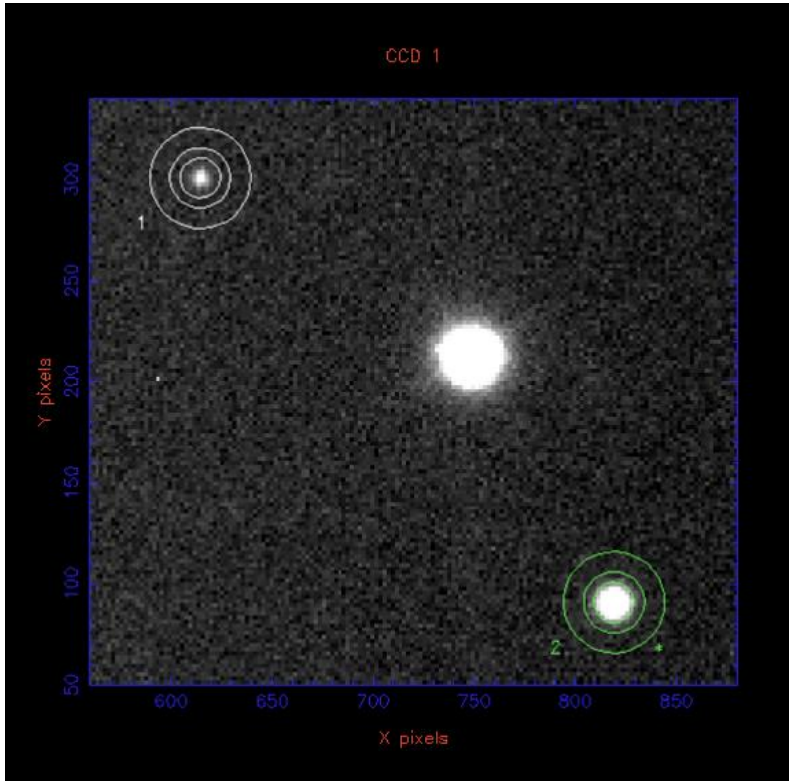
Extracting Photometric Data

- Determining the number of counts from a source in an image is usually a three-step process (we follow a technique known as aperture photometry):
  1. To measure the centre of the source, which we shall assume is a star.
  2. To estimate the sky background at the position of the star.
  3. To calculate the total amount of light received from the star.
- The sky annulus is unlikely to contain counts from the sky alone. There will also be contributions from cosmic rays, hot pixels, faint stars, and the wings of the PSF of the central star. All of these will add a positive skew to the histogram of pixel values in the annulus. The mean of these pixel values will then not be an accurate representation of the sky background.
  Instead, the sky level is usually determined using a more robust estimator, such as **the median**.
- The total signal from the star can then be calculated by summing the counts from each pixel that falls inside the aperture (usually a **circle** or **ellipse**), and then subtracting the determined sky background from each pixel.



Differential (Relative) Photometry

- Even if the intrinsic brightness of the star is constant, the number of counts we detect might change, due to transparency variations or seeing variations. To obtain an accurate light-curve we need to correct for these effects, using a **comparison star**.
- This is a second star which is known (or assumed) to have a constant flux. We assume that the comparison star is affected in the same way as our target star by seeing and transparency variations (this is a very good assumption).
- Therefore, if transparency or seeing variations cause the counts from the target star, N_t, to halve, they will also cause the counts from the comparison star, N_c, to halve.
  The ratio N_t/N_c is therefore corrected for transparency and seeing variations. Correcting aperture photometry this way is known as differential or relative photometry.



Courtesy of Vik Dhillon

Calibrated Magnitudes (1)

- Once the sky-subtracted signal from an object, measured in **counts**, is extracted from an image, it is useful to convert it to a calibrated **magnitude** in a photometric system:
  - Calculate the instrumental magnitude, from the counts per second.
  - Determine the extinction coefficient, and correct the instrumental magnitude to the above-atmosphere value.
  - Repeat the above steps for a standard star and use the resulting above-atmosphere instrumental magnitude of the standard star to calculate the zero point.
  - Use the zero point to transform the above-atmosphere instrumental magnitude of the target star to the required photometric system.

Procedures for photometry

- If we have standard stars in the CCD field that we are observing, then its fairly easy to calibrate, as we can just use the N_t values as our measure of intensity.
- If not then we need to observe standard stars in separate CCD frames, and as the PSF will vary between different frames, we need to find a true measure of the brightness of the stars.

$$N_t = C_i \int_0^{rmax} 2\pi r \left(1 + r^2/R_0^2 \right)^{-\beta} dr$$

r_{max} is chosen so that we get all of the light.

(The Moffat function, see the previous lecture)

Instrumental Magnitudes (1)

- We have the sky-subtracted signal from our target object, in counts, N_t. We convert this to an instrumental magnitude, using the formula:

$$m_{inst} = -2.5 \log_{10} (N_t / t_{exp})$$

where t_{exp} is the exposure time of the image in seconds.

- The instrumental magnitude depends on the characteristics of the telescope, instrument, filter and detector used to obtain the data.
- We need to compare the instrumental magnitudes of stars of known magnitude with their true magnitudes, to calculate the offset, and thus to calculate the true magnitudes of all of the stars in the frame.
- The relationship between instrumental magnitudes and calibrated magnitudes can be understood as follows. The counts per second N_t / t_{exp} is proportional to the flux, F_λ. Hence

$$m_{inst} = -2.5 \log_{10} (c F_\lambda) = -2.5 \log_{10} (F_\lambda) + c'$$

Instrumental Magnitudes (2)

- Therefore, instrumental magnitudes are offset from calibrated magnitudes by a constant:

$$m_{calib} = m_{inst} + m_{zp}$$

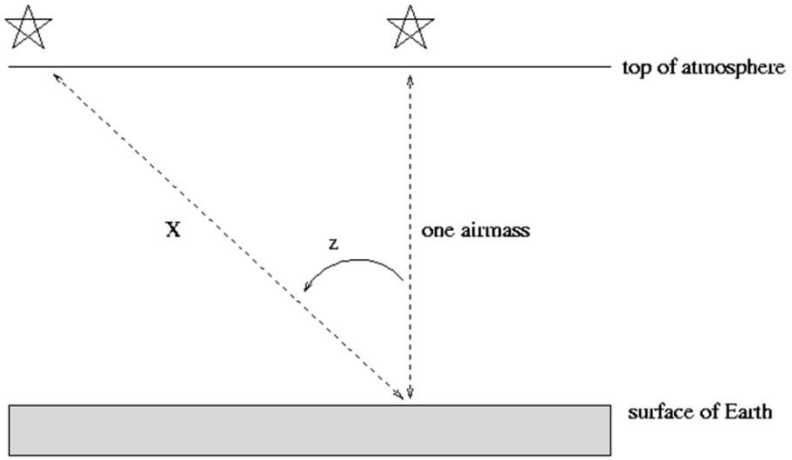
where the constant, m_{zp}, is known as the **zero-point**.

- The zero-point depends upon the telescope and filter used. We can understand this because if we used a larger telescope to observe a star, the instrumental magnitude would change, but the calibrated magnitude must not!
- An object with a calibrated magnitude equal to the zero-point gives one count-per-second at the telescope.
- For example, suppose the zero-point of a telescope/filter combination is m_{zp}=19.0. If we observe a star with m_{calib}=19.0, then it follows that, for this star m_{inst}=0. By definition then, this star gives one count-per-second.

$$m_{inst} = -2.5 \log_{10} (N_t / t_{exp})$$

Atmospheric absorption

- The next step is to convert the instrumental magnitude, which is measured on the surface of the Earth, to the instrumental magnitude that would be observed above the atmosphere.
- Atmospheric absorption is proportional to the airmass, which is proportional to the secant of the angular distance from the zenith. Strictly this assumes a plane parallel atmosphere, but this is a good approximation for z < 70°.



$$X = \sec Z = [\sin \varphi \sin \delta + \cos \varphi \cos \delta \cos h]^{-1}$$

φ is the latitude of the observatory, h is the hour angle of the source, and δ is the declination of the source.

Atmospheric extinction

- The effect of atmospheric extinction on photometry is usually expressed as:

$$m_{obs} = m_{true} + k(\lambda) \sec Z = m_{true} + k(\lambda)\, X$$

Here, m_{true} is the magnitude of the source outside the Earth's atmosphere, m_{obs} is the magnitude observed,

$k(\lambda)$ is the "extinction coefficient" [magnitudes per unit airmass].

- The dominant source of extinction in the atmosphere is Rayleigh scattering by air molecules. This mechanism is proportional to λ^{-4}, which means that extinction is much higher in the blue than in the red.
- The extinction coefficients $k(\lambda)$ have been carefully measured and tabulated for a number of observatories.

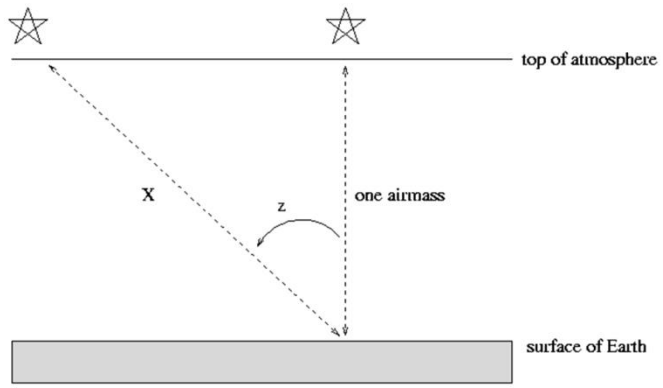
The extinction coefficients on a typical (undusty) night on La Palma

Filter	λc (Å)	k
U	3660	0.55
B	4380	0.25
V	5450	0.15
R	6410	0.09
I	7980	0.06

Extinction coefficient (1)

- However, the extinction can vary from night to night depending on the conditions in the atmosphere, e.g. dust blown over from the Sahara can increase the extinction on La Palma during the summer by up to 1 magnitude.
- How do we find k in practice?
  if we plot instrumental magnitudes vs airmass for a particular star, k is just the slope of the line that passes through the observed points:

$$k = \frac{\Delta m_{inst}}{\Delta X}$$



Extinction coefficient (2)

- Even two measurements of the instrumental magnitude of a star at two different zenith distances is enough to estimate *k* (although less accurate): subtracting $m_{z1} = m_0 + k \sec Z_1$ from $m_{z2} = m_0 + k \sec Z_2$ eliminates m_0, allowing k to be derived.
- For more accurate photometry: observe a set of standard stars (of known magnitude) at different airmass.



- Note that no explicit extinction correction is required when performing differential photometry: the target and comparison stars are always observed at the same airmass and hence suffer the same extinction. Hence, the variation due to extinction present in the comparison star is removed from the target star.

Colour term (1)

- For **very accurate** photometry, the wide bandpass of broad-band filters has to be taken into account when correcting for extinction. **Even for differential photometry!**
- There is a colour term, caused by the variation in spectral profile of the stars and the filter response over the passband.
- Because extinction is so strongly colour dependent, a blue object actually loses more light to the atmosphere than a red one. The solution is to introduce additional colour-dependent secondary extinction coefficients, which modify the above extinction correction equation to:

$$m_{obs} = m_{true} + k_1(\lambda) \sec Z + k_2(\lambda)\,(B\text{-}V) + k_3(\lambda)\,(B\text{-}V)\sec Z$$

(B-V) is the colour of a star. Other colours can be used, e.g. (V-R).

Colour term (2)

$$m_{obs} = m_{true} + k_1(\lambda)\ \sec Z + k_2(\lambda)\ (B\text{-}V) + k_3(\lambda)\ (B\text{-}V)\sec Z$$

- Solve for k_1, k_2 , k_3 from stars of known magnitude.
- Usually k_2 is negligible (usually of order a hundredth of a magnitude), often k_3 is too.
- In this case we can now simply convert the values of m_{inst} to m_{true} using the value of k_2 that we solve for, and the value of z for each observation.
- However, if k_2 and/or k_3 is not zero, we need to know **(B-V)** for the star to calculate the true magnitude, but we do not. In this case we must observe a set of standard stars (of known magnitudes) at different Z and in two passbands, for instance B and V, and use:

$$V_{obs} = V_{true} + k_1(\lambda)\ \sec Z + k_2(\lambda)\ (B\text{-}V) + k_3(\lambda)\ (B\text{-}V)\sec Z$$
$$B_{obs} = B_{true} + k_4(\lambda)\ \sec Z + k_5(\lambda)\ (B\text{-}V) + k_6(\lambda)\ (B\text{-}V)\sec Z$$

Calibrated Magnitudes (2)

- Now we can find the above-atmosphere instrumental magnitude of any object. If the above-atmosphere instrumental magnitude of our standard star is $m_{std,0,i}$, then:

$$m_{zp} = m_{std} - m_{std,0,i}$$

- The calibrated magnitude of our target star, m_{calib}, can then be found using:

$$m_{calib} = m_{zp} + m_{0,i}$$

where $m_{0,i}$ is the above-atmosphere instrumental magnitude of our target star.

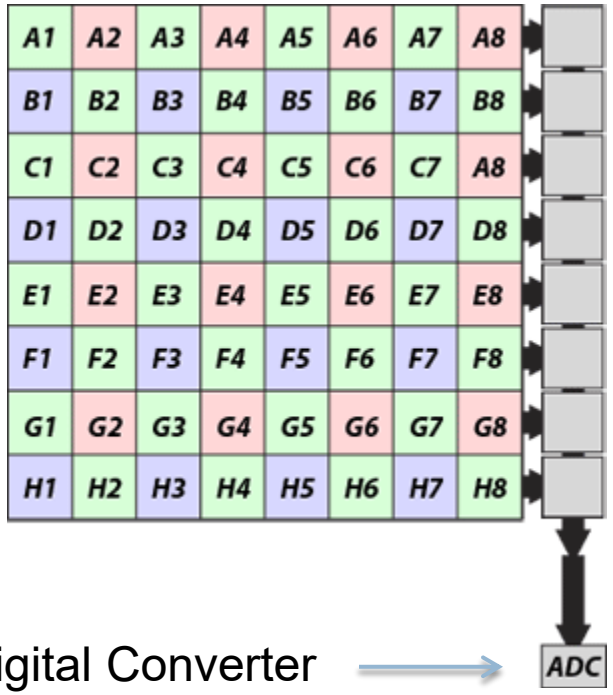
- Each filter in a photometric system will have a different zero-point. Once the zero-point has been measured for a particular telescope, instrument, filter and detector combination, it **should** remain unchanged, although dirt and the degradation of the coatings on the optics will cause minor changes to the zero point on long timescales. To determine the zero points for the UBVRI system, the photometric standards measured by Landolt can be used.

CCD Gain

CCD Gain (1)

- What is relationship between electrons in a CCD and pixel values?

The readout register is shifted to the right by one pixel, and the pixel at the bottom right is shifted into a readout capacitor. What's next?



CCD Gain (2)

- The steps involved in reading the value of a pixel are something like this:
  1. Electrons transferred to "amplifier"; really a capacitor. Units are **coulombs**.
  2. The voltage induced by this charge is measured. Units are **volts**.
  3. An Analog-To-Digital (A/D) unit converts the voltage into some other voltage, which may have only one of several discrete levels. Units are still **volts**.
  4. The voltage is converted into a number which is passed from the hardware to the computer software as the pixel's value. Units are **counts**, also called "Analog-to-Digital Units" (**ADUs**).

CCD Gain (3)

- In both steps 3 and 4, the result may be multiplied by an arbitrary factor without altering the relative pixel values. Some software packages allow this scaling factor to be adjusted dynamically, whereas others use a fixed value.
- The end result is that there is some factor which relates the initial number of electrons in a pixel to the final number of counts reported by camera software. The ratio of these two numbers is the **gain** of the camera:

$$gain = \frac{Number\ of electrons\ per\ pixel}{Number\ of\ counts\ per\ pixel}$$

CCD Gain factor (1)

- How should one choose the gain factor? There are several criteria.

1. **Full-well depth vs. largest pixel value:**
   Each CCD is designed to hold only so many electrons within a pixel before they start to leak outwards to other pixels. This maximum size of a charge packet on the chip is called the **full well depth.**
   There is also a "maximum possible number" in the Analog-to-Digital converter.
   Most CCDs use 14-bit, 15-bit, 16-bit, sometimes 32 A/D units:
   the corresponding maximum pixel values are
   $2^{14} = 16384$, $2^{15} = 32768$, $2^{16} = 65536$, $2^{32} = 4294967296$.
   It is logical to arrange the gain so that very roughly, the number of electrons in the full-well depth corresponds to the maximum pixel value.

CCD Gain factor (2)

- How should one choose the gain factor? There are several criteria.

2. **Readout noise vs. smallest pixel value:**
   What are the SMALLEST values that make sense? A typical **readout noise** is 3 – 10 electrons. Therefore, if two pixels have values which differ by only 2 electrons, it's not easy to tell the difference between them. The smallest difference one can represent in an integer image is 1 count. To some extent, it makes sense to arrange the gain so that 1 count corresponds to some moderate fraction of the readout noise:

$$gain = \frac{Number\ of electrons\ per\ pixel}{Number\ of\ counts\ per\ pixel} \sim 4\ (\text{typical readout noise})$$

   Any finer measurement of the pixel values would yield differences which would be essentially random.

Why should we care about the Gain?

- For everything what we discussed previously and where we used the **Counts** we **must** use the number of **electrons** (photons). They can be obtained from the counts using the gain factor (photon per data unit).

$$Number\ of\ electrons = gain * Number\ of\ counts$$

Example: CCD Gain (ALFOSC)

- ALFOSC - the Alhambra Faint Object Spectrograph and Camera – is an instrument on the NOT telescope built to allow the acquisition of both images and spectra.

- The detector (CCD14) is an **CCD231-42-g-F61** back illuminated, deep depletion CCD with 2048 x 2064 pixels.
- Readout Noise is **~4.3 e^-/pix**
  the gain is **0.16 e^-/ADU**.
- Dark current is **1.3e^-/pix/hour**
- Dynamical range is **32 bit.**
- Full-well capacity – **135000 electrons (700 kADUs).**
  Good (linear better than ±1%) up to ~113500e^- (600 kADUs)

Example: CCD Gain (BFOSC)

- BFOSC - Bologna Faint Object Spectrograph & Camera – is an instrument built to allow, with a simple configuration change, the acquisition of both images and spectra.

- The detector is an **EEV LN/1300-EB/1** CCD with 1300 x 1340 pixels, AR Visar coated, back illuminated.
- The detector Readout Noise is **3.06 e^-/pix** and the gain is **2.22 e^-/ADU**.
- Dynamical range is **16 bit**.
- Full-well capacity – **117000 electrons.**



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 13

Vitaly Neustroev

Spectroscopy

Spectral analysis is the source of most of our astrophysical knowledge.

Outline

1. General introduction to spectroscopy
2. Practical spectroscopy
3. Spectral reductions (calibration)

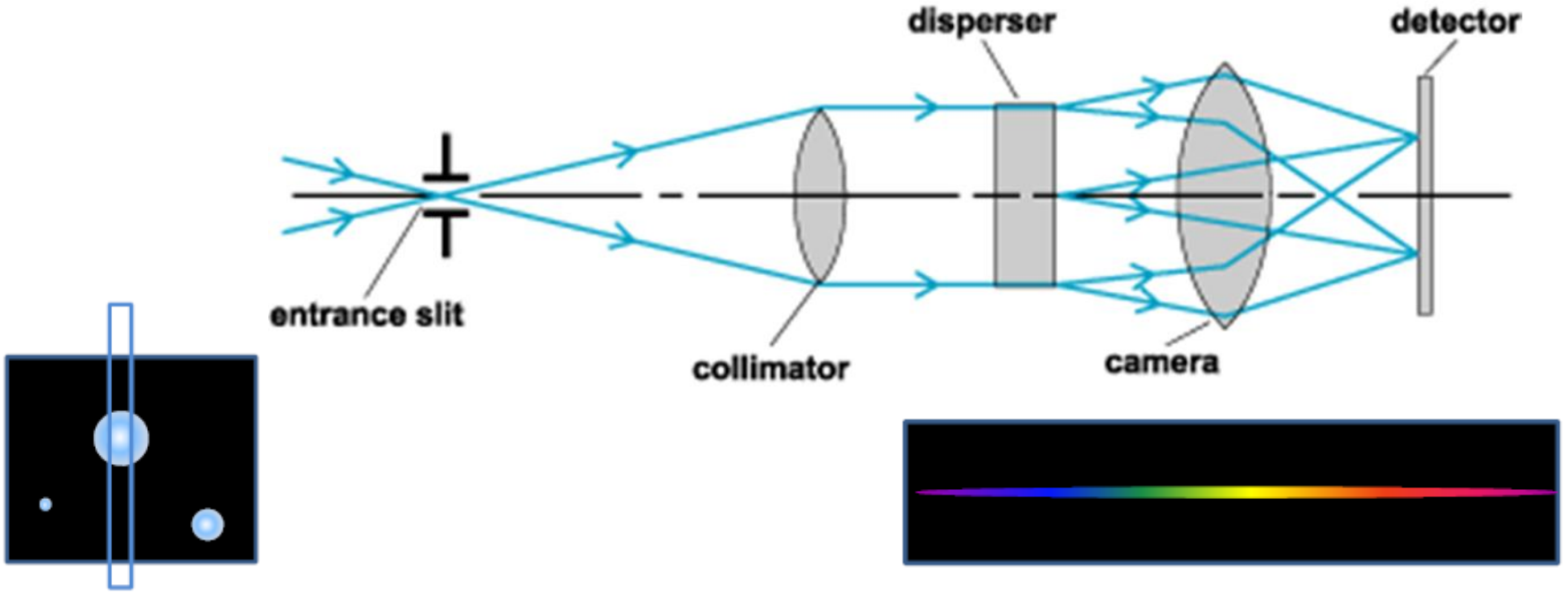
Techniques of Spectroscopy

Spectroscopy (spectrophotometry) is the most informative technique of light analysis, that measures how much light an object produces at various wavelengths of light.

- At X-ray and gamma ray wavelengths detectors have intrinsic energy resolution.
- At lower energies we must use different techniques to separate radiation of different wavelength/energy/frequency spatially.
- These techniques are largely those of interference, so considering radiation as waves.

Spectrometers

- All spectrometers have essentially the same basic design, but many different implementations are possible depending on the constraints and choice of spectral disperser.

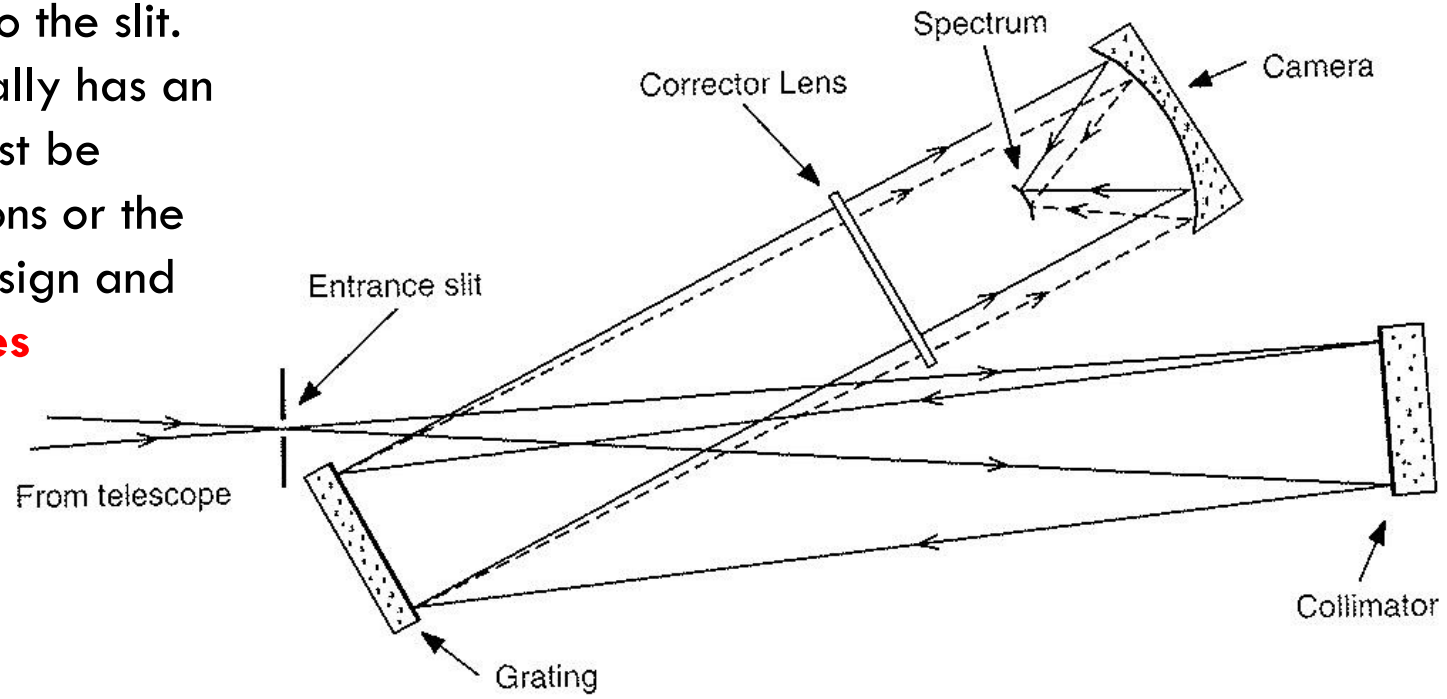


Spectrometers: main elements

- **Entrance aperture:**
  The image of a target is focused onto the slit. The slit is in the focal plane, and usually has an adjustable width w. The slit width must be matched to either the seeing conditions or the diffraction disc depending on the design and application. A narrower slit **improves** resolution ~1/w ×.

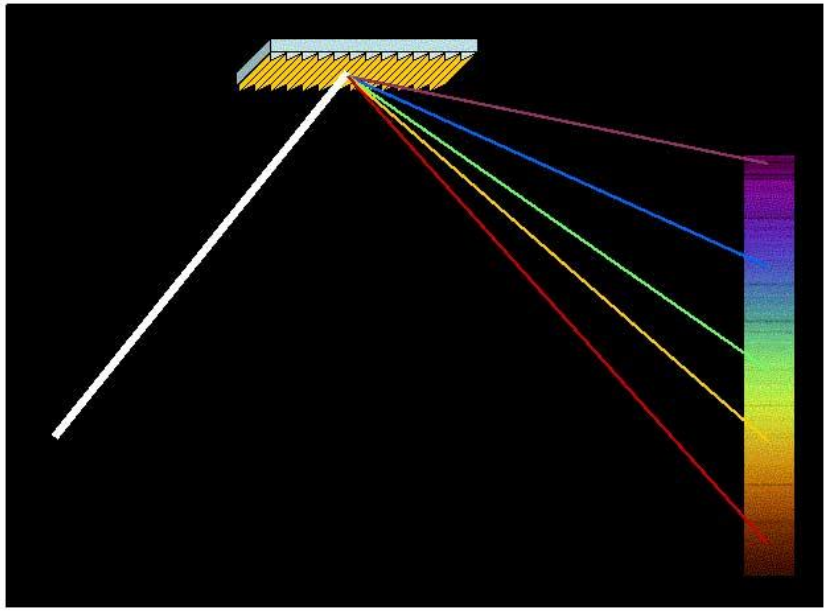
- **Collimator:** makes the rays parallel

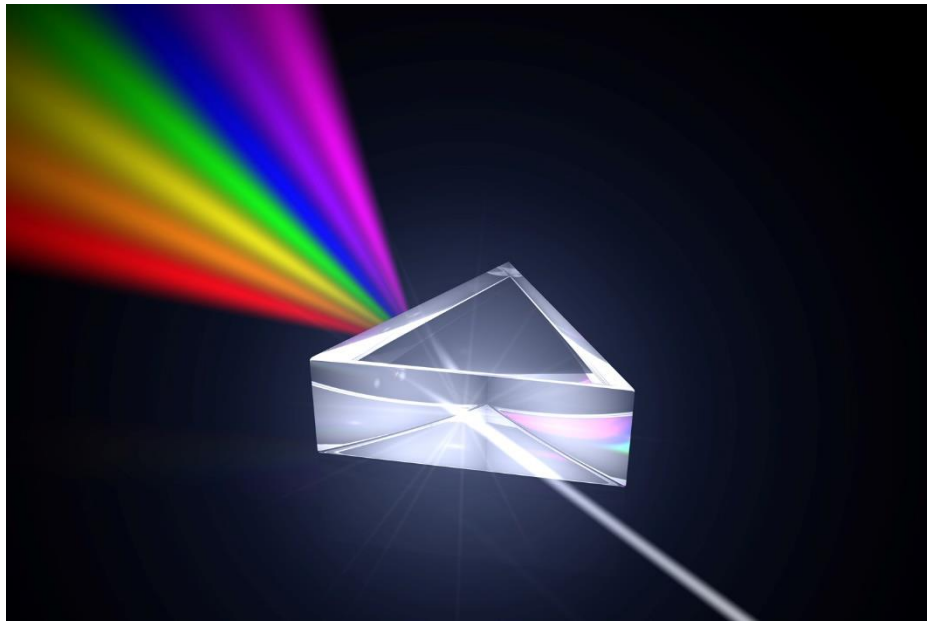
- **Disperser** disperses the light into colours: grating or prism, usually on rotating stage so can adjust central wavelength.



- **Camera:** to re-focus parallel output beam from disperser onto focal plane of detector (CCD)

Disperser: Grating vs Prism

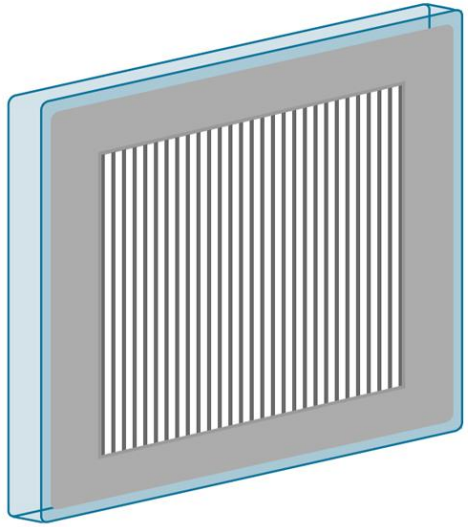




Resolution of a prism is low compared to what is possible with a grating, therefore **grating is usually the primary dispersive element in a modern spectrograph.**

Diffraction Grating (1)

- A diffraction grating is a set of multiple, identical slits (transmitting or reflecting) separated by a distance comparable to the wavelength of light.
- Each slit can be considered as radiating secondary waves (Huygens' secondary wavelets).
- The amplitude at any point on the image side of the slit can be calculated by summing the amplitude contributed by each set of secondary wavelets.



Diffraction Grating (2)

- The theory of Fraunhofer diffraction from a plane grating predicts that the diffracted light is distributed as:

$$I(\theta) = I_0 f_1 f_2,$$

where I is the output intensity leaving the grating in direction ϑ with respect to the normal,
I_0 is the input intensity at the grating,
f_1 is the diffraction pattern for a single grating slit, and
f_2 is the pattern for a set of N identical apertures.

- The two patterns are given by:

$$f_1 = \frac{\sin^2(\pi\alpha)}{(\pi\alpha)^2}, \qquad \alpha = \frac{a \sin\theta}{\lambda}$$

$$f_2 = \frac{\sin^2(N\pi\delta)}{\sin^2(\pi\delta)}, \qquad \delta = \frac{d \sin\theta}{\lambda}$$

where a is the linear width of the apertures (assumed rectangular) and d is the linear separation between them (the groove spacing).
Here, we assume normal incidence of the incoming light.

Diffraction Grating (3)

- Consider monochromatic light.
- Principal maxima ("orders") in the multi-slit pattern occur for $\delta = n$, where n is any integer, i.e. $\sin(\pi\delta)=0$.

$$f_2 = \frac{\sin^2(N\pi\delta)}{\sin^2(\pi\delta)}$$

- Indeed,

L'Hôpital's rule $\lim_{x\to c}\frac{f(x)}{g(x)} = \lim_{x\to c}\frac{f'(x)}{g'(x)}$

$$\lim_{\beta\to\pm n\pi}\frac{\sin(N\beta)}{\sin(\beta)} = \lim_{\beta\to\pm n\pi}\frac{\frac{d}{d\beta}(\sin(N\beta))}{\frac{d}{d\beta}(\sin(\beta))} = \lim_{\beta\to\pm n\pi}\frac{N\cos(N\beta)}{\cos\beta} = N\lim_{\beta\to\pm n\pi}\frac{\cos(N\beta)}{\cos\beta} = N$$

- Thus, the condition for principal maxima $\sin(\pi\delta)=0$

Diffraction Grating (4)

- Consider monochromatic light.
- Principal maxima ("orders") in the multi-slit pattern occur for $\delta = n$, where n is any integer.

  This implies the path difference Δ between adjacent slits will be n wavelengths:

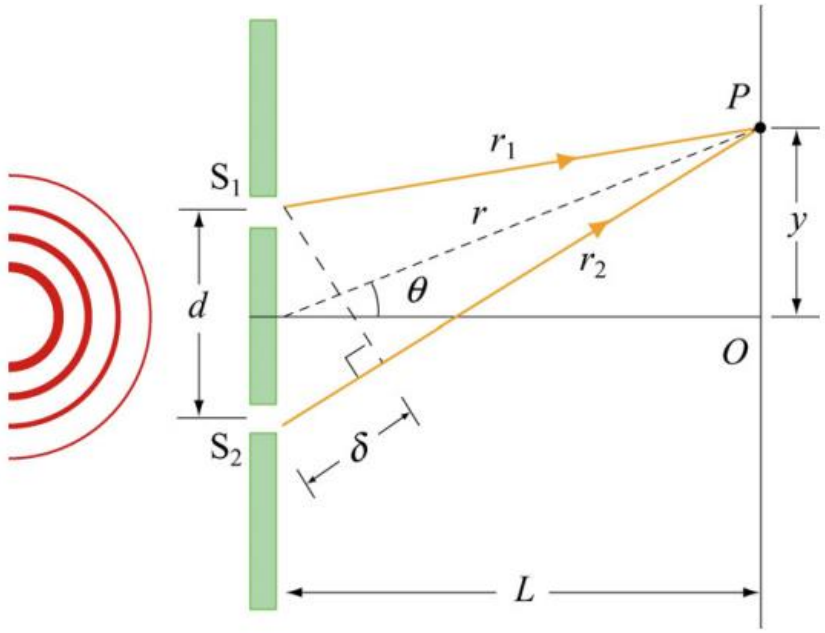
Using the law of cosines,

$$r_1^2 = r^2 + \left(\frac{d}{2}\right)^2 - dr\cos\left(\frac{\pi}{2} - \theta\right) = r^2 + \left(\frac{d}{2}\right)^2 - dr\sin\theta$$

$$r_2^2 = r^2 + \left(\frac{d}{2}\right)^2 - dr\cos\left(\frac{\pi}{2} + \theta\right) = r^2 + \left(\frac{d}{2}\right)^2 + dr\sin\theta$$

Subtracting

$$r_2^2 - r_1^2 = (r_2 + r_1)(r_2 - r_1) = 2dr\sin\theta$$



If L≫d, i.e. the distance to the screen is much greater than the distance between the slits):

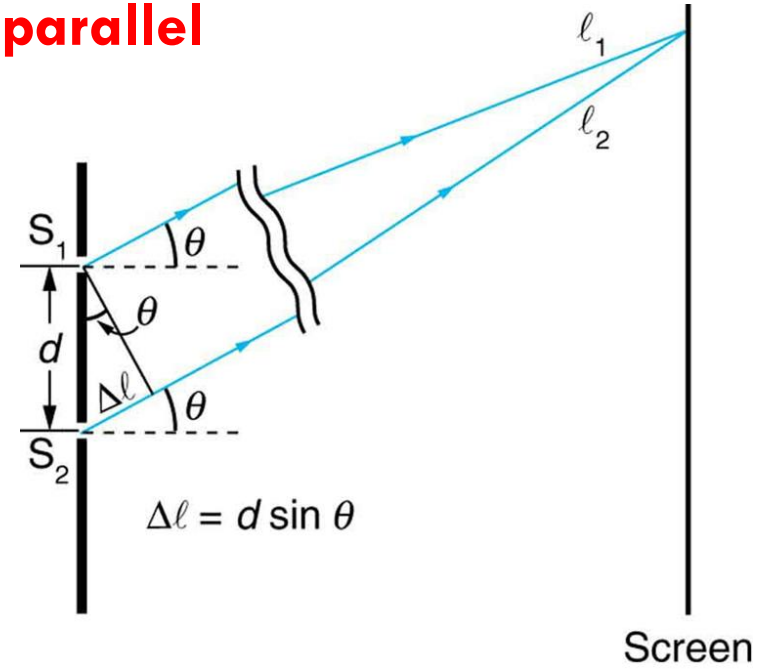
$$\delta = r_2 - r_1 \approx d\sin\theta$$

Diffraction Grating (5)

- Consider monochromatic light.
- Principal maxima ("orders") in the multi-slit pattern occur for $\delta = n$, where n is any integer.
  This implies the path difference Δ between adjacent slits will be n wavelengths.
  If L≫d, the two rays and are essentially treated as being **parallel**

$$\Delta = \Delta l = d \sin \vartheta$$

ϑ – is the angle of diffraction

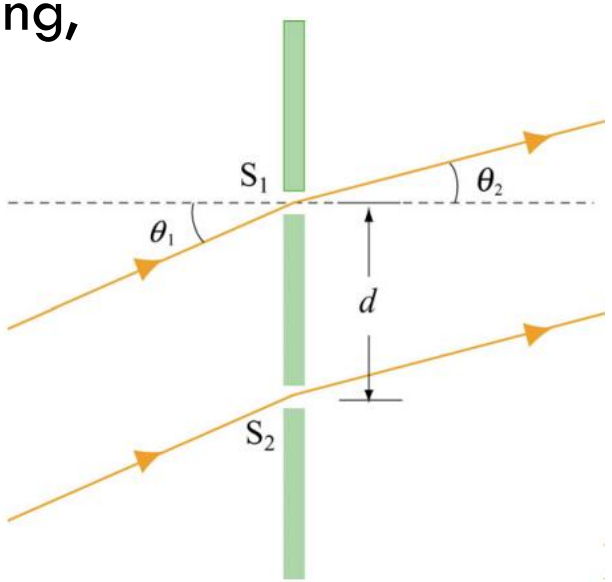


Diffraction Grating (6)

- Consider monochromatic light.
- Principal maxima ("orders") in the multi-slit pattern occur for $\delta = n$, where n is any integer.

If the normal to the plane of the apertures is at an angle ϑ_1 to the incoming radiation, which corresponds to the often utilized inclined grating, then

$$\Delta = d\sin\vartheta_1 + d\sin\vartheta_2$$



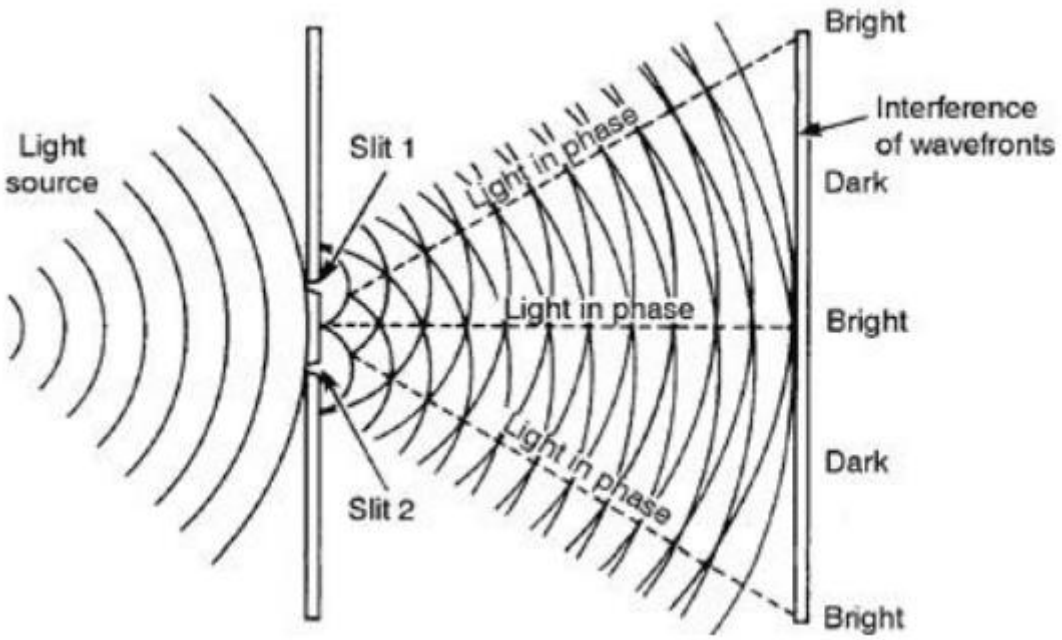
ϑ_1 – is the angle of incidence

ϑ_2 – is the angle of diffraction

Diffraction Grating (7)

- Thus, if the path difference Δ between adjacent slits is n wavelengths, this will produce constructive interference. Maxima in the output intensity occur at a sequence of angles $\sin\theta_n = n\lambda/d$.

$$\delta = n = \frac{d\sin\theta}{\lambda}$$



Diffraction Grating (8)

- Principal maxima are given by the general grating equation:

$$n\lambda = d(\sin\alpha + \sin\beta)$$

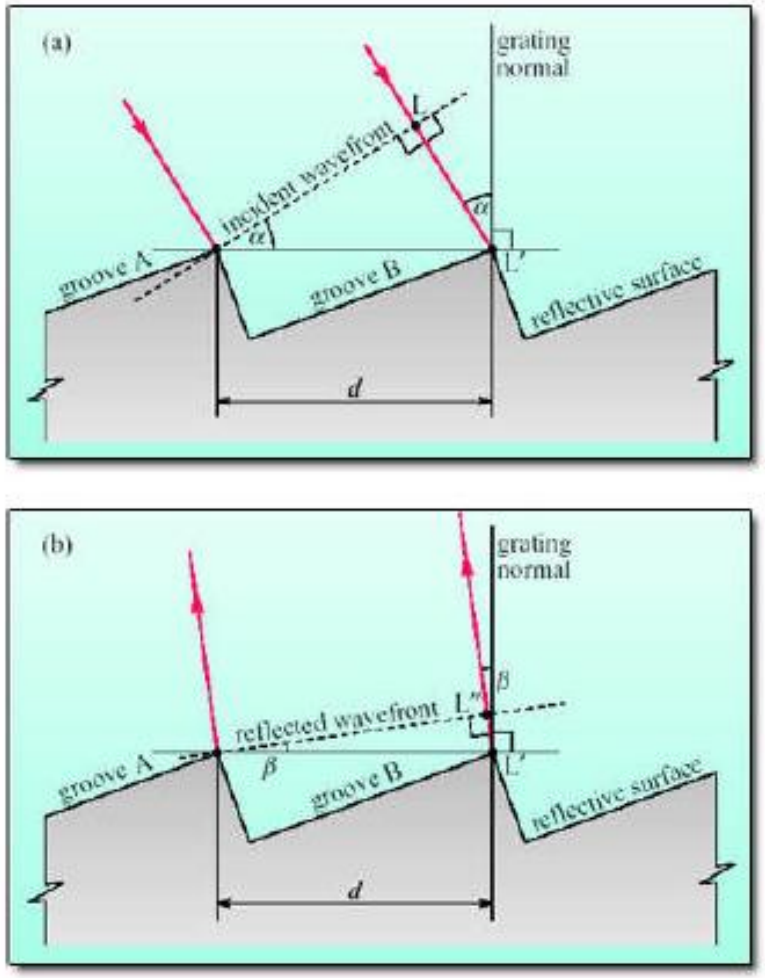
n is an integer representing the order in which the grating is being used.

n is called the order of diffraction.

d – the groove spacing.

α – is the angle of incidence

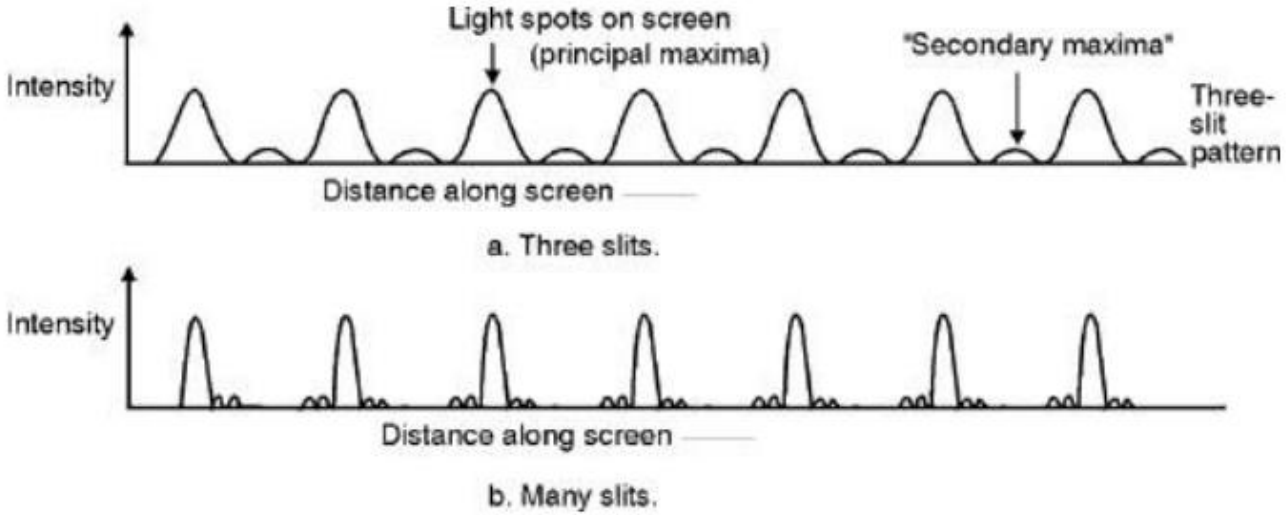
β – is the angle of diffraction



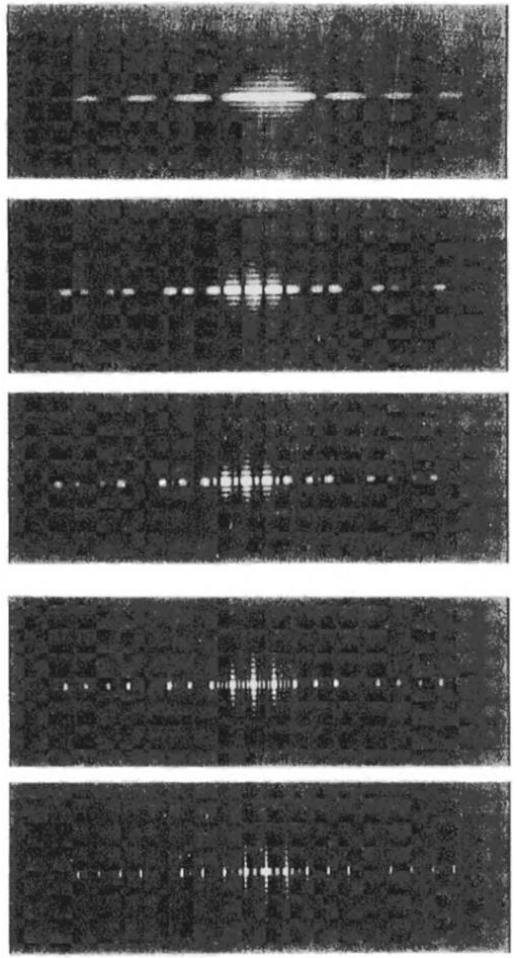
Diffraction Grating (9)

- The single slit diffraction pattern modifies this by affecting the heights of the maxima, the strongest maximum is that at $n = 0$.
- This maximum is of no use to us, because it does not provide any discrimination in wavelength, it is at the same angle for any λ.
- Gratings are designed to concentrate radiation in orders with $n \neq 0$ (note that positive and negative n are equivalent).

Diffraction Grating (10)



The monochromatic multi-slit pattern for 3 slits and a large number of slits. Each peak corresponds to a particular order. The addition of slits **increases the sharpness** and brightness of the peaks **but leaves the locations** of the orders unchanged.



The monochromatic multi-slit pattern for 1 slit (top), 2, 3, 5, and 10 slits (bottom).

Credit: Kitchin C. R.

Resolving power & Spectral resolution

- Spectral resolution for order n is determined by the wavelength shift needed to place the diffraction pattern maximum for $\lambda + \delta\lambda$ on the first minimum in the pattern for λ. The **resolving power** is

$$R = \frac{\lambda}{\delta\lambda} = nN$$



it depends both on the order *n* and on the total number *N* of slits illuminated on the grating.

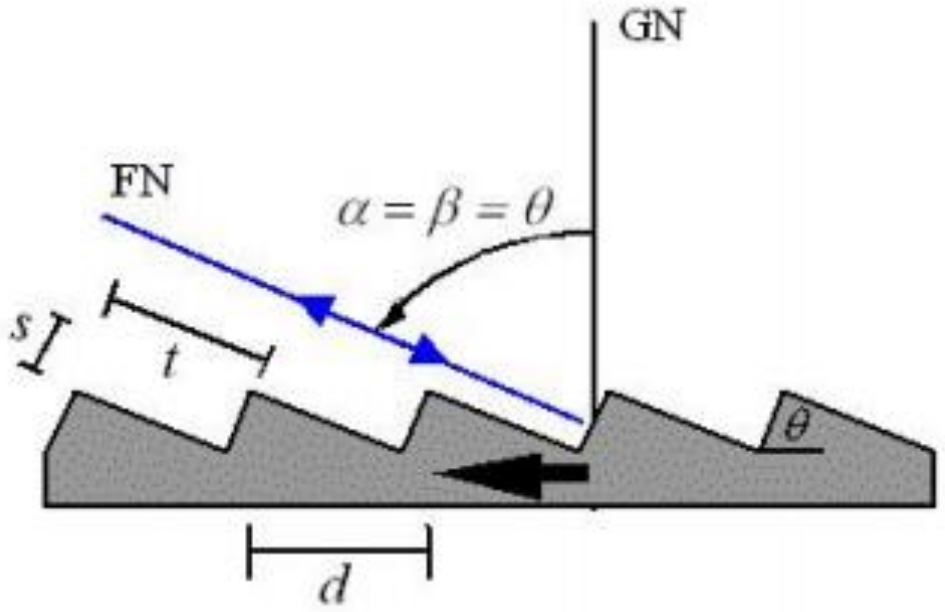
- Astronomers often use the word "dispersion" to refer to $d\lambda/dx$ in the spectrograph focal plane, usually quoted in Å per mm, or Å per pix.
  It is inversely related to the resolving power, so lower values correspond to higher resolving power.

Which resolving power to use for your observations?

- Always „the larger the better“ is not the answer.
- High resolution needs a lot of photons, so to get any signal one needs a bright source and/or a large telescope.
- Also, in some cases there is no need for high resolution. If the process you want to study produces velocities of 1000 km/s, there is not much point studying it with resolution of 1 km/s.
- Still, with high resolution you might discover surprising things about your object.

Diffraction Grating: Resolution

- "Echelle" gratings reach very high resolutions by operating at large n ~ 50 – 100 and angle of incidence α~90°. Yield R > 10^5.



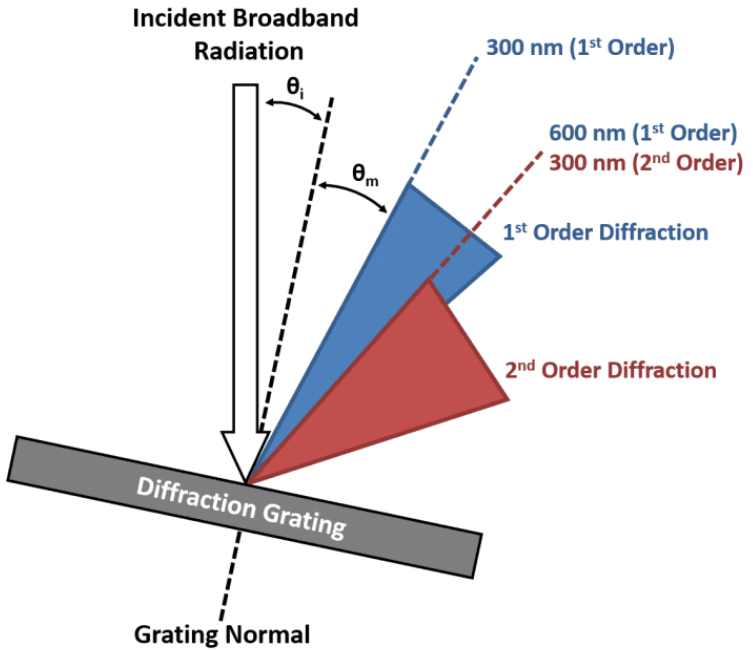
Order overlap

- Red light of a given order is spatially coincident with blue light from a higher order. Wavelength λ_m in order m is superposed on light from wavelength λ_n in order n if

$$\lambda_m = \frac{n\lambda_n}{m}$$

For instance, $\lambda_1 = 10000\text{Å}$, $\lambda_2 = 5000\text{Å}$, and $\lambda_3 = 3330\text{Å}$ are coincident.

Solution: Use "order separating" filters to block out the unwanted orders, (through this becomes difficult for large n).

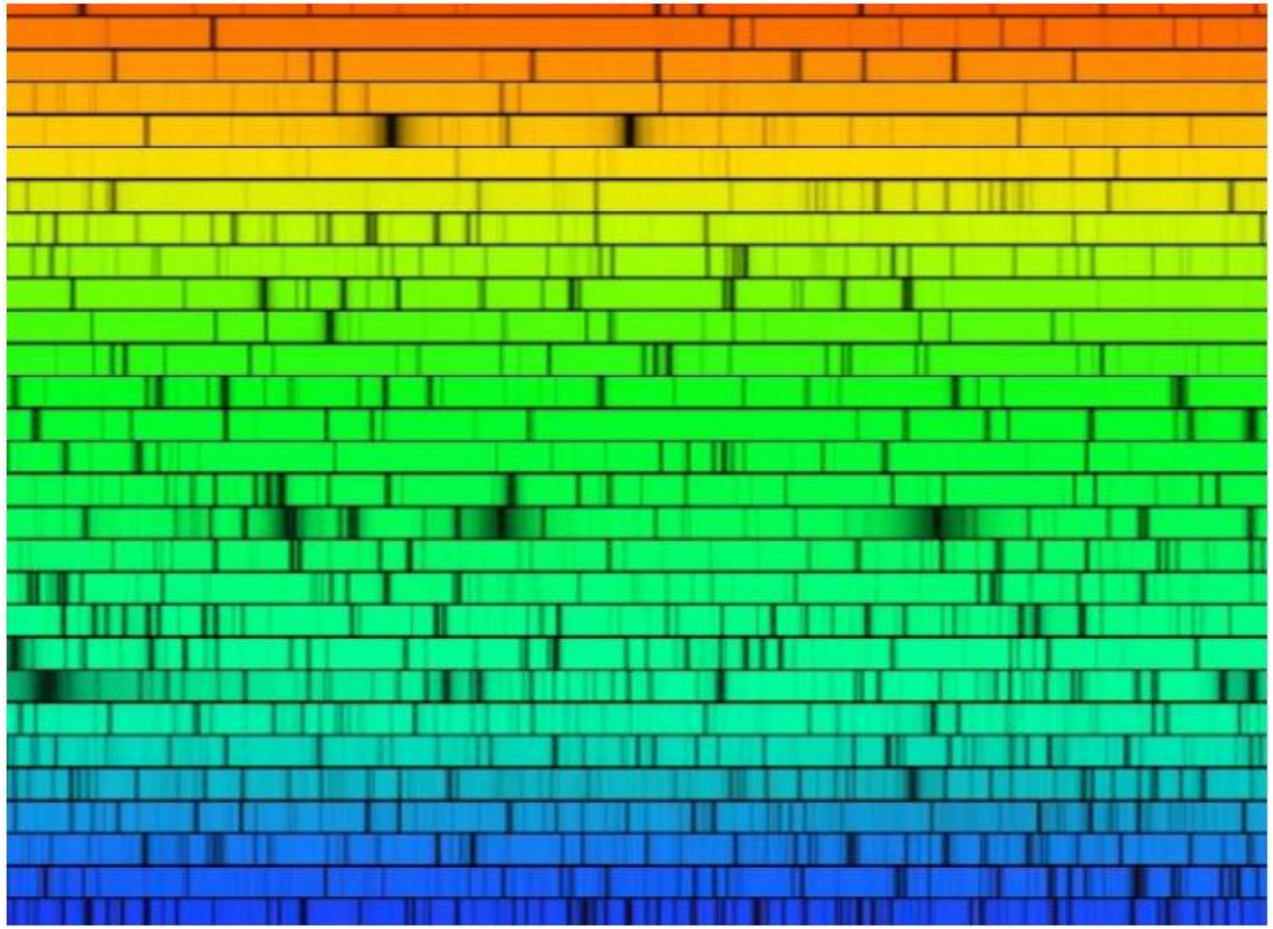


Echelle gratings (1)

- Order overlap is much worse, because adjacent orders differ in wavelength by small amounts (e.g. Order 6 @ 500nm is coincident with order 5 @ 600nm, order 7 @ 429nm, order 8 @ 375nm etc)
- Must separate these orders by cross-dispersion, usually dispersing with a prism at right angles to the grating dispersion.
- Echelle spectrum consists of a number of spectral orders arranged side by side on the detector.
- Echelles can only be used for point sources (stars and quasars) or for small objects, otherwise the light from different orders still overlaps.

Echelle gratings (2)

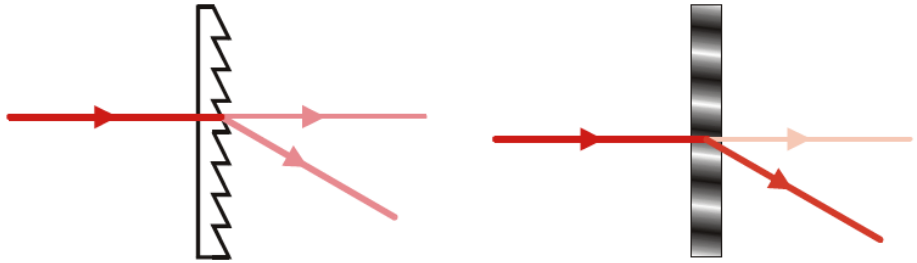
High resolution, optical band solar spectrum



Transmission gratings and grisms

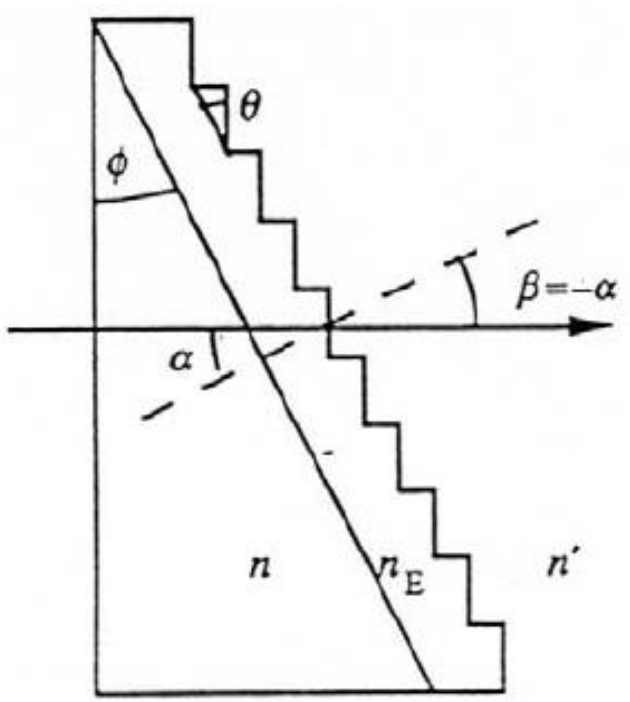
There are different versions of transmission gratings:

- Transmission grating
- Volume Phase Holographic Gratings: **VPH** - use modulations of the index of refraction rather than surface structures to produce dispersion. High efficiency.



Relief Diffraction Grating *VPH Diffraction Grating*

- **Grisms** - a very popular way to convert a camera into a spectrograph is to deposit a transmission grating on the hypotenuse of a right-angled prism and use the deviation of the prism to bring the first order of diffraction on axis.
  - The advantage of a grism is that it can be placed in a filter wheel and treated like another filter. Resolving powers of R ~ 500 – 2000 are practical.



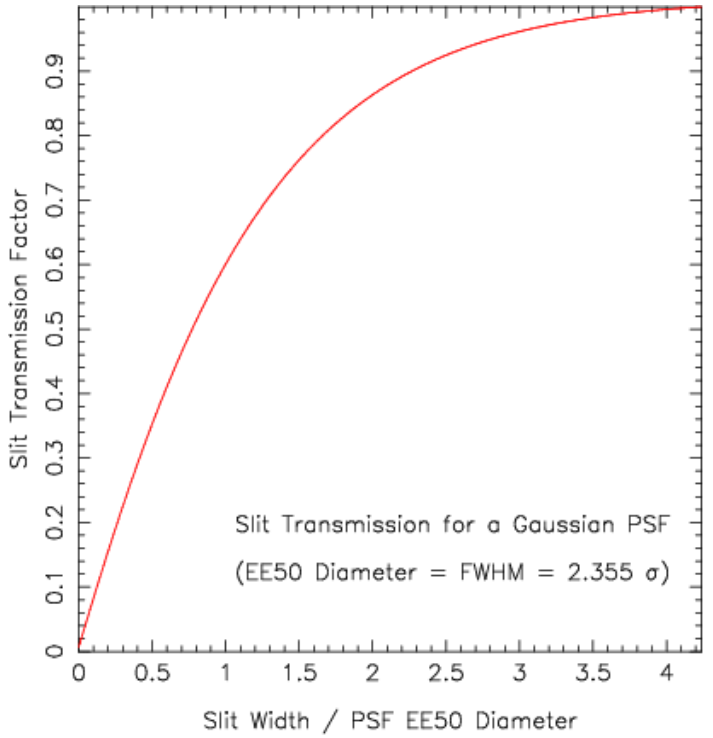
Spectrometer Throughput

- Spectrometer throughput ranges from a few percent to ~50%. The losses accumulate fast. Dispersing elements are usually a big hit, then the losses at multiple surfaces go like (transmission)^n where n is the number of surfaces in the collimator and camera elements (n can be pretty big)

$$0.98^8 \cdot 0.7 \cdot 0.8 = 0.47$$

Camera/collimator with 8 surfaces — grating — ccd

- Another throughput issue: slit losses can be very significant!



ALFOSC at the NOT telescope



- UV-optical imaging, low resolution spectroscopy and polarimetry
- 2048x2048 CCD (0.19”/pixel)
- 6.4’x6.4’ field of view (FOV)
- Large selection of broad, intermediate and narrow band filters
- Several grisms:
  - R~200-10,000 (typically 1000)
  - Velocity resolution 30-1500km/s
- Multi Object Spectroscopy with masks
- Imaging polarimetry and spectropolarimetry
- Spectroscopy of objects brighter than R~20
- Imaging of objects brighter than R~23.5



OBSERVATIONAL ASTRONOMY

FALL 2026

Lecture 14

Vitaly Neustroev

Practical spectroscopy

Practical spectroscopy

- The purposes of spectroscopy are:
  - To measure accurate wavelengths of emission and absorption lines.
  - To measure the relative strengths of emission lines.
  - To measure equivalent widths of spectral lines.
  - To measure the spectral energy distribution of the continuum radiation.

Practical spectroscopy (2)

- Science goals must come first:
  - What are the resolution and S/N requirements?
  - Is there a restriction on exposure time?
  - Decide on the best compromise between these constraints, you will soon enough run out of photons.
- Identify a slit-width/disperser combination that provides the required dispersion and sampling.
  - Seeing or slit-width limited?
- Work out calibrations required
  - Always try to take calibration data through the same/similar lightpath.

Slit-filling effects

- A slit-limited setup ensures that the slit is always illuminated uniformly [provided the object is centered]
- A partially illuminated slit (because image quality is better than the slit-width) may introduce shifts in the projected spectrum as different areas of the slit are illuminated as a function of time
- This will lead to shifts in both the spatial and dispersion direction of the spectrum when comparing to calibration data that are obtained with a fully illuminated slit

  Not good if you are after accurate radial velocities!

Atmospheric dispersion (1)

- Differential atmospheric refraction will deflect a source by an amount that is dependent on wavelength:
  - the index of refraction is a function of wavelength.
- A point source position on the sky is dependent on wavelength!
- The displacement is towards the zenith and larger for shorter wavelengths.
- This obviously affects acquisition and slit-angle strategies when obtaining spectroscopy.

Atmospheric dispersion (2)

- Index of refraction: $n(\lambda,T,p,f)$
  - wavelength, temperature, pressure, water vapour
- Angle displacement:

$$\Delta\theta = 206265 \times [(n_{\lambda 1}-1)-(n_{\lambda 2}-1)] \times \tan z$$

index of refraction

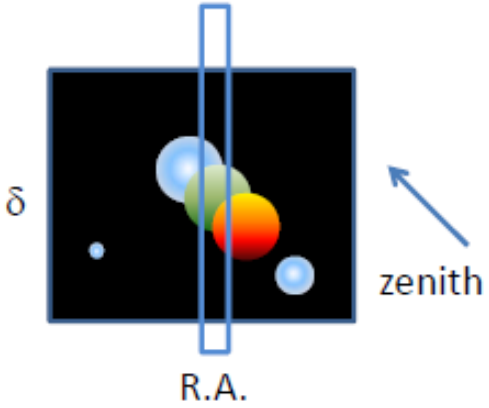
zenith angle (airmass)

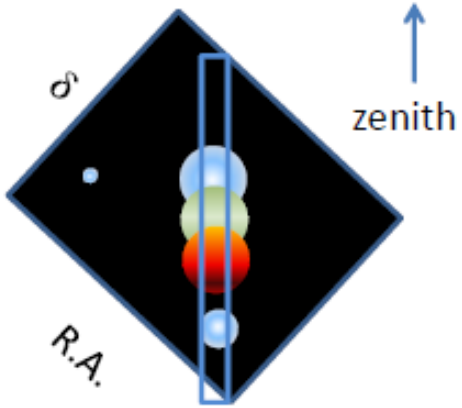
- Some example shifts (“) relative to image at 5000Å:

airmass	3000Å	4000Å	6000Å	10000Å
1.00	0.00	0.00	0.00	0.00
1.25	1.59	0.48	-0.25	-0.61
2.00	3.67	1.10	-0.58	-1.40

Atmospheric dispersion (3)

- Make sure you acquire the target at a wavelength relevant for your spectral range
- Differential refraction will mean differential slit-losses: can only centre object at one λ
- If the slit is vertical (relative to horizon/ zenith line), differential refraction will occur purely along the slit
- This means that the slit P.A. (sky angle) must change with time. The vertical P.A. is the *parallactic* angle





Calibration

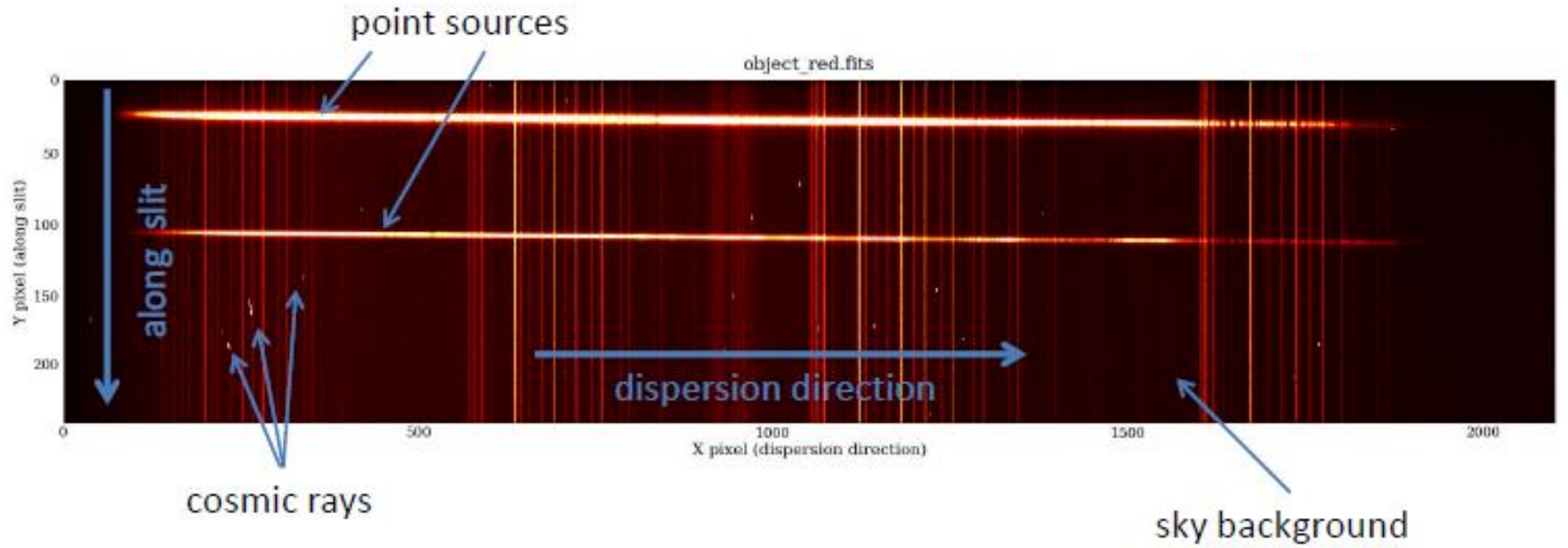
Primary reduction

Extracting the spectrum

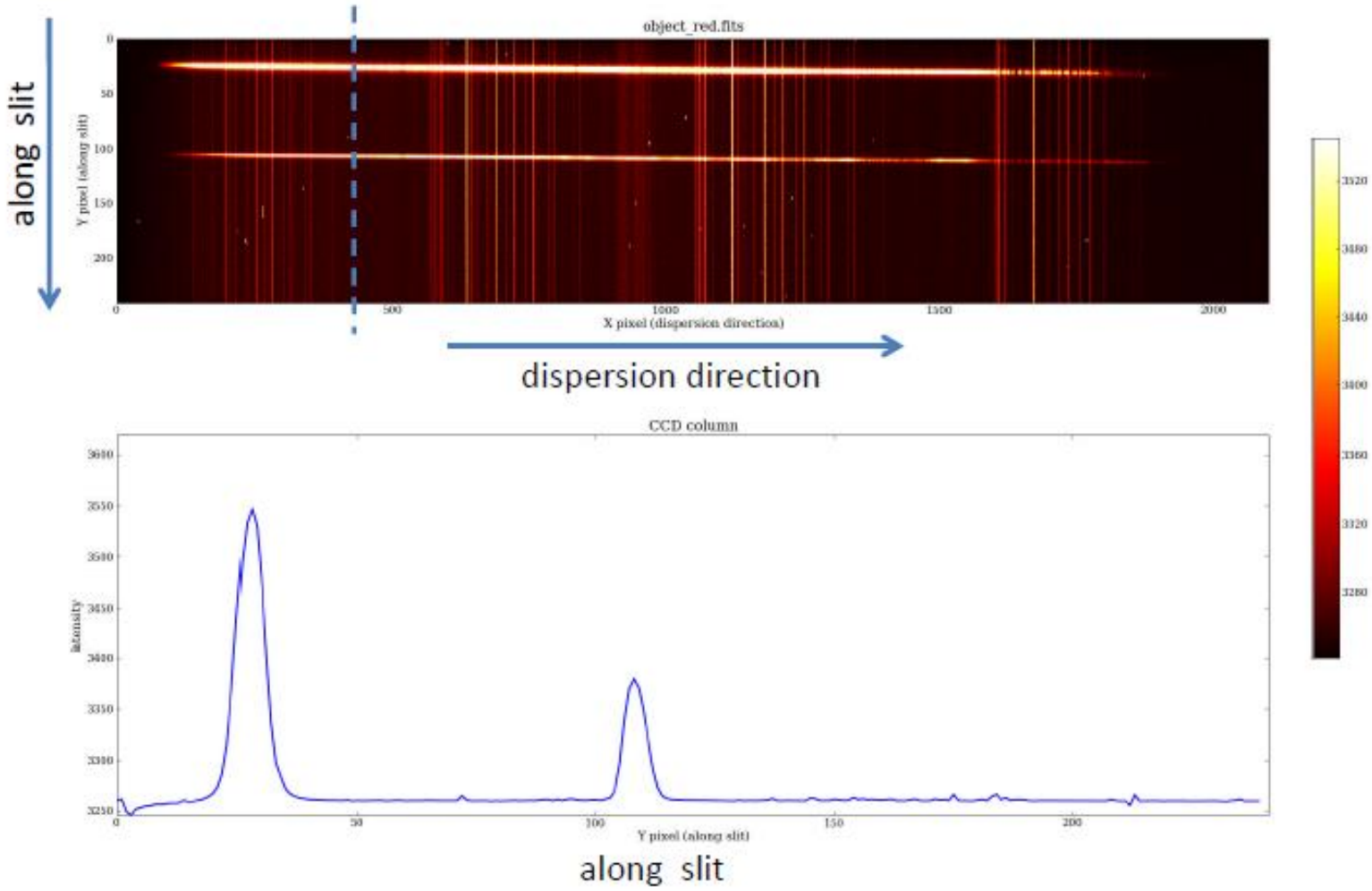
Wavelength calibration

Spectrophotometric calibration

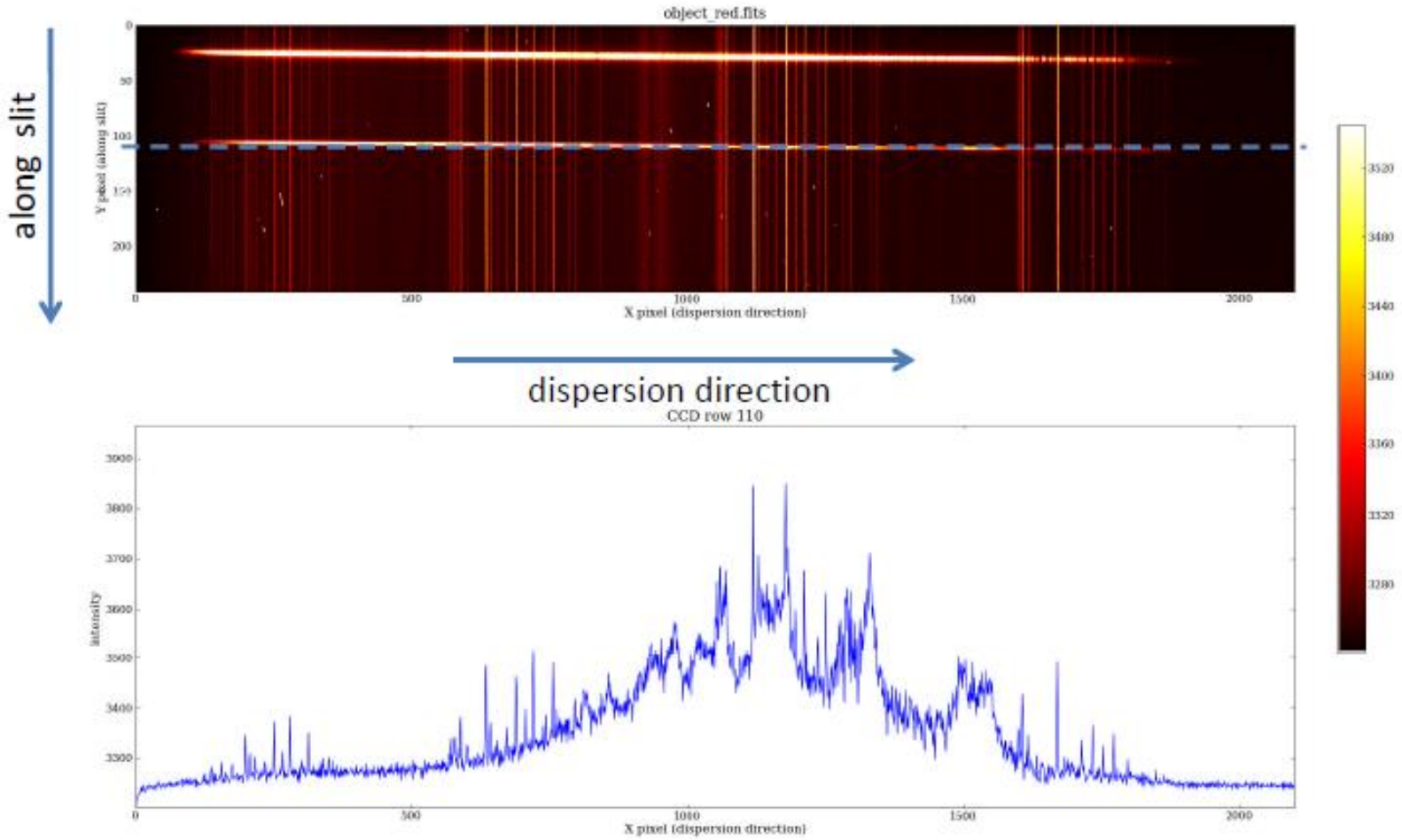
A long slit spectrum



A long slit spectrum: spatial slice



A long slit spectrum: spectral slice



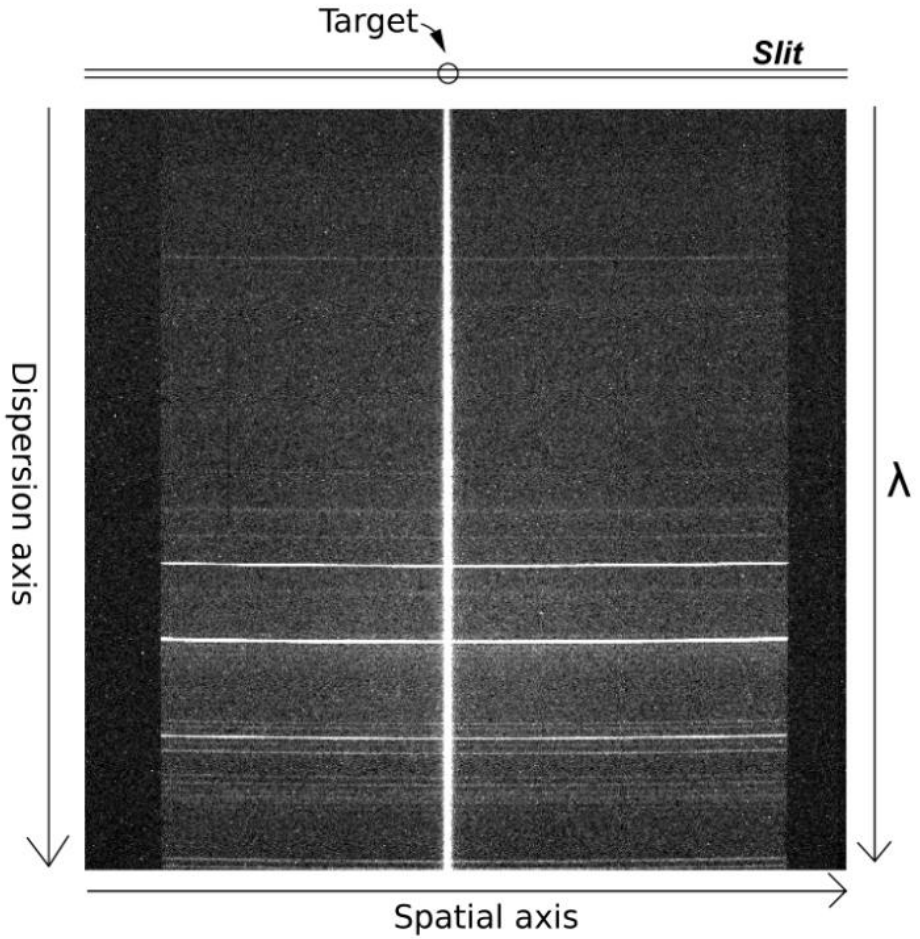
Reducing spectra

- In addition to the science frames you will need:
  - Bias frames
  - A continuum lamp image (for flat-fielding)
  - A line lamp (so-called arc) frame (for wavelength calibration)
  - A standard star spectrum (for flux calibration)
- The continuum and line lamps are inside the instrument in a special calibration unit. They are obtained immediately before or after the science observation

Calibrations: primary reduction

- The first stage of calibration is to calibrate the detector, and the steps in doing this are exactly the same as for photometric observations.
  - Subtract off the CCD bias signal, either as a constant value or as a frame. This step is not needed for photon counting detectors
  - Subtract off the dark current, either as a constant value or as a frame. As spectroscopic exposure times are longer than photometric exposure times, this step is now more often needed.
  - Divide by the flat field frame to correct for variations in the sensitivity of the detector.

An ALFOSC long slit spectrum

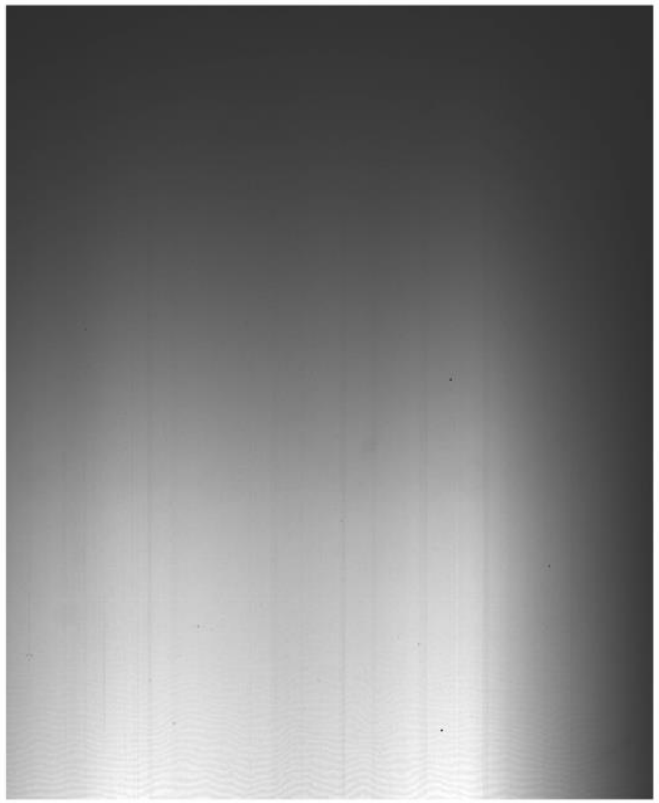


A simple rule:

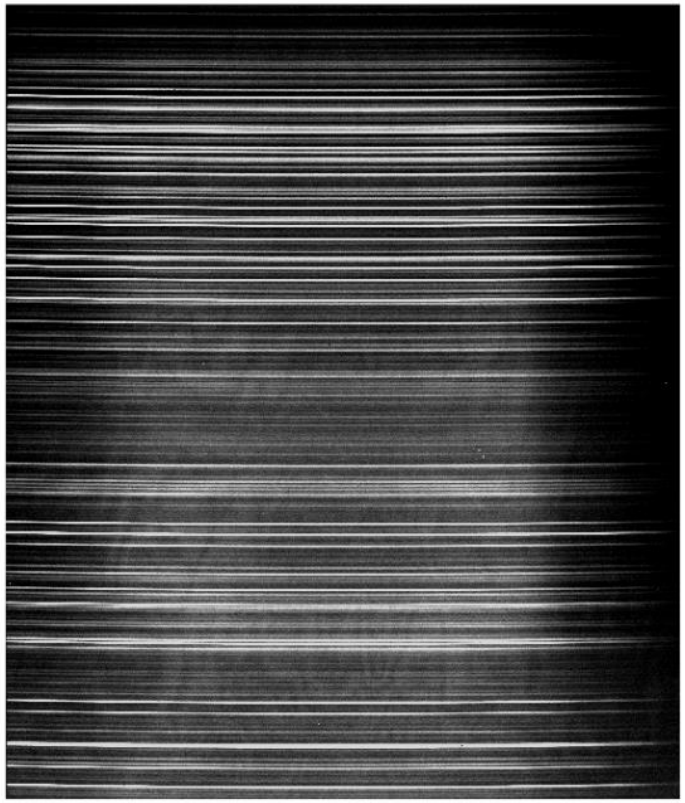
The spectrum on the CCD can be thought as the *image of the slit at different wavelengths*.

Calibrating spectra

Continuum lamp
(For flat-fielding)

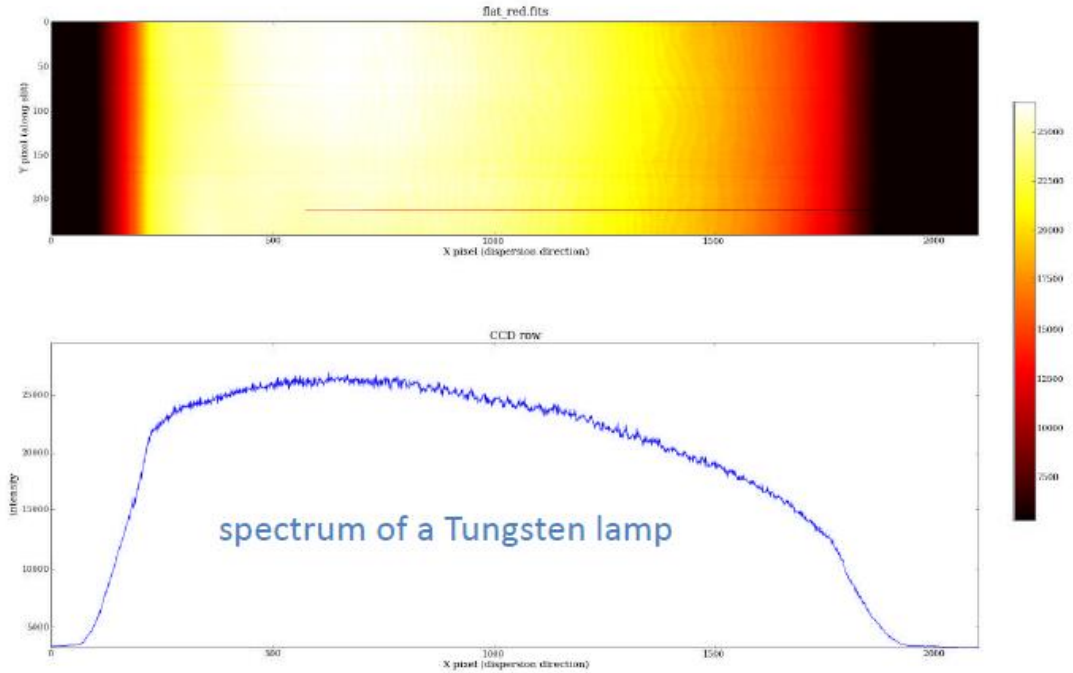


Line lamp
(For wavelength calibration)



Flat-fielding (1)

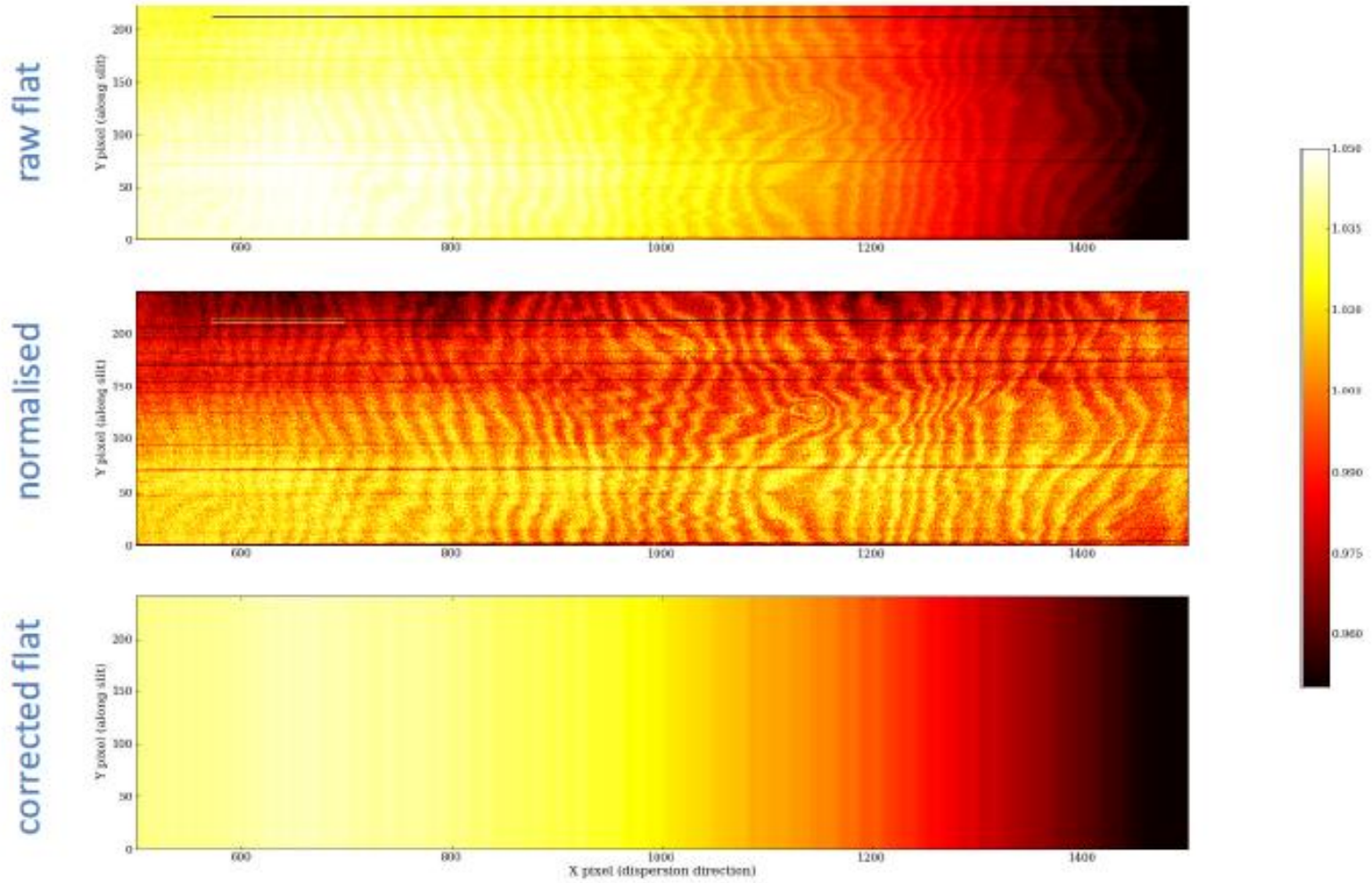
- Flat-fielding is probably one of the trickier steps
  - Uniform illumination along the slit
  - Uniform illumination along the dispersion direction
    - Need a light source with a smooth/simple spectrum



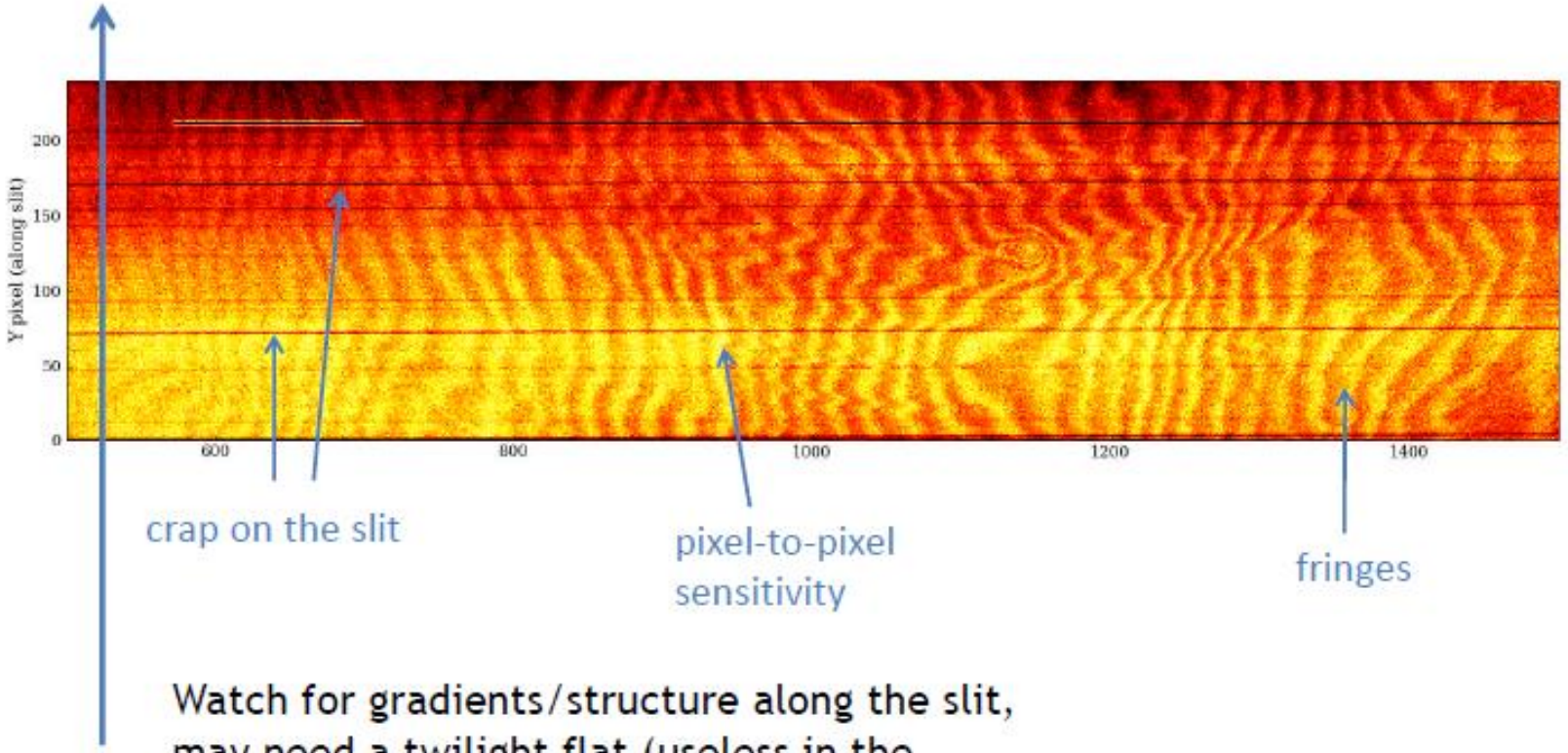
Flat-fielding (2)

- The trick is to remove the spectrum of the calibration lamp and normalise the flatfield
  - Not always possible to distinguish between broad CCD sensitivity features and features in the lamp

Flat-fielding (3)



Flat-fielding (4)



Watch for gradients/structure along the slit, may need a twilight flat (useless in the spectral direction) to correct spatial profile

make sure slit width, grating angle, filters are all in place, replicating as much the light path to the science frames

Extracting the spectrum

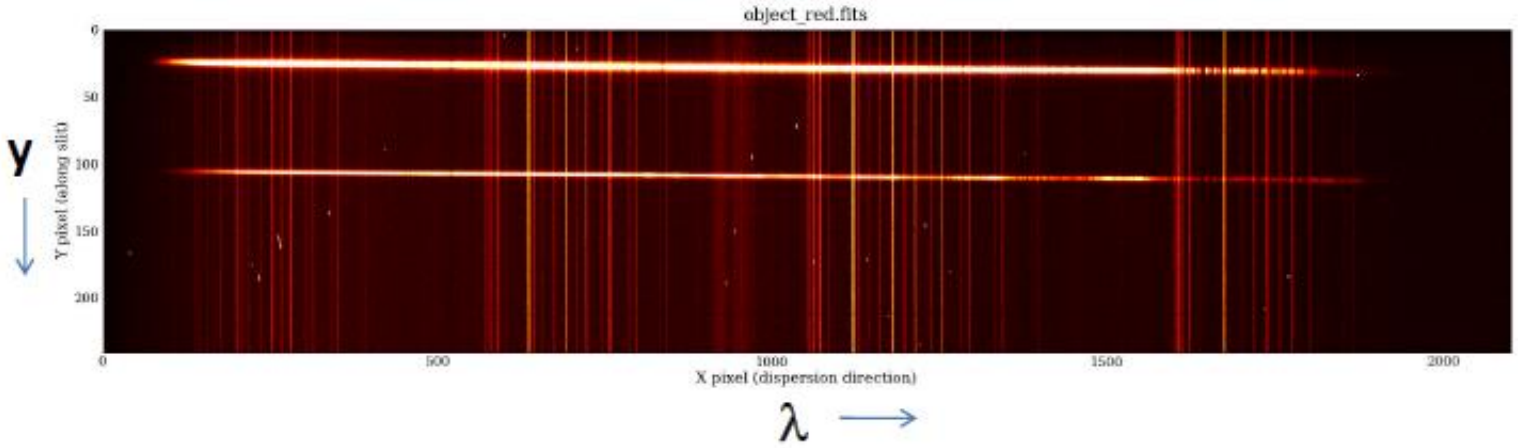
signal = (source + background) – background@source

$$\mathbf{S}(\lambda) = \sum I(y,\lambda)\, p(y) - \sum I(y,\lambda)\, b(y)$$

object profile weight — sky profile weight

$$D(\lambda) = f(x,y) \approx f(x)$$ relates λ to x,y

dispersion relation

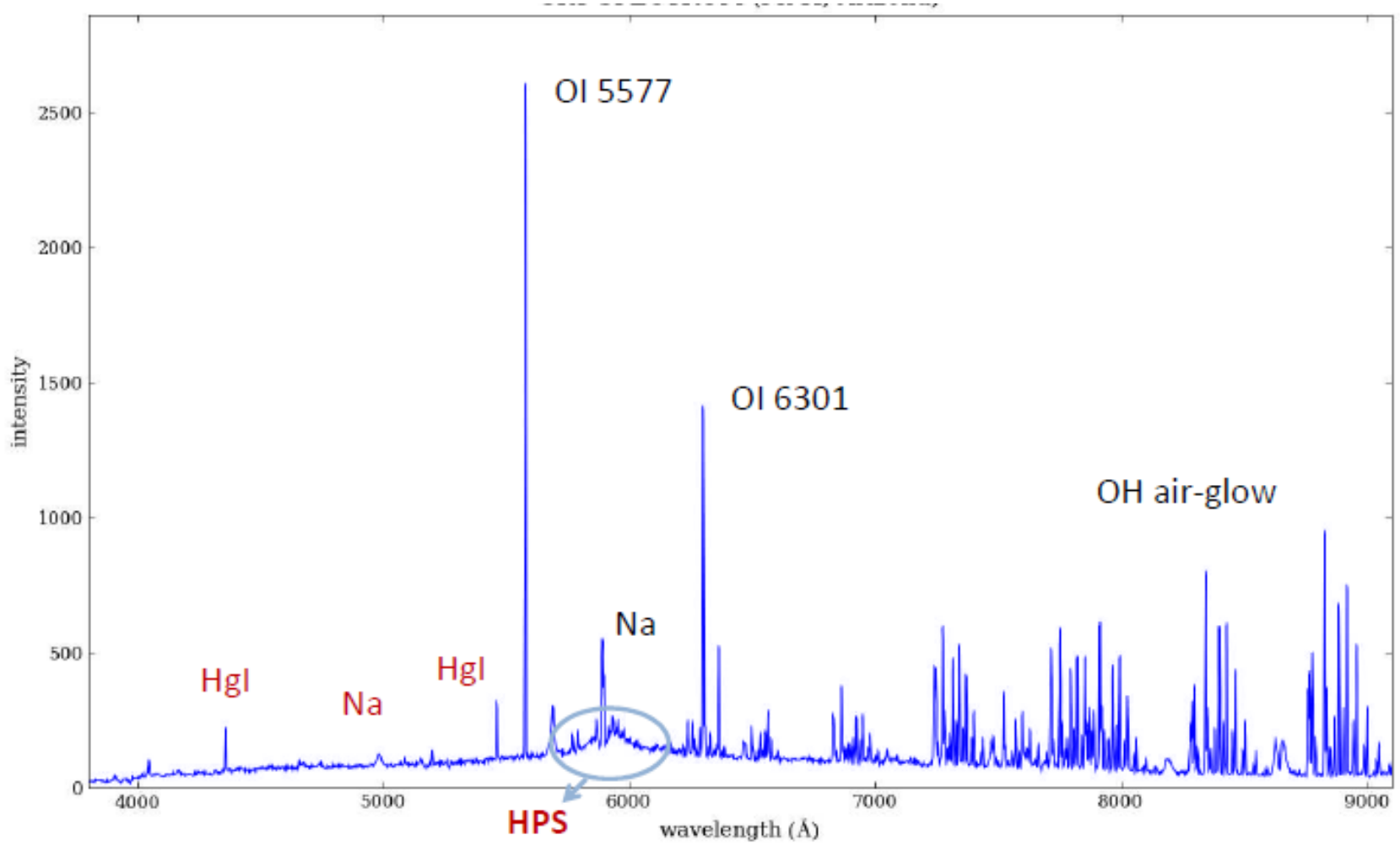


Sky background (1)

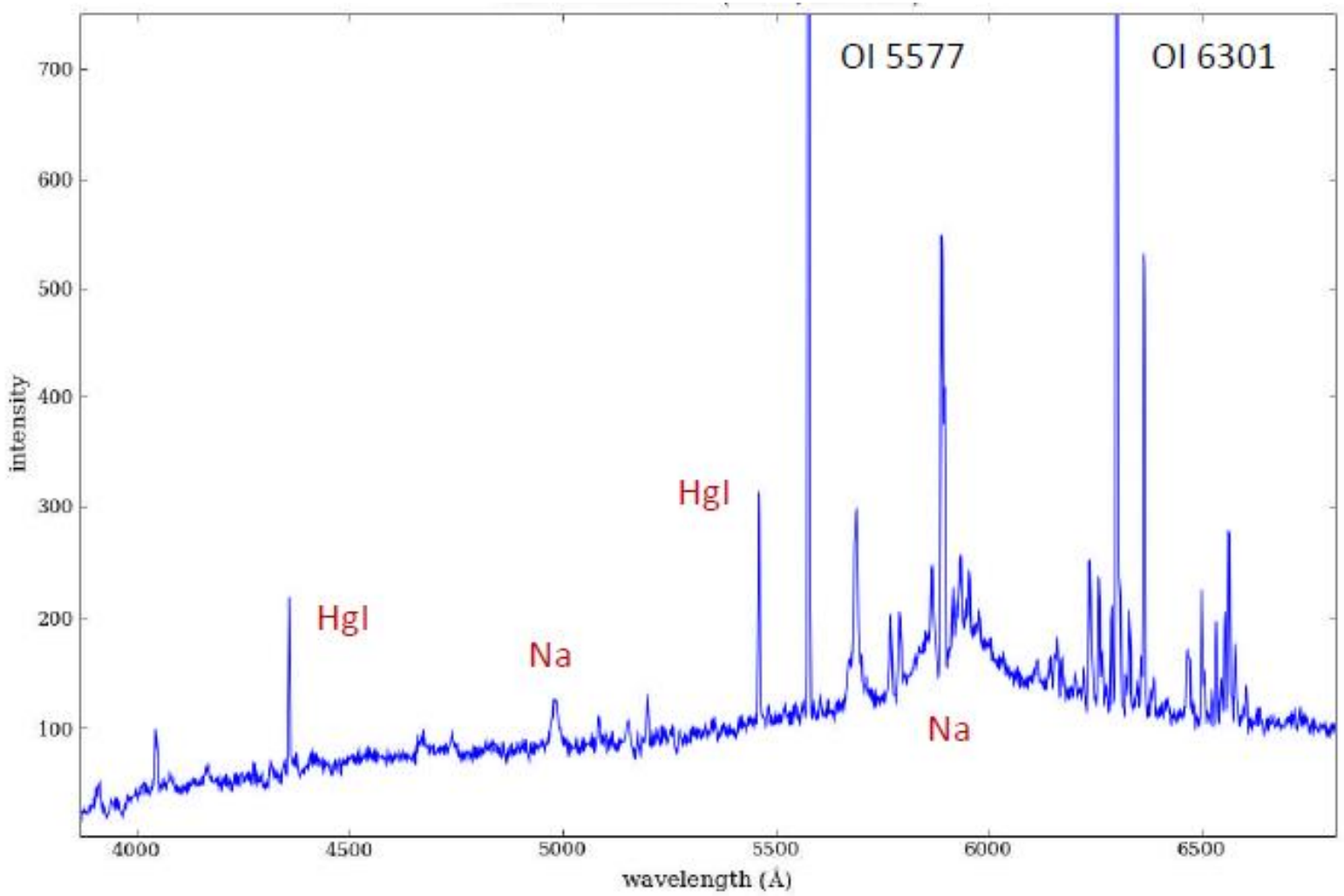
Background has contributions from many sources;

- Air glow ; strong discrete emission lines
- Zodiacal light ; $m_V \sim 22.-23.5$
- Sun/Moonlight
  - new moon : $m_V \sim 21.9$
  - full moon : $m_V \sim 19.9$
- Aurorae
- Light pollution
- Thermal emission from sky, telescope and buildings
- Non-resolved astronomical background

Sky background (2)

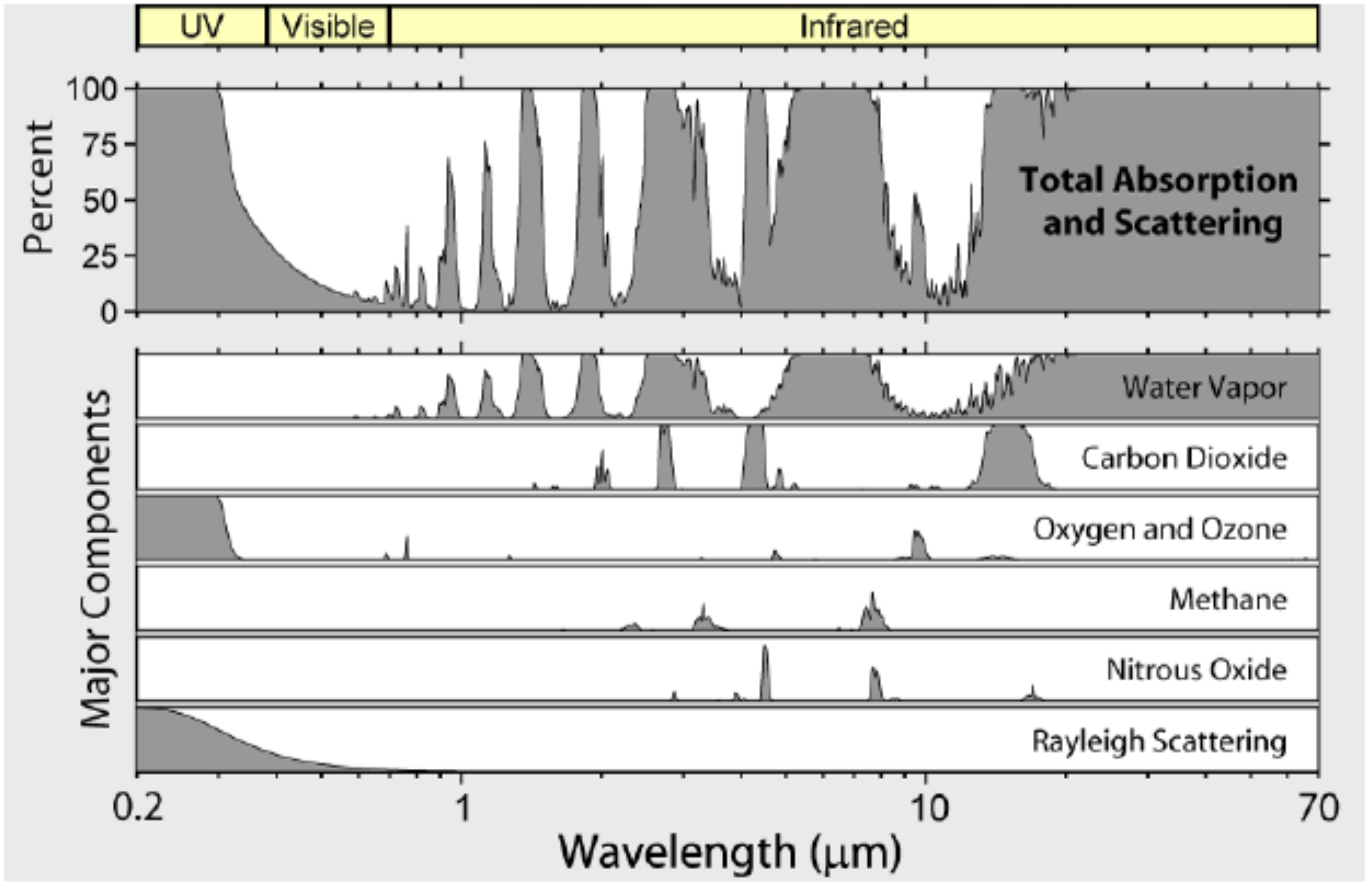


Sky background (3)

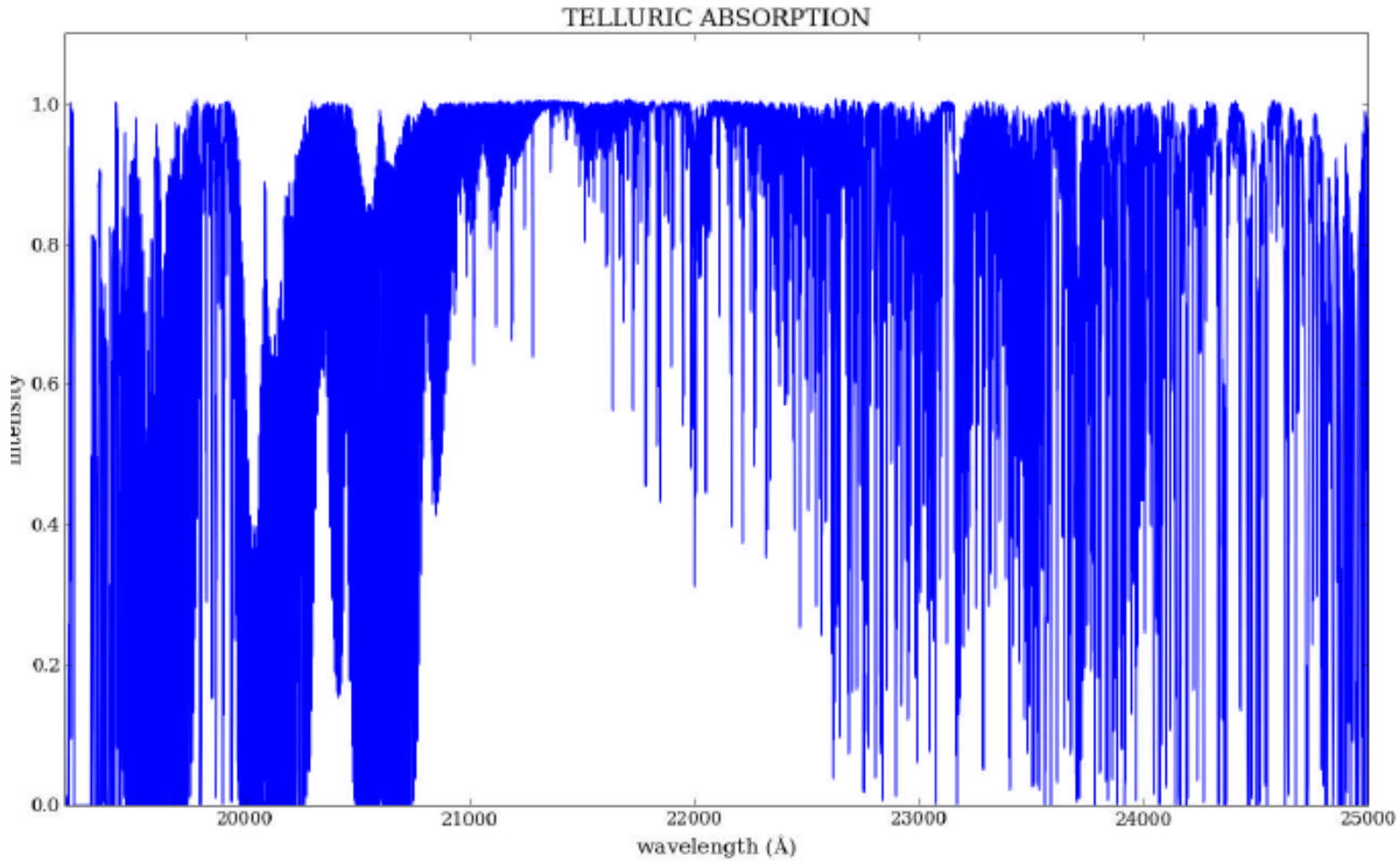


Atmospheric transmission

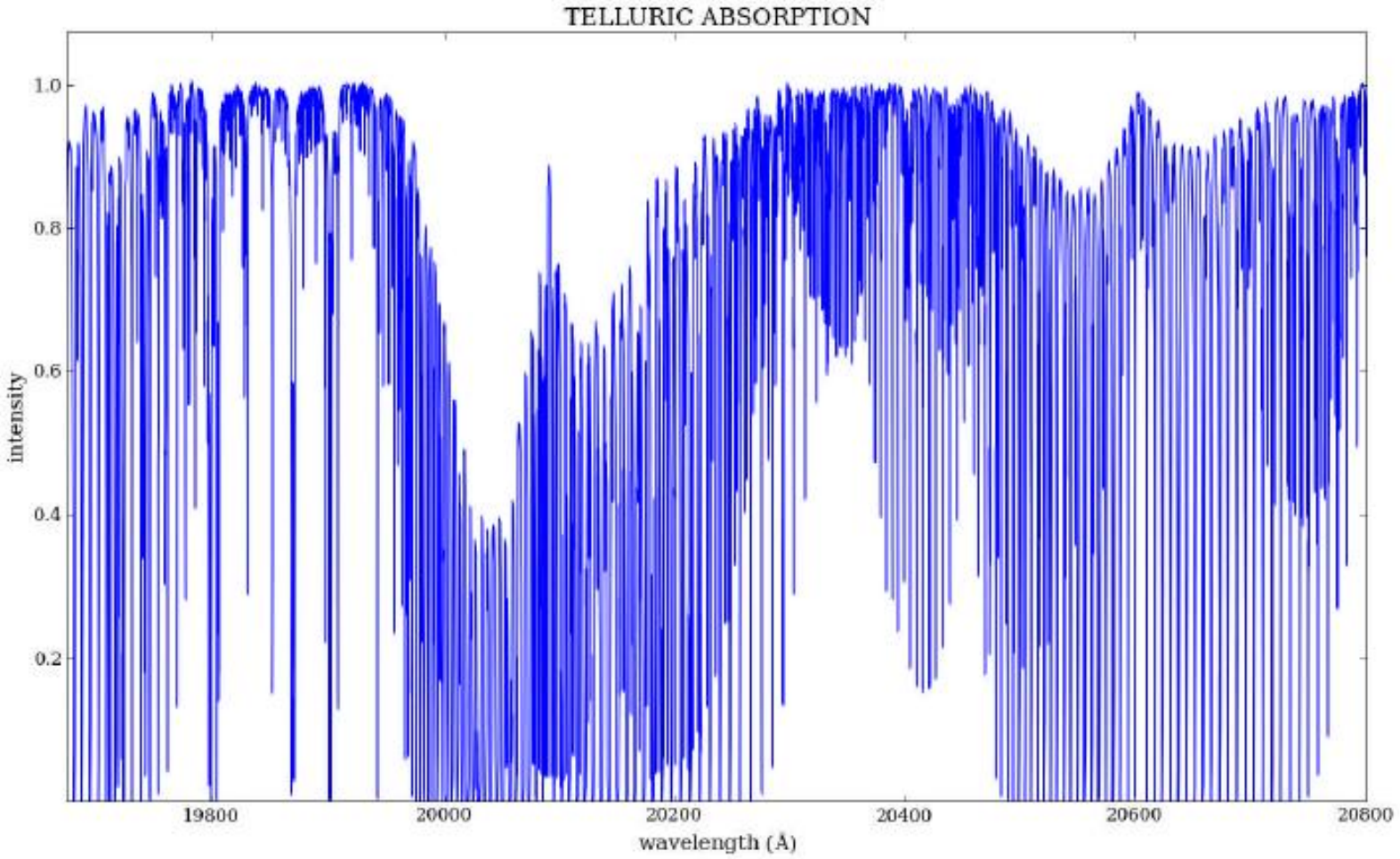
Atmospheric transmission is strongly dependent on wavelength



Telluric absorption (1)



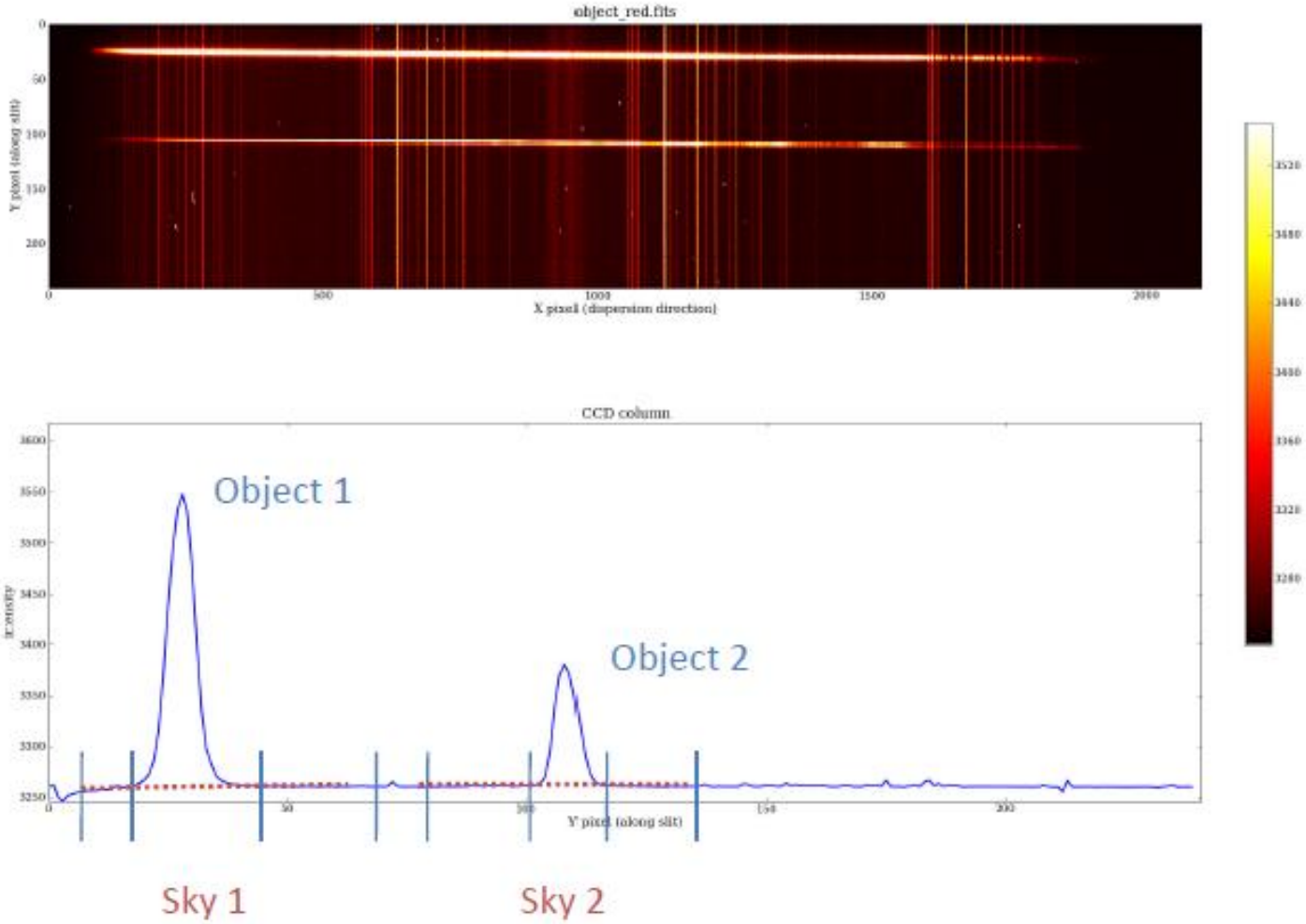
Telluric absorption (2)



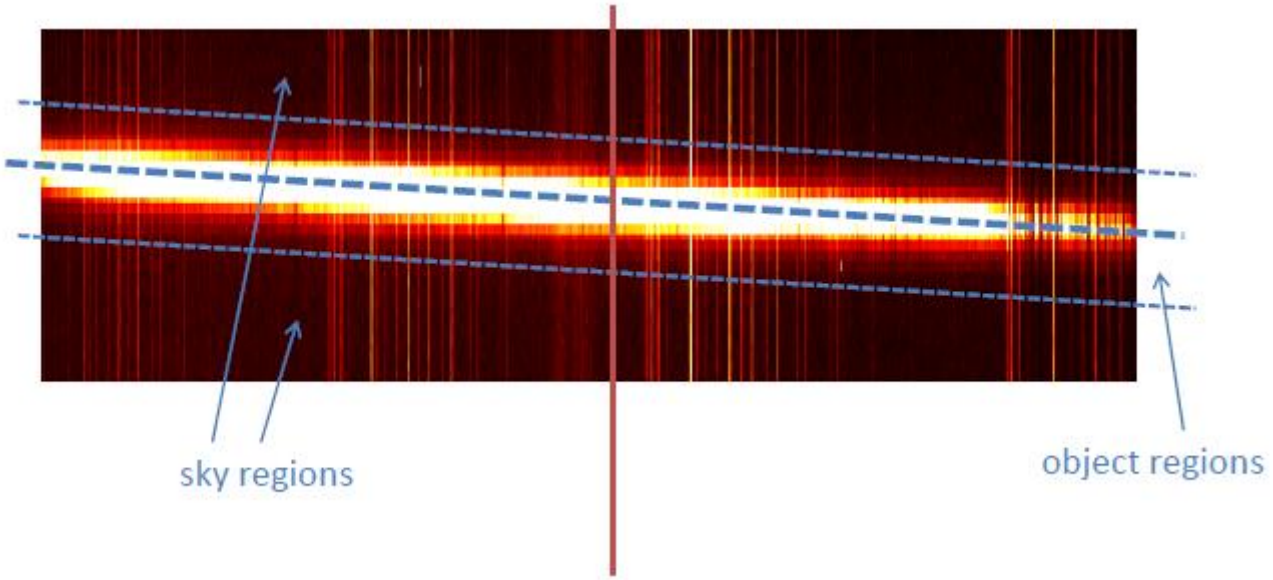
Summary: Background

- The background is a composite of many sources
- All of these are dependent on wavelength and their strength varies with time
- Some correlate with lunar cycle, airmass, solar activity cycle, etc, but many variations are erratic
- *Background subtraction needs to be done on a wavelength by wavelength basis* and ideally is measured simultaneously with the object exposure
- Some parts of the spectrum may be background dominated, others not*; error propagation*

Locating the object and sky

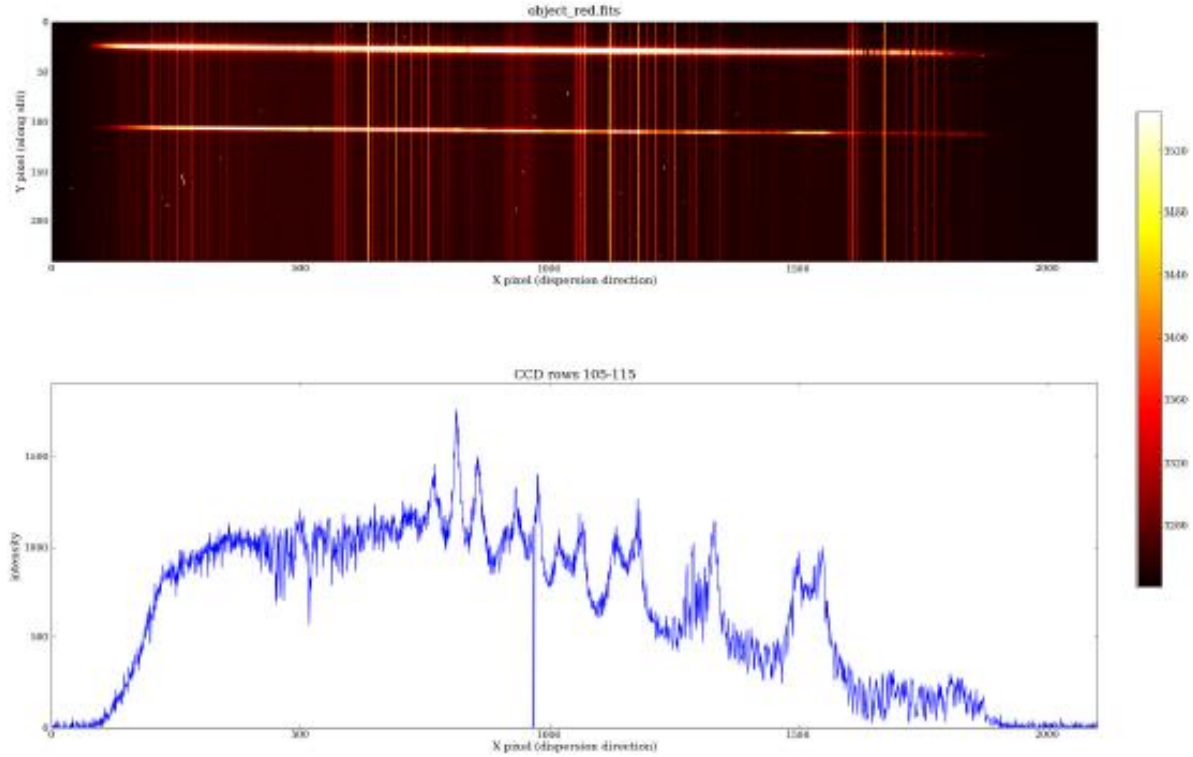


Tracing and skyfit



- Evaluate sky background at each wavelength by considering the sky pixels around the shifting object *[if you are lucky, sky lines are well-aligned with the CCD columns]*
- This gives you the *fitted* background value **at the location of the object**

Net signal: naïve method



$$\mathbf{S}(\lambda) = \sum \left(I(y,\lambda) - sky(y,\lambda) \right)$$

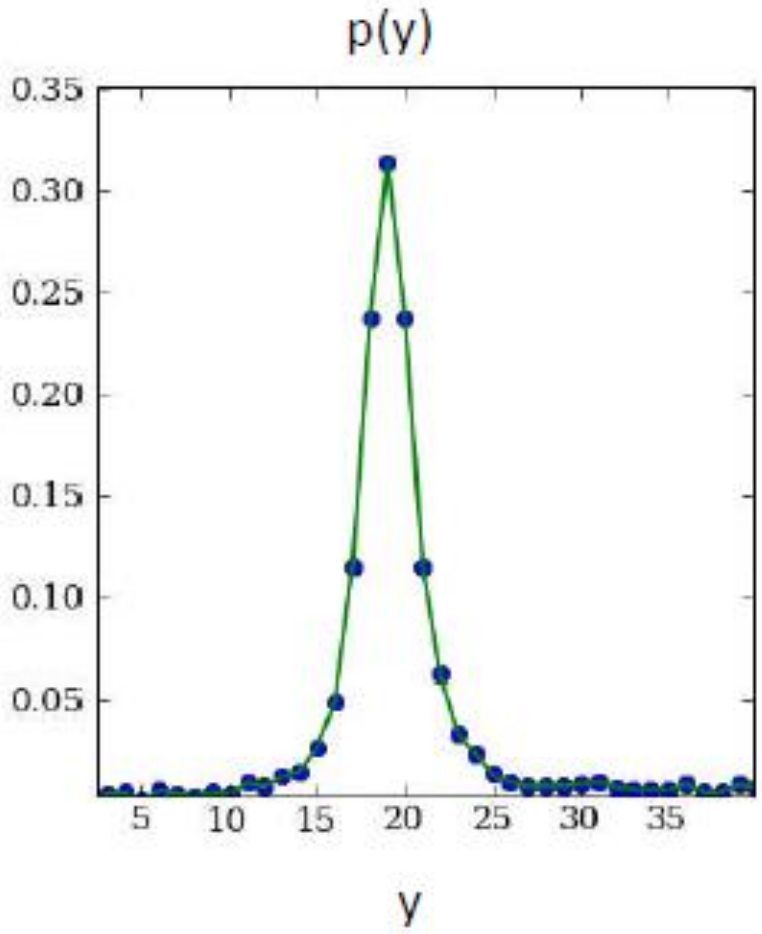
unweighted sum over object region

skyfit at each object pos

Net signal: optimal extraction (1)

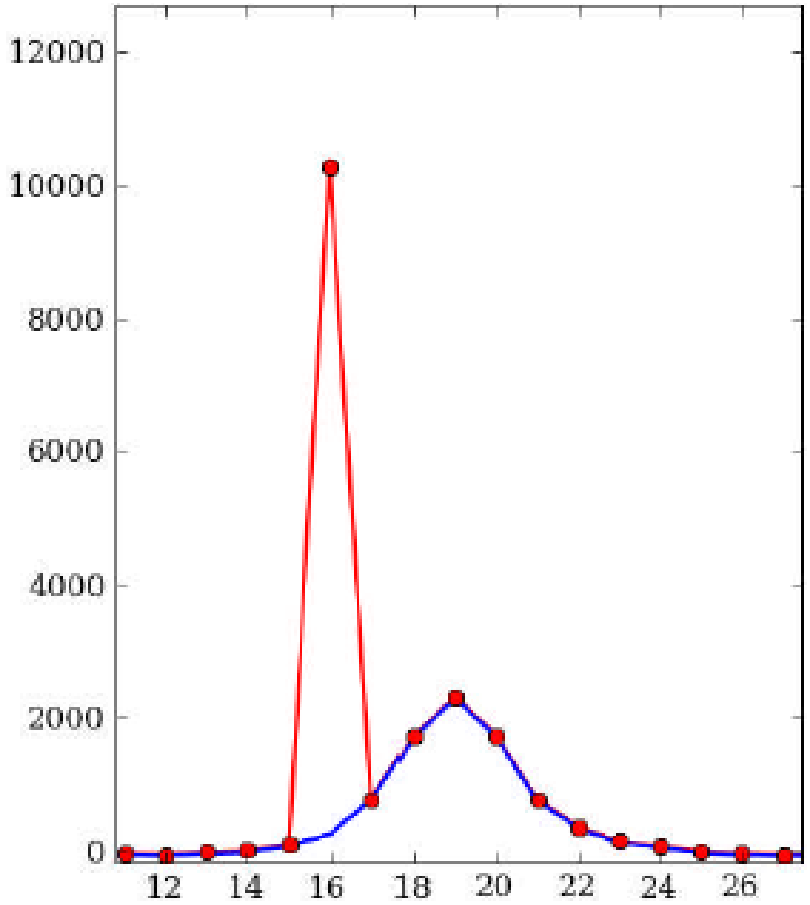
- Optimal sum across the object profile considers the fractional contribution of a given pixel to the total light at that wavelength and weighs its contribution
- Profile p(y) is measured from the 2D frame and normalised such that Σ p(y) = 1 when summed along the extraction region
- Optimal weights (minimising variance):

$$w(y) \propto p(y)/\sigma(y)^2$$



Net signal: optimal extraction (2)

- Need to estimate $\sigma(y)$ reliably from readout noise and gain such that Poisson noise and background subtraction errors are properly propagated
- Not only provides the optimal sum with a significant S/N improvement in the extracted spectrum, but also allows the easy flagging of outlier pixels due to cosmic rays or CCD defect



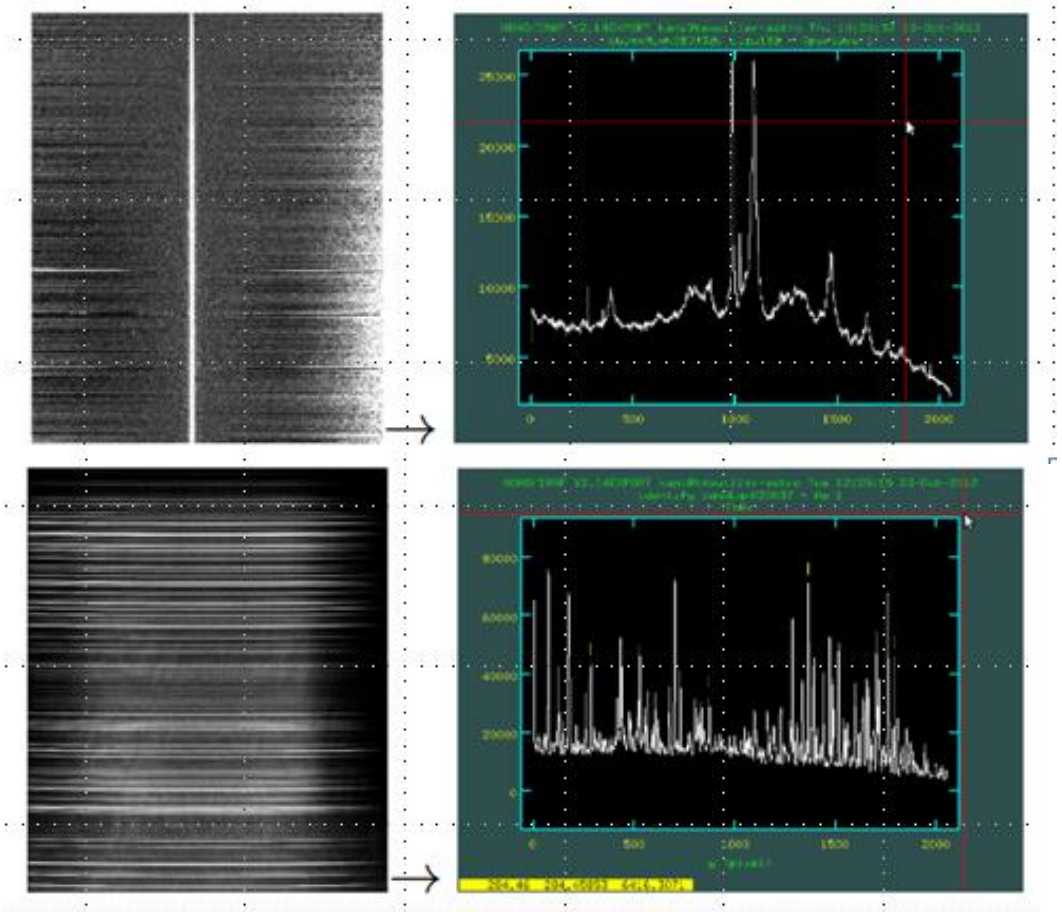
Wavelength calibration (1)

- **Wavelength calibration** – a comparison spectrum usually of a hollow cathode discharge lamp, gives a series of emission lines of the gas in the lamp, plus the metal or metals that the cathode is made from.
- Typically the gas is a noble gas (helium, argon, neon etc.) and the metal is copper, iron or thorium.
- Using the laboratory determined wavelengths of these lines a functional fit of wavelength against position on the detector is made.
- In principal this is a two dimensional fit, although in practice the dispersion direction is usually accurately aligned with one of the principal axes of the detector (usually vertical on a CCD), so this reduces to a series of one dimensional fits, one per CCD column.

Wavelength calibration (2)

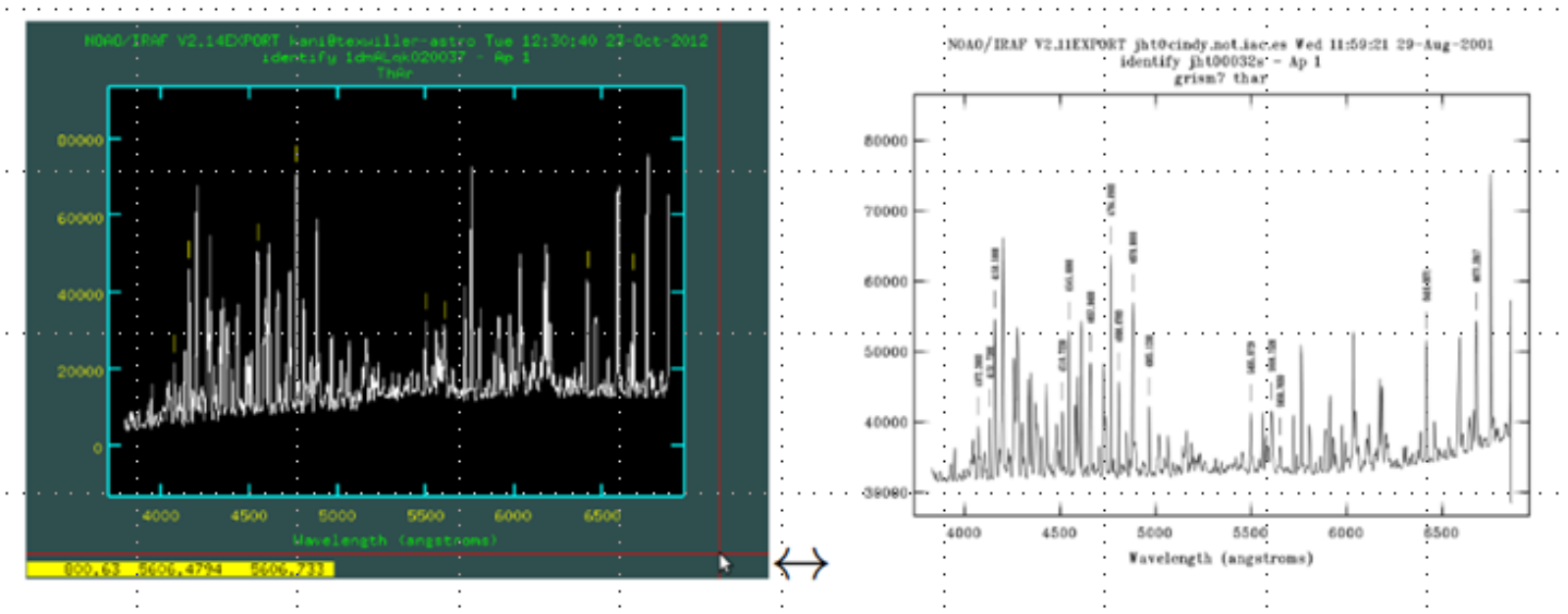
Extract the 1-d spectrum of the star

Extract the lamp (arc) spectrum



Wavelength calibration (3)

- Determine wavelengths of arc-spectrum lines and the transformation equation from pixel coordinates to λ



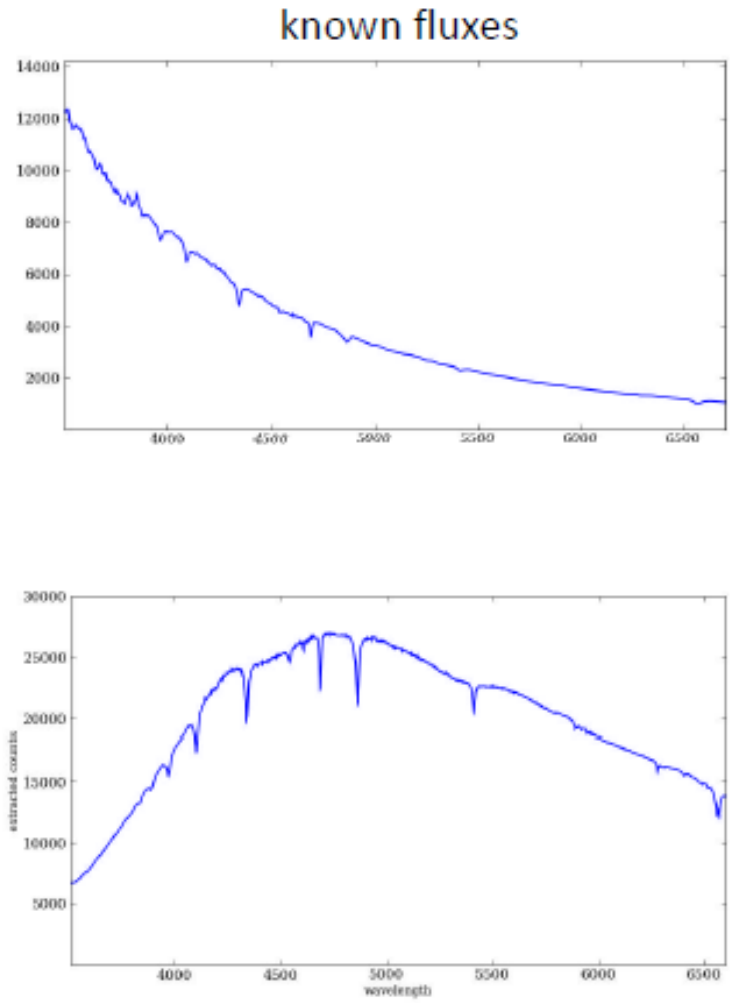
- Wavelength calibrate the science and standard spectra using the pix → λ transformation

Spectrophotometric calibrations

- **Spectrophotometric calibration or Flux calibration.** This is a calibration of sensitivity and efficiency, and is carried out in the same way as the photometric calibration, by observing a number of standard stars whose flux as a function of wavelength is accurately known, at a variety of airmass values. There is one extra quite serious problem.
  - The slit size α is set so that the size projected on the detector p is of order 2 detector pixels, this gives the maximum spectral resolution. α is normally smaller than the resolution set by seeing, so light is lost at the entrance slit.
  - The amount of light lost varies between exposures, making an absolute flux calibration very difficult.
  - The amount of light lost is also wavelength dependent, due to the weak dependence of seeing on wavelength, and due to atmospheric dispersion.

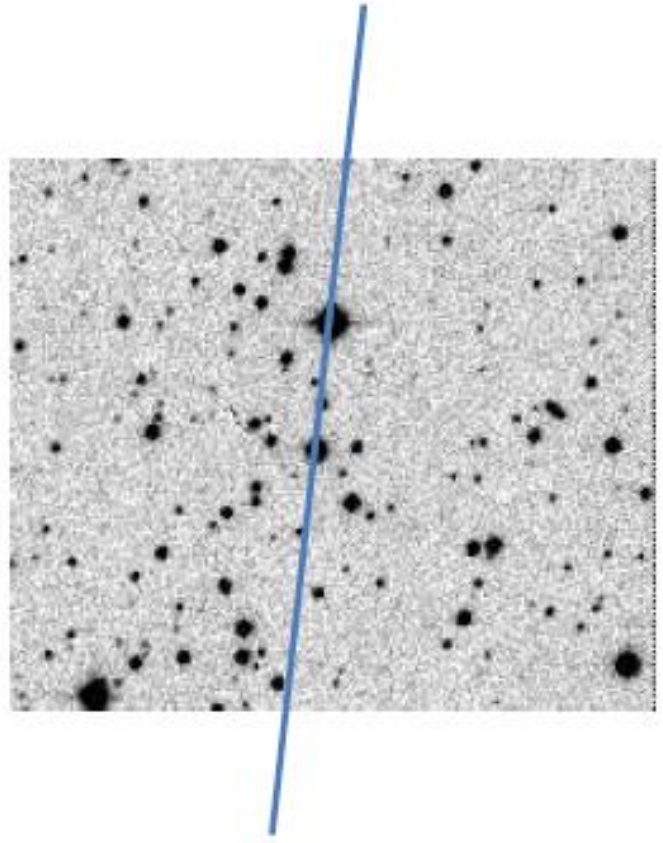
Flux calibration (1)

- This is a calibration of sensitivity and efficiency, and is carried out in the same way as the photometric calibration, by observing a number of standard stars whose flux as a function of wavelength is accurately known, at a variety of airmass values.



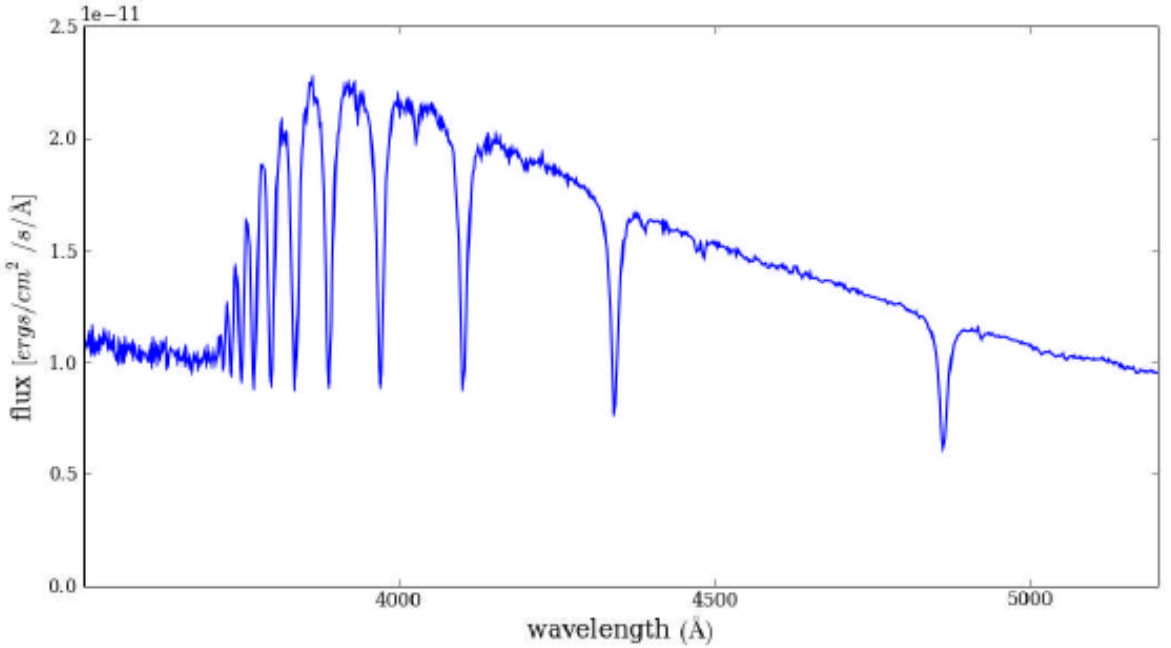
Flux calibration (2)

- Flux calibration is subject to variable slit-losses since target observations are observed through a narrow slit
- A second reference star can be aligned to fall along the slit such that both target and reference star spectra can be extracted
- A differential flux correction can then be made by comparing the narrow slit observations with a wide slit observation of the reference star
- This requires the slit-angle to be fixed, and thus not be at the parallactic angle!



“Final” spectrum

Flux calibration of the science spectrum



- Velocity rest-frame: heliocentric frame
- Extinction/telluric correction

Now the fun can begin: velocities, abundances etc.

