Interstellar Astrophysics
Summary notes: Part 2

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The main reference source for this section of the course is Chapter 5 in the Dyson and Williams (The Physics of the Interstellar Medium) book. Beware: this book is very technical!!

Note: Figure numbers refer either to the chapter and figure number (N-n) in Freedman and Kauffman, or to the Dyson and Williams textbook.

2 Photoionised nebulae: H II regions

H II regions (like the Orion nebula) are regions of ISM gas in which an OB star (or OB stars) are embedded. For simplicity we shall start by assuming an idealized nebula which is spherical and at its centre is a single O or B star. The radiation from the central OB star photoionises the nebula gas, and it is this process and its inverse (recombination) which provides the physical basis for much of the nebula's physical properties.

2.1 The photoionisation process

Photons from the central OB star, which has a high effective temperature ($\gtrsim$20 000 K at B0 and up to $\sim$50 000 K for an O5 star) can have a sufficiently high energy such that when they are absorbed by atoms in the gas they cause the removal of a bound electron from the atom. This is termed ionisation.

We will consider hydrogen since it is the most abundant element in the gas: our basic picture of the atomic structure of H is that it consists of a proton and a single electron, which can occupy any bound energy state, denoted by a 'quantum number' n =1,2,3,4, etc. The $n = 1$ state is known as the ground state and has zero energy ($E = 0$; see Fig. 5-25). As we go to very large n-values the energy levels become very close together and there comes a point where the electron is no longer in a bound state but can occupy a range of possible continuum states where it has a certain free kinetic energy.

The removal of a bound electron from an atom to form an ion and a free electron is called ionisation (in the case of hydrogen, this ion is a single proton). If the energy supplied to ionise the atom comes from an incoming photon, the process is called photoionisation.

Using hydrogen as an example (Fig. 5-25) we see that the energy required to take the electron from the ground state ($n = 1$, $E = 0$) to the continuum is 13.6 eV (electron volts). This value is known as the ionisation energy for hydrogen, I_{H}. Thus to ionise H from the ground state we need photons with energy >13.6 eV, or in wavelength terms photons with wavelengths less than 912 Å*. These photons are called Lyman continuum photons. Only hot (OB) stars have radiation fields with significant numbers of photons at these short (UV) wavelengths. Thus only OB stars can photoionise the hydrogen gas in the surrounding nebulae (Fig. 1).

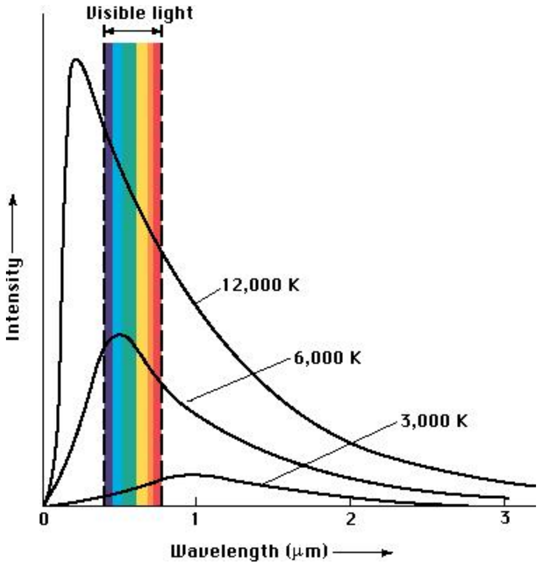

Figure 1: Hotter stars produce more photons in the UV range

NB. the low density of the gas in H II regions means that the excitation of the H-atoms is very low, so that nearly ALL the H-atoms will have their electrons in the ground state ($n = 1$). Thus photoionisation need ONLY be considered from the $n = 1$ state.

The strong (Coulomb) attraction between the negatively charged electron (e^-) and the positively charged proton (p^+) produces the inverse process of recombination, so we have a reversible reaction:

$$\mathrm{H} + \mathrm{h}\nu \;\rightleftharpoons\; \mathrm{p}^+ + \mathrm{e}^- \tag{1}$$

NB. the notation "hν" is often used to represent a photon.

In recombination the energy of the photon which is produced depends on two things: (a) the kinetic energy (KE) of the recombining electron (in its continuum state) and (b) the binding energy of the bound-level into which the recombination occurs.

If the recombination occurs directly to some upper level n', the electron will try and go to the ground state by jumping down to lower n-levels, and each $n' \rightarrow n''$ jump produces a photon at a specific energy or wavelength which corresponds to some particular spectral

*Photon energy is described using the formula: $E = h\,\nu$, where h is the Planck constant and ν is the frequency of the light (photon). This can be related to the wavelength using $c = \nu\,\lambda$, where λ is the wavelength, thus $E = h\,c\,/\,\lambda$. In this way we can relate energy in eV to wavelength.

Table 1: Hydrogen recombination lines (wavelengths in nm)

	Lyman $n' \to n = 1$	Balmer $n' \to n = 2$	Paschen $n' \to n = 3$
α	122	656.3	1875
β	103	486.1	1282
γ	97	434.0	1094
δ	95	410.2	
ϵ	94		
	(UV)	(optical)	(IR)

line in hydrogen. Transitions from $n' \to n = 2$ give rise to the Balmer series of lines seen at optical wavelengths (Hα is the 3→2 transition and occurs at 6563 Å, Hβ is the 4→2 transition and occurs at 4861 Å). Transitions from $n' \to n = 3$ give the Paschen series in the near IR, whilst those from $n' \to n = 1$ give the Lyman series in the far-UV (see Table 2.1). This is the main process by which the hydrogen emission lines are produced in H II region spectra – and this produces the recombination line spectrum of the nebula.

Equilibrium occurs when there is a balance between the forward and backward rates of the above equation – this is called the condition of ionisation balance.

2.1.1 Thermalisation of gas in a pure hydrogen nebula

As Eqn. 1 shows, the nebula gas contains three types of particles: neutral atoms (H), protons (p^+) and electrons (e^-).

Photoionisation feeds energy continuously into the nebula via the kinetic energy of the ejected (photo)electrons. The energy of the electrons is determined by the energy distribution of the ionising photons from the ionising OB star. This is determined by its effective temperature, $T_{\rm eff}$ (see Fig. 1).

Collisions between the particles take place which thermalises the particle velocity distributions and transfers energy from one type of particle species to another, and this occurs on such short timescales as to be regarded as instantaneous compared to the timescale for successive photionisation/recombination.

After thermalization, all the particles in the photoionised nebula have a 'Maxwellian Distribution' of velocity (and hence KE), and thus we can assign a single "kinetic temperature", $T_{\rm k}$, of the gas particles which is the same for all three types of particle.

The speeds at which the particles interact and share energy follows the order: (i) hot, photoelectrons share their energy with other electrons, (ii) electrons transfer energy to protons via collisions (needs a longer time since the proton mass is > electron mass), (iii) protons share their energy by collisions with neutral atoms.

2.1.2 Ionisation balance in (pure H) H II regions

We need to be able to calculate the forward and background rates to get the hydrogen ionisation ratio.

The photoionisation rate is given by:

$$\mathrm{N_{PI}} = a\ n_{\mathrm{HI}}\ J \qquad [\mathrm{m^{-3}\ s^{-1}}] \tag{2}$$

– n_{HI} is the number density of neutral H (H I)
– J is the number of ionising photons crossing unit area per unit time at some point in the nebula,
– a is the 'photoionisation cross-section' for H (from level $n = 1$). To a first approximation we can assume $a = 6.8 \times 10^{-22}$ m^2.

The cross-section of an interaction can be thought of as the 'effective area' that the interacting particle sees. The term 'cross-section' is just a way to express the probability (likelihood) of interaction. See Fig. 2.

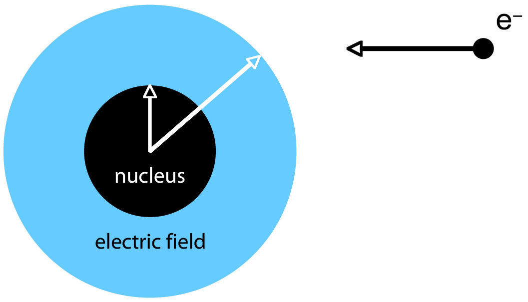

Figure 2: Interaction cross-section. Geometric vs. effective area.

The recombination rate into level n of an H-atom is:

$$\mathrm{N_R} = n_{\mathrm{e}}\ n_{\mathrm{p}}\ \alpha_n = (n_{\mathrm{e}})^2\ \alpha_n(T_{\mathrm{e}}) \qquad [\mathrm{m^{-3}\ s^{-1}}] \tag{3}$$

– n_{e} is the electron density
– n_{p} is the proton density†
– α_n = recombination coefficient of level n, and is the rate of ion–electron recombination per unit volume (units cm^3 s^{-1}). It varies as a function of T_{e} (the electron temperature of the gas). Note: here the interaction cross-section is already incorporated in α_n.

To get the total rate of recombination we need to sum Eqn. 3 over all levels of H ($n = 0 \rightarrow \infty$)

Photoionisation is assumed to only occur from level $n = 1$ (due to the low excitation of the gas) but recombination can occur to any n-value level. Thus any recombination directly to the $n = 1$ level is immediately followed by a photoionisation.

†$n_{\mathrm{e}}\ n_{\mathrm{p}} = (n_{\mathrm{e}})^2$ because in equilibrium, the number of electrons equals the number of protons

For this reason, astronomers make it easy for themselves when calculating equation 3, and ONLY take recombinations to level $n = 2$ and above. This is known as case B recombination. Under the assumption of case B, α_n becomes α_B, and a good approximation is: $\alpha_\mathrm{B} = 2 \times 10^{-16}\ T_\mathrm{e}^{-3/4}$.

Calculations often use the degree of ionisation (or ionisation fraction), X, of H which is given by the simple relation:

$$n_\mathrm{p} = n_\mathrm{e} = X\ n$$

where n is the number density of hydrogen nuclei (i.e. the sum of the proton and neutral atom number densities). Thus $n_\mathrm{HI} = n - Xn$, and note that X lies between 0 and 1.

The condition of ionisation equilibrium says that Eqn. 2 = Eqn. 3, i.e. the total photoionisation rate = total recombination rate.

This allows us to calculate X for any given value of J and for a specified gas density (and hence n_e) at some distance r from the star. If the star is emitting $S_\star$ ionising photons per second (i.e. λ below 912 Å), then

$$J = S_\star/(4\pi r^2) \qquad [\mathrm{m}^{-2}\ \mathrm{s}^{-1}]$$

If we take a typical O-star energy distribution for an O6V star (like those of the Trapezium stars in Orion), we find $S_\star \sim 10^{49}$ photons per sec, and adopting for example a gas density of $n = 10^8\ \mathrm{m}^{-3}$ and $r = 1$ parsec, gives $J \sim 8.4 \times 10^{14}\ \mathrm{m}^{-2}\ \mathrm{s}^{-1}$. This gives $X = 0.999999$, i.e. the H gas is almost completely ionised.

This can be done for a range of r-values (distances from the central star) to give X as a function of r – see Fig. 3.

2.2 Sizes of photoionised H II regions: Strömgren Spheres

A particular star cannot ionise an indefinite volume of ISM gas. The volume it can ionise depends on the volume at which the total recombination rate is exactly equal to the rate at which the star emits ionising photons ($S_\star$).

If this region has a radius R_S, the condition of ionisation balance says that:

$$S_\star = \frac{4\pi}{3}\ R_\mathrm{S}^3\ n^2\ \alpha_\mathrm{B}$$

For a given value of $S_\star$ and n (the nebula gas density) we can solve the equation easily to find R_S, which is called the Strömgren radius for the ionised gas around an OB star.

Typical values are: $S_\star = 10^{49}$ photons per sec, $n = 10^8\ \mathrm{m}^{-3}$, which gives $R_\mathrm{S} = 3$ parsecs.

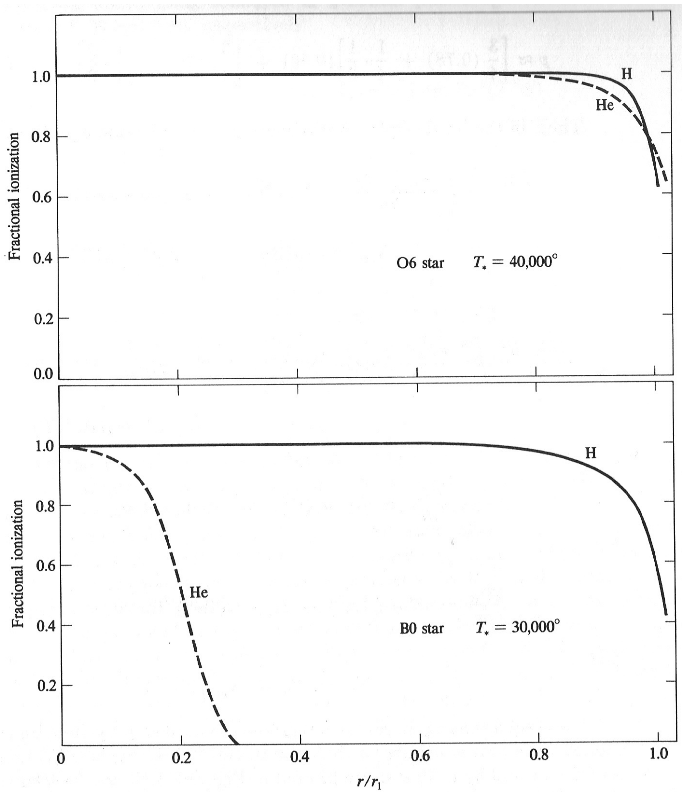

Figure 3: Ionisation structure of two homogeneous H + He model H II regions.

The radius of a Strömgren Sphere (the region of fully ionised H gas around an OB star) will increase if the gas density is lower, and will also increase the hotter the central OB star is – since it will emit more ionising photons. For low densities (say $n = 10^6$) and for an O6V star we can get $R_\mathrm{S} \sim 100$ parsecs.