Interstellar Absorption Lines

OPTICAL & UV ABSORPTION LINES
DIFFUSE INTERSTELLAR BANDS
RADIATIVE TRANSFER IN IS LINES
COLUMN DENSITY
CURVE OF GROWTH
21 CM HYDROGEN LINE

Interstellar Absorption Lines

Optical & UV Absorption Lines:

- Familiar solar optical doublets, e.g., Fraunhofer Ca II K and Na I D provided early evidence for a pervasive ISM.

- These are "resonance" transitions with an electron going from the ground state to the next energy level (from an *s* to a *p* orbital).
- Similar transitions occur across UV and optical wavelengths. Some important examples are:

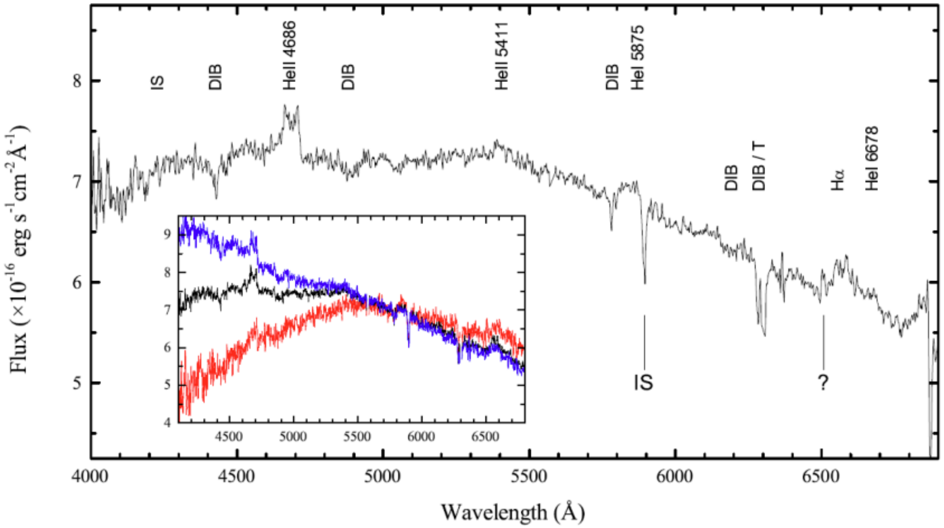
H I	1216 Å
C IV	1548, 1551 Å
Na I (Na D)	5890, 5896 Å
Mg II	2796, 2803 Å
K I	7665, 7645 Å
Ca II	3934, 3968 Å

Diffuse Interstellar Bands (DIBs)

DIBs are the most mysterious of the UV/Visible absorption components of the ISM. Since their discovery by Merrill in 1938, more than 200 DIBs have been identified in stellar spectra, with the strongest appearing at λ4430 Å.

They have not been identified conclusively with any atomic or molecular species (neutral or ionized). They are characterized by being extremely broad (by the standards of interstellar absorption lines).

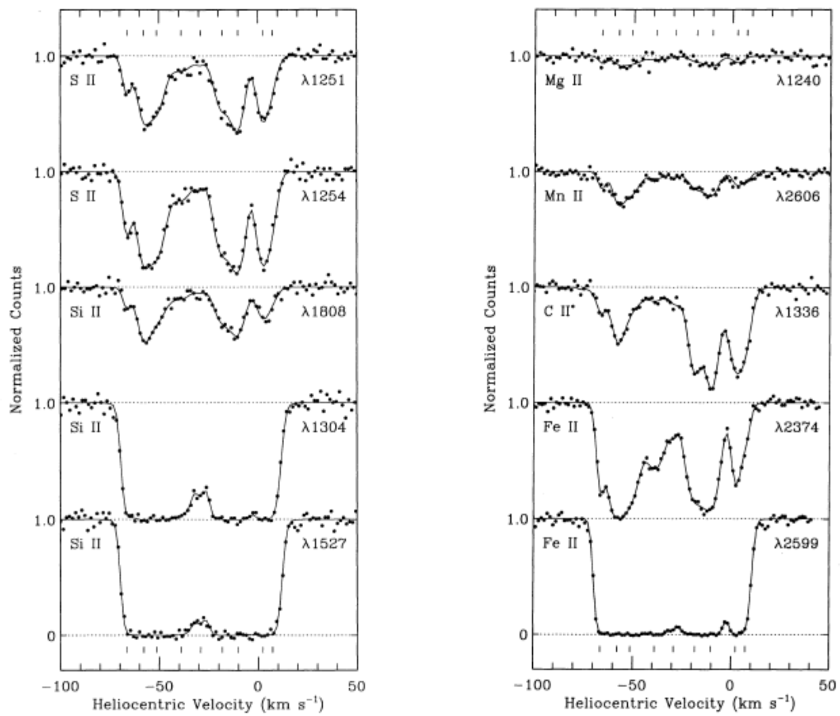
Some ideas are exotic molecular bands, transition from stuff on dust grain surfaces, exotica like ionized Fullerenes (3-D aromatic C molecules shaped like geodesic spheres), but none have produced consistent predictions of wavelengths.



A spectrum of the distant Black Hole binary Swift J1753.5-0127, showing a number of DIBs [from Neustroev et al. 2014]

Observations of IS Absorption Lines

At high spectral resolutions ($R=\lambda/\Delta\lambda \gtrsim 10^4$) interstellar absorption lines resolve into narrow absorption lines that are Doppler shifted relative to each other.



Interstellar absorption lines towards the halo star HD93521.
These spectra reveal velocities spanning ~ 90 km/s and show multiple velocity components and effects of line saturation in different species.
[From Spitzer & Fitzpatrick 1993]

Our ultimate aim is to deduce from measurements of these absorption lines the column densities and velocity dispersions of the absorbing atoms and ions.

Radiative Transfer in IS lines

We have already learned many things in this course which can also be applied to studying the ISM:

- We defined the specific intensity I_λ, emission (j_λ and ε_λ) and absorption coefficients (κ_λ and α_λ), optical depth $d\tau_\lambda = \kappa_\lambda \rho ds = \sigma_\lambda n ds = \alpha_\lambda ds$, the source function $S_\lambda = j_\lambda/\kappa_\lambda$.

- For the ISM, we can use the parallel-ray equation of radiative transfer:

$$\frac{dI_\lambda}{d\tau_\lambda} = -I_\lambda + S_\lambda$$

- We formally solved it (slides 147-148), assuming that S_λ is constant along the path:

$$I_\lambda = S_\lambda(1 - e^{-\tau_\lambda}) + I_{\lambda 0} e^{-\tau_\lambda}$$

- The more accurate (constant S_λ along the path is a rude assumption), formal solution is

$$I_\lambda(\tau_\lambda) = \int_0^{\tau_\lambda} S_\lambda(t_\lambda) e^{-(\tau_\lambda - t_\lambda)}\, dt_\lambda + I_{\lambda 0} e^{-\tau_\lambda}$$

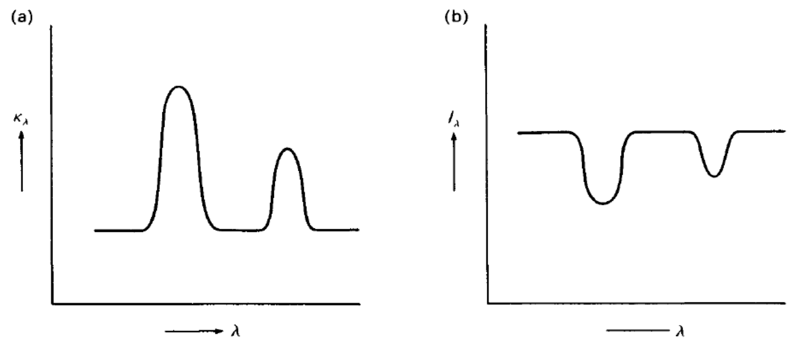
Based on slide I.151

$I_\lambda = S_\lambda(1 - e^{-\tau_\lambda}) + I_{\lambda 0} e^{-\tau_\lambda}$

Imagine $I_{\lambda o} \neq 0$ (background source)

If $I_{\lambda o} > S_\lambda$, so there is something subtracted from the original intensity which is proportional to the optical depth – we see absorption lines on the continuum intensity I_λ.

EXAMPLE: stellar photospheres or **interstellar absorption lines**



Opacity κ versus λ ➔ Intensity versus λ

An extreme case: Optically thin case ($\tau_\lambda \ll 1$):

$$I_\lambda = I_{\lambda 0}(1 - \tau_\lambda) + \tau_\lambda S_\lambda = I_{\lambda 0} + \tau_\lambda (S_\lambda - I_{\lambda 0})$$

Column Density

- Consider an arbitrary structure of gas – a "cloud" – with number density of atoms and ions which is a function of position within the cloud: $n(cm^{-3}) = n(x, y, z)$.
- A line-of- sight to a background source of light (e.g., a star) will probe a finite distance through the cloud for a length L.
- The column density N is defined as the integral of the number density along the line of sight:

$$N = \int_0^L n(l)dl$$

- The quantity N can be deduced from the strengths and profiles of IS absorption lines.

- If column densities can be measured for different ionization stages of the same ion, e.g. Si, Si^+, Si^{++}, Si^{3+} from absorption lines of Si i, Si ii, Si iii, Si iv, the ionization conditions of the gas can be inferred.
- In some physical environments, for example clouds where H is predominantly neutral, the ionization structure of the gas can be simple, with most of the atoms/ions of a given element being concentrated in a dominant ionization stage.

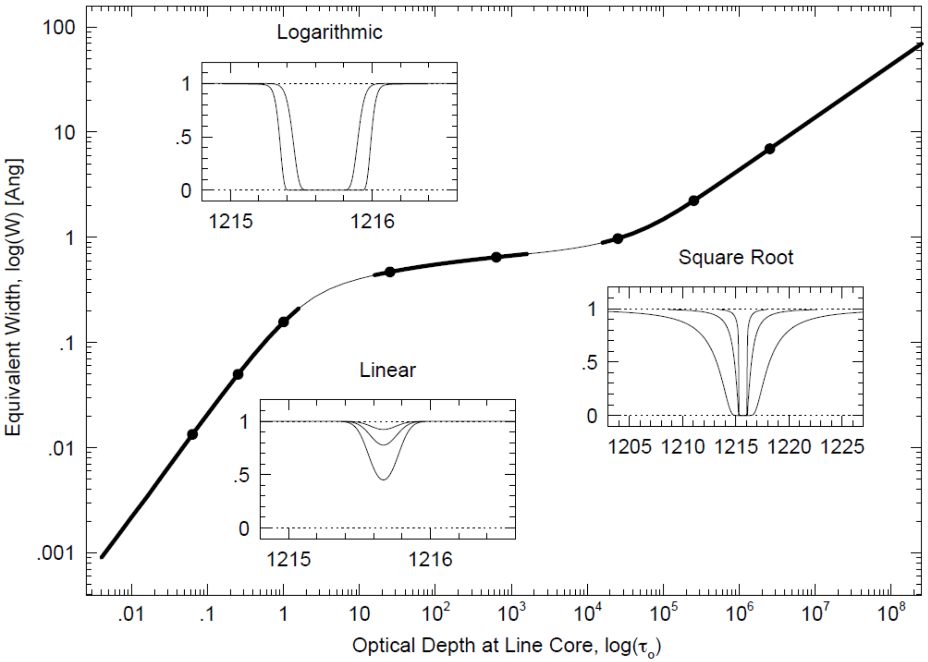
- In this circumstances, if absorption lines from different elements are available, the chemical composition of the absorber can be investigated.

The Equivalent Width Curve of Growth

A traditional method of analyzing absorption line data is via the Curve of Growth (Lecture 24).

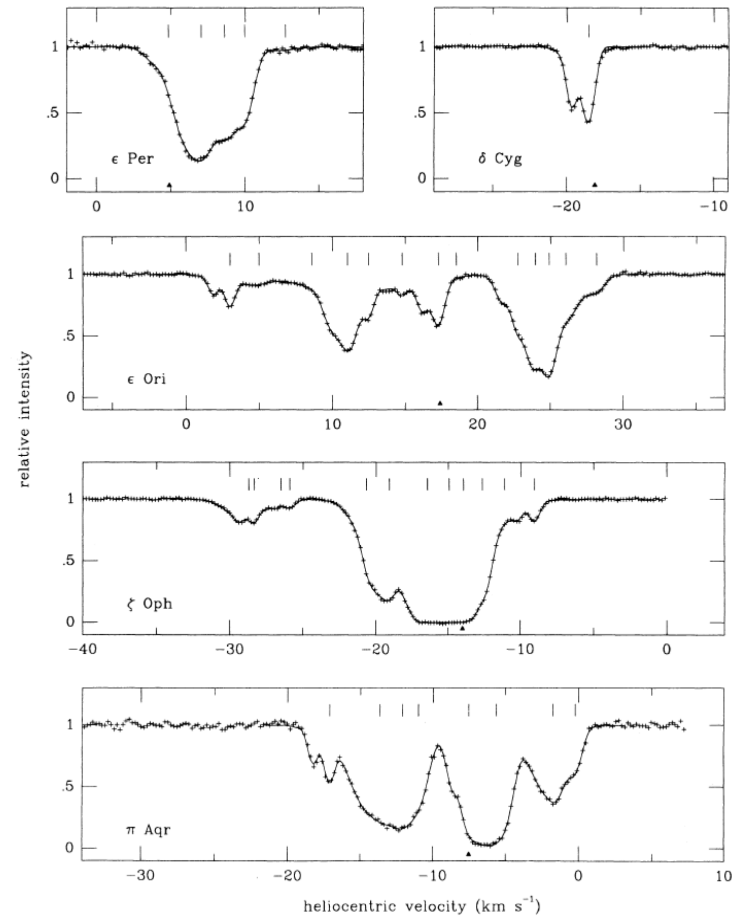
The precise functional dependence of the equivalent width W on N is sensitive to the optical depth at the line core, τ_0 and the line profile. This behaviour defines three distinct portions of the Curve of Growth:

- The linear part, where $\tau_0 < 1$ and $W \propto N$. The absorption line is optically thin and W is a sensitive measure of N, irrespectively of the values of the Doppler width $\Delta\nu_D$ or the damping widths $\gamma/2$.
- The flat (logarithmic), or plateau, where $10 \lesssim \tau_0 \lesssim 1000$ and $W \propto (\ln N)^{1/2}$. The absorption line is optically thick and W is not a good measure of N, but is sensitive to the Doppler width $\Delta\nu_D$.
- The damping, or square root part, where $\tau_0 \gtrsim 10^4$ and $W \propto N^{1/2}$. In this regime, the optical depth in the damping wings provides an accurate estimate of N.



IS absorption line profiles

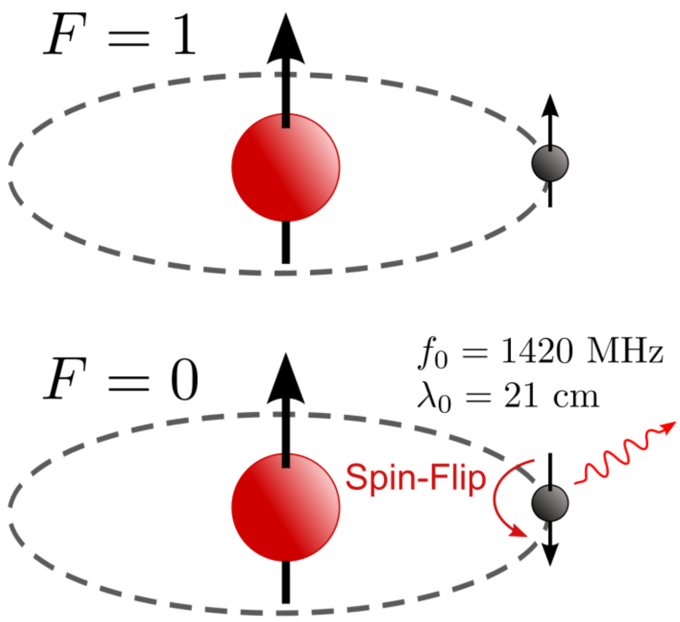
- In practice, curve of growth methods are quite powerful, as they can relate the product N to a direct observable, W, that is relatively insensitive to the choice of spectral resolution.
- In principle, two different spectrometers working at very different resolutions and on different telescopes with different detectors should be able to measure the same equivalent widths to within the irreducible measurement uncertainties.
- Real interstellar absorption lines are often highly structured with a mixture of both saturated and unsaturated components because the line of sight to a particular star will often intersect a number of interstellar clouds with a wide range of column densities.



These Na D profiles show a mix of linear (δ Cyg), flat (ε Per & ε Ori), and square-root (ζ Oph) absorption lines.

21 cm hydrogen line (1)

- One can study ISM using absorption lines, but we need a bright background ”lamp”! Which we don’t have for all directions!!
- To map the distribution of interstellar atomic gas in any direction we wish to study emission lines.
- Most of our knowledge of the distribution of neutral atomic hydrogen in the ISM of the Milky Way and other galaxies come from observations of the strong 21-cm line (1420 MHz).
- This line arises from transitions between the hyperfine structure levels in the ground state of Hydrogen, and is seen in both emission and absorption.



21 cm hydrogen line corresponds to a flip of the electron spin relative to the spin of the proton:
Upper state: electron and proton spins are parallel, statistical weight g_u=3.
Lower state: electron and proton antiparallel, g_l=1.

21 cm hydrogen line (2)

Upper state: electron and proton spins are parallel, statistical weight g_u=3.

Lower state: electron and proton are antiparallel, g_l=1.

A_{lu} = transition probability = 2.9 × 10^{-15} sec^{-1} →

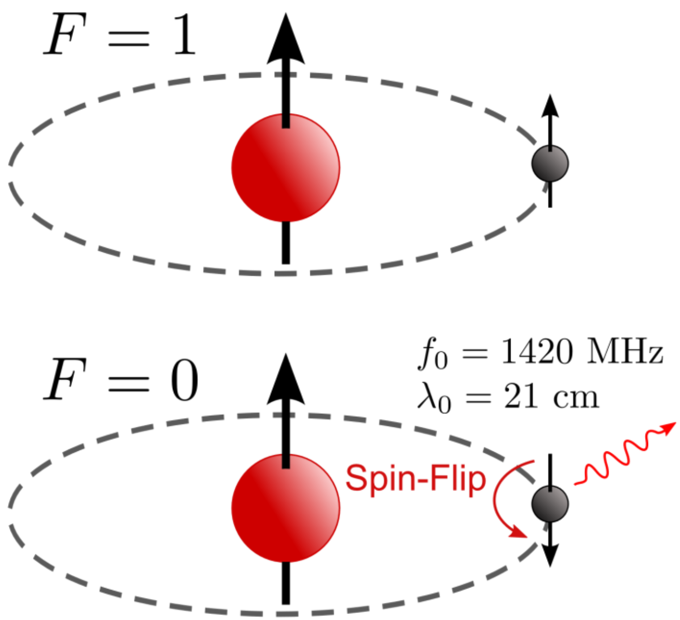
Lifetime of upper level = 11 million years!

Thus, for $n_H \approx 1\ cm^{-3}$ collisions dominate - levels are populated according to the Boltzman equation:

$$\frac{N_u}{N_l} = \frac{g_u}{g_l} e^{-(E_u - E_l)/kT} \approx \frac{g_u}{g_l} \approx 3$$

Since the energy difference between levels is very small!!

Thus, the populations of the levels are essentially independent of temperature in the ISM.



The transition energy in units of Kelvin is 0.07 K, which is much lower than in most astrophysically interesting conditions.

ISM: Ionized regions

SOURCE OF IONIZATION
IONIZATION AND RECOMBINATION
STRÖMGREN SPHERES

The non-equilibrium ISM

The ISM is generally not in thermodynamic equilibrium due to its low density and consequent low collision rates and low optical depths. Collisions between particles may result in radiative decay rather than redistributing internal energies. Photons can escape the system, or enter from external sources. Whereas the distribution of velocities remains generally Maxwellian, and described by a kinetic temperature on scales greater than a mean free path, the distribution of energy levels may be significantly different from the Boltzmann distribution. This can be formulated either as departure coefficients,

$$b_i = \frac{n_{i,actual}}{n_{i,LTE}}$$

or, more commonly, by defining an excitation temperature, T_{ex}, such that

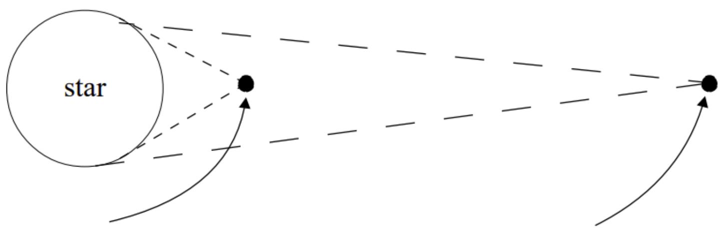
$$\frac{n_u}{n_l} = \frac{g_u}{g_l} e^{-(E_u - E_l)/kT_{ex}}$$

Note that although T_{ex} has units of K, it is not a physical temperature and may not be equal to the kinetic temperature. It is a function of the energy level and parameterizes how far the distribution of states is from Boltzmann.

Radiation field

Interstellar radiation field is characterized by a huge discrepancy between the frequency-integrated radiation density and spectral composition.

Compare photons close to and far from an extended source like a star:



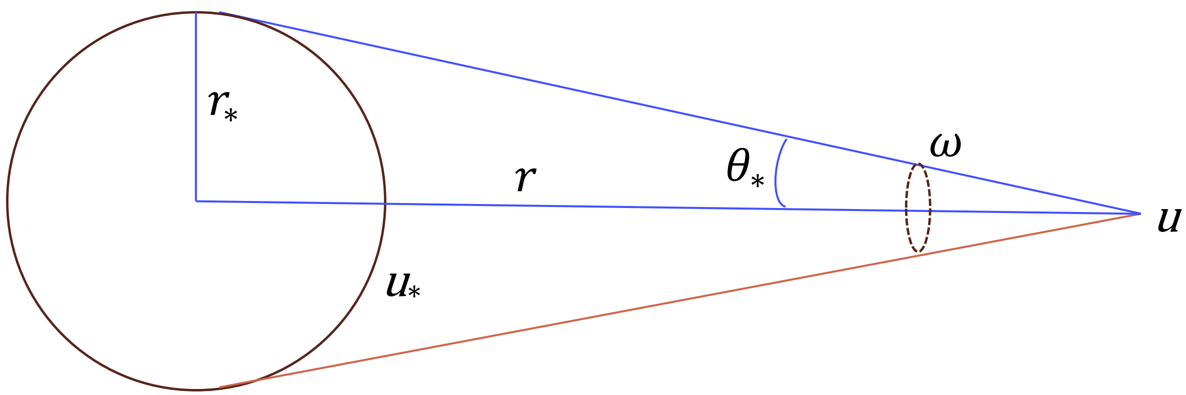
Photons here are crowded in physical space (high density) but spread out in solid angle (direction)

Photons here are diffuse in physical space (low density) but highly collimated in angle (high density).

It is sometimes convenient to approximate the density of the background starlight as a "diluted blackbody" of "dilution factor" W and "color temperature" T_c.

Dilution factor (1)

Compare photons close to and far from an extended source like a star:



The dilution factor W is defined to be the ratio of the actual energy density u to the energy density of (undiluted) blackbody radiation of temperature T_c.

ω is the solid angle, then $u = \frac{\omega}{4\pi} u_* = u_* W$, where W is the **dilution factor**.

$$W = \frac{\omega}{4\pi} = \frac{1}{4\pi}\int_0^{2\pi} d\varphi \int_0^{\theta_*} \sin\theta \, d\theta = \frac{1}{2}(1-\cos\theta_*) = \frac{1}{2}\left(1-\sqrt{1-\frac{r_*^2}{r^2}}\right) \approx \frac{1}{2}\frac{r_*^2}{r^2}$$

$r \gg r_*$

Dilution factor (2)

The dilution factor $W \approx \frac{1}{2}\frac{r_*^2}{r^2}$

Planetary Nebula? $r_* \approx 10^9$ cm, $r \approx 10^{16}$ cm → W~10^{-14}

Slide I.156: the energy contained in volume element per wavelength interval is $u_{*,\lambda} = \frac{4\pi}{c} J_\lambda \left[\frac{erg}{cm^3 Å}\right]$

Assume Blackbody, then $u_{*,\lambda} = \frac{4\pi}{c} B_\lambda(T_*)$, where $T_* = T_{\mathrm{eff}}$ of a star.

The integrated radiation density $u_* = aT_*^4$ for the star and $u = \int_0^\infty u_\lambda \, d\lambda = aT_d^4$ far from the star.

Then $T_d^4 = WT_*^4$, or

$$T_d = W^{1/4} \, T_*$$

T_d is the temperature corresponded to the integrated energy density of diluted blackbody radiation of temperature T_*. For W~10^{-14}, T_d=a few K.

But a spectrum corresponds to $T_C = T_*$! The ISM must work as a transformer, decreasing T_C.

Indeed, Rosseland's theorem dictates that

in low-density regions, short-wavelength radiation is transformed into long-wavelength radiation.

This immediately explains a PNe observable, that the nebula emits much more energy in the optical than does the central star. This is because UV radiation is being processed into optical photons.

The non-equilibrium ISM (2)

The interstellar radiation field is very dilute but populated by discrete energetic sources (e.g. OB stars) and can contain multiple spectral features from the gas. It is therefore generally very different from the Planck function which we can parameterize in terms of a brightness temperature.

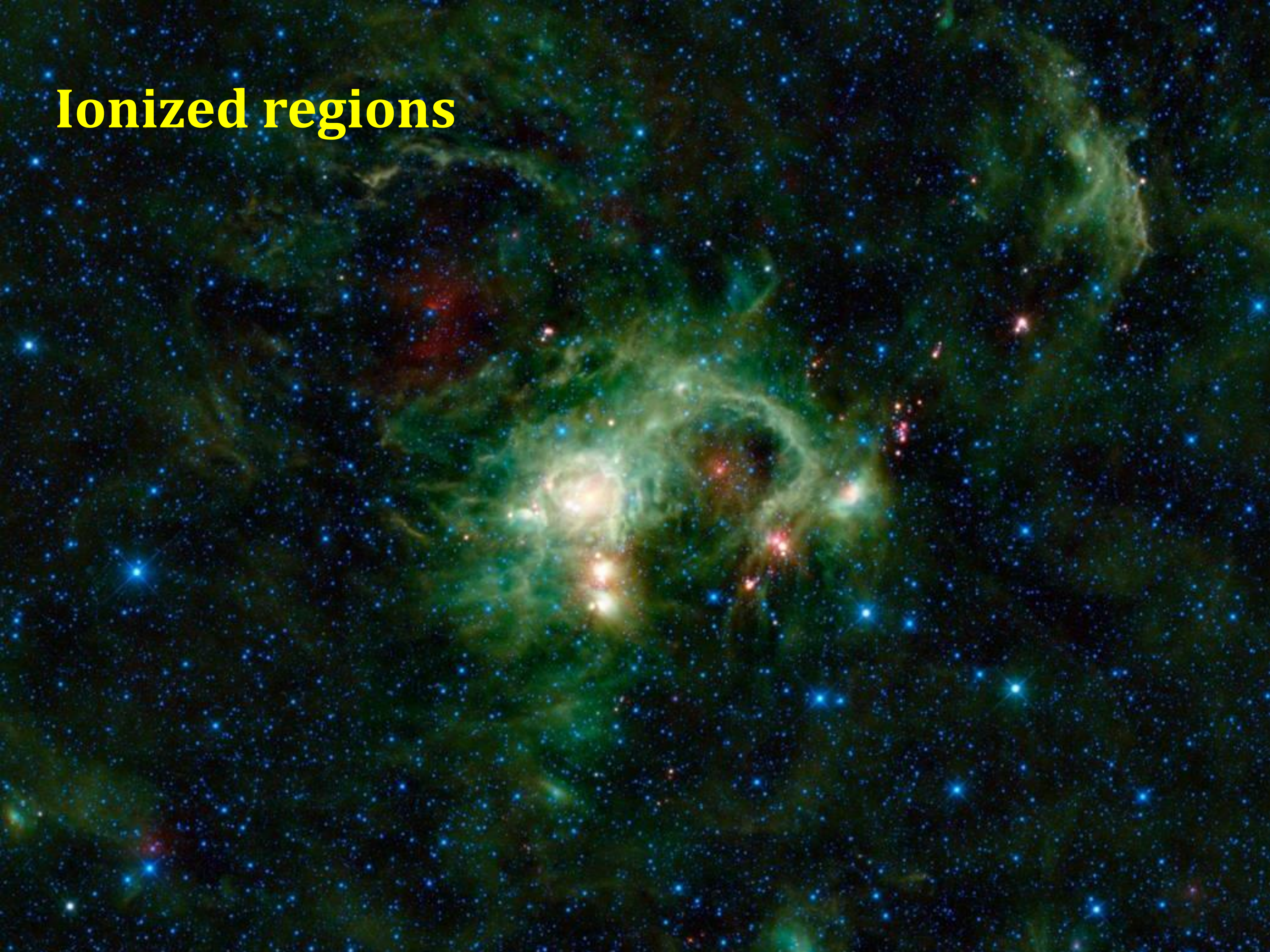
Despite these complications, we can still look for a statistical equilibrium solution to the distribution of energy levels,

$$\frac{dn_i}{dt} = \sum_{j\neq i}^{N} n_j P_{ji} - n_i \sum_{j\neq i}^{N} P_{ij} = \sum_{j\neq i}^{N} n_j (R_{ji} + C_{ji}) - n_i \sum_{j\neq i}^{N} (R_{ij} + C_{ij}) = 0$$

where P_{ji} is the (radiative R_{ji} plus collisional C_{ji}) rate from level j to i. This matching of forward and reverse rates is also known as the principle of detailed balance.

Remember? See lecture 25.

Ionized regions



Source of ionization?

H II regions are zones of ionized atomic hydrogen. They are often associated with nebulae, and require gas that contains a continuous source of ionizing radiation.

Ionization potential of hydrogen $h\nu_0 = 13.6$ eV

($\nu_0 = 3.29 \times 10^{15}$ Hz, $\lambda_0 = 912$ Å)

Thus, for ionization we need $h\nu > 13.6$ eV
(it is $\gg kT$ in neutral ISM => collisions unimportant)

This could be a massive star or a white dwarf, but either way it must be very hot.

Are OB stars hot enough?

$T \approx 3\text{-}5 \times 10^4$ K => $E = kT \approx 3\text{-}5$ eV < 13.6 eV => cannot ionize H…?

$$B_\lambda(T) = \frac{2hc^2}{\lambda^5} \frac{1}{e^{\frac{hc}{\lambda kT}} - 1}$$ => can ionize H!

Yes, the OB stars are hot enough.

H II regions

Without that **hot** central source, the protons and electrons in the ISM will quickly recombine – even with the ionizing source, it can only ionize a region of some given volume before recombinations will be happening as quickly as ionizations.
The result will be a bubble of ionized gas, termed a **Strömgren Sphere**.
Such H II regions are easily observable via the strong emission lines resulting from recombination. Thus, to add further to the nomenclature, they are also sometimes known as emission-line nebulae (remember the spectrum of the Cat's Eye Nebula?)

Our goal now is to understand the size of the ionized bubble and its detailed ionization structure. In this effort, we define the ionization fraction

$$f \equiv \frac{n_{H^+}(r)}{n_H(r)}$$

For a fully neutral, atomic ISM $f = 0$, while full ionization implies $f = 1$.
As hinted at in the preceding argument, to maintain a constant f we will want to make use of ionization equilibrium, where

# of recombinations / sec = # of ionizations / sec

Ionization and Recombination

- Consider pure hydrogen gas
- Reaction: H+ hν ↔ p + e
- Ionization produces an electron of energy $E_{\mathrm{kin,e}} = h\nu - \chi_{\mathrm{ion}}$
  where χ_{ion} is the ionization potential (one can assume all atoms to be in the ground state)
- Recombination to level i gives a photon

$$h\nu = E_{\mathrm{kin,e}} + \chi_{\mathrm{ion}} / i^2 \approx \chi_1 \, (1/i^2 + 0.07 \; T_{\mathrm{e}}/10^4 \; \mathrm{K})$$

where $\chi_{\mathrm{ion}} \approx kT_{\mathrm{e}}$ is kinetic energy of electron, i – is the main quantum number, plus a cascade of line (Balmer+Lyman etc.) photons.
Photons recombining to the 1st level can still ionize hydrogen.

- Equilibrium: # of recombinations = # of ionizations

Energy redistribution and temperature

- We have e, p, and neutral H.

- Why should we assume that all of them are described by the same temperature $T_e = T_p = T_H$?

- Why should they have Maxwellian distribution?
- The distribution of injected electrons depends on incident photon spectrum.
- Electrons thermalize rapidly by collisions.
- Electrons and protons exchange energy much slower (since the masses are very different).
- Protons and neutrals exchange energy even slower $\tau_{ee} \ll \tau_{ep} \ll \tau_{pH}$ but still $\ll \tau_{dyn}$ $\Rightarrow$
- Electrons lose energy in inelastic collisions (by radiation). Energy is taken from the gas energy as a whole. We associate gas temperature with T_e.

HII regions: recombination (case A)

- Recombination rate proportional to $n_e n_p$. Depends on T_e, since the recombination probability and the flux of electrons both depend on T_e
- e^- can be captured in any level
- Recombination rate to level i is

$$N_{rec,I} = n_e n_p \, \alpha_{i,rec}(T_e) \text{ cm}^{-3} \text{ s}^{-1}$$

where $\alpha_{i,rec}$ is the recombination coefficient.

- Total "case A" recombination coefficient

$$\alpha_A = \sum_{n=2}^{\infty} \sum_{l=0}^{n-1} \alpha_{nl}(T) \approx 4.2 \times 10^{-13} \, (T_e/10^4)^{3/4} \text{ cm}^3 \text{s}^{-1}$$

- Thus,

$$N_{rec} = n^2 \alpha_A (T_e) \text{ cm}^{-3} \text{ s}^{-1}$$

However

HII regions: recombination (case B)

- Photon produced by recombination to level 1 still can ionize gas. These photons produce "**diffuse**" radiation field. Assuming that these photons are absorbed nearby ("on-the-spot" approximation), one can just neglect these recombinations in the total recombination rate:

$$N_{\rm rec} = \sum_{i=2} N_{{\rm rec},i} = n_{\rm e}^2\, \alpha_{\rm B}\,(T_{\rm e})\ {\rm cm^{-3}\ s^{-1}}$$

where "case B" recombination coefficient

$$\alpha_B = \alpha_A - \alpha_1$$

$$\alpha_{\rm B}\,(T_{\rm e}) \approx 2.6 \times 10^{-13}\,(T_{\rm e}\,/\,10^4\ {\rm K})^{-3/4}\ {\rm cm^3\ s^{-1}}$$

Steady State in HII Regions

$H^+ + e^- \rightarrow H + h\nu$

$H + h\nu \rightarrow H^+ + e^-$

Case A recombination: gas optically thin to radiation produced after recombination

Case B recombination: gas optically thick to radiation just above 13.6 eV;
photons produced in recombination are absorbed for photoionization of another atom

Photoionization/ Recombination

- Transitions down are fast : $A \sim 10^8$ s^{-1}
  All atoms are on ground state (absorption in Balmer lines is negligible).

- Total ionization rate at distance r from the star is:

$$N_{\rm ion} = n_{\rm H}\,\Gamma_i \approx n_{\rm H}\,\sigma_0\,N_* / (4\pi r^2)\ \text{cm}^{-3}\ \text{s}^{-1}$$

where N_* is the number of stellar ionizing photons (with $h\nu$ >13.6 eV) per second,
$\sigma_0 = 6.3 \times 10^{-18}$ cm^2 is the ionization cross-section at $\nu = \nu_i$,
and Γ_i is the number of ionizations per atom per second from level i

$$\Gamma_i = \int_{\nu_i}^{\infty} \frac{F_\nu}{h\nu}\sigma_i(\nu)\,d\nu, \qquad \text{where} \quad F_\nu = L_\nu / 4\pi r^2$$

σ_i is the ionization cross-section: $\sigma_i = 6.3 \times 10^{-18}\,(\nu_i / \nu)^{3.5}$ cm^2

HII regions: ionization

- Let introduce $n_e = x\, n$

  $n = n_p + n_H = n_e + n_H$

- Then $n_H = (1-x)\, n$

$$N_{rec} = N_{ion} \Rightarrow x^2 n^2 \alpha'_{rec} (T_e) = (1-x)\, n\, \sigma_0\, N_* / (4\pi r^2)$$

- For O6.5 star: $N_* = 10^{49}$ s^{-1}, r = 1 pc, $n = 10^2$ cm^{-3}, $T_e = 10^4$ K, we get $(1-x) = 3 \times 10^{-5}$, i.e. $x \sim 1$ and hydrogen is almost fully ionized.

Strömgren Spheres

- Two components to local ionizing flux near a star:
  - direct ionizing flux
  - diffuse flux from recombinations to ground state

- Calculate radius of ionized sphere in uniform density:
  balance total number of ionizations (=number of ionizing photons the star produces) to the total number of recombinations to levels above ground $\alpha^{(2)}=\alpha'$

$$\frac{4\pi}{3} r_S^3 x^2 n^2 \alpha_B = N_* \quad \Longrightarrow \quad \boldsymbol{r_S = \left(\frac{3N_*}{4\pi n^2 \alpha_B}\right)^{1/3}}$$

- Strömgren radius for an O star: r_S=70 pc n$^{-2/3}$; n=10^2 cm^{-3}, $r_S \sim$ 3 pc

→ OB stars have an enormous impact on the ISM

Edges of HII regions

- n_H ↑, absorption ↑, ionizing photon flux ↓, n_H ↑
- The optical depth $\tau = \int n\,\sigma\,dl$ for ionization is huge.
- At Lyman edge, $\sigma = 6.3 \times 10^{-18}$ cm^2, so if $n = 1$ cm^{-3}, the mean free path Δr (where τ=1) is

  $\Delta r = \tau / n\,\sigma = 1.5 \times 10^{17}$ cm / n = 0.05 pc / n

$$\Delta r / r_S = 10^{-3}\, n^{-1/3} \ll 1$$

- Observed H II regions limited:
  - ionization bounded: all photons are used up for ionization, interstellar cloud has a larger extent than the nebula
  - density bounded: all atoms ionized, there are still photons left
- Ionization bounded H II regions have sharp edges

Edges of HII regions