Interstellar Astrophysics
Summary notes: Part 3

Dr. Paul M. Woods

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

2.3 Heating and cooling processes in photoionised nebulae

Photoionisation feeds energy into the nebula gas since it creates hot, fast photoelectrons. Without some energy loss mechanisms the gas temperature would rise indefinitely. Observational analysis of H II region spectra suggest electron temperatures of about 10 000 K, significantly lower that the effective temperatures of the ionising OB stars.

If we have a pure H nebula, energy is released in H recombination. The kinetic energy of the recombining electron is converted to photons, which may escape from the nebula and cool it. However, by comparing the heating and cooling rates from this process we find that higher nebula temperatures are expected.

On average each recombination removes an energy $(3/2)\,k\,T_{\rm e}$ from the gas, and the energy loss, or cooling rate, is:

$$L = \frac{3}{2}\,k\,T_{\rm e}\,{\rm N_R} \qquad [{\rm J\ m^{-3}\ s^{-1}}]$$

The energy input, the heating rate, is given by:

$$G = {\rm N_I}\,Q \qquad [{\rm J\ m^{-3}\ s^{-1}}]$$

where Q is the energy injected into the gas in each photoionisation. For star with an approximate Blackbody spectrum, one can show that $Q \sim (3/2)\,k\,T_{\rm eff}$. In equilibrium, $L = G$, and ${\rm N_R} = {\rm N_I}$, and we expect to find $T_{\rm e} \sim T_{\rm eff}$, i.e. a nebula gas electron temperature comparable to the star's effective temperature. For a typical OB star $T_{\rm eff} \sim 30\,000$–$60\,000$

K, which is much higher that values of $T_e \sim 10\,000$ K for nebulae derived from their spectra.

Inclusion of other possible H-cooling process like Balmer-line emission, and hydrogen free-free emission (which occurs mainly at radio frequencies) still gives too high temperatures,

This implies that we have neglected some important cooling process ...

2.3.1 Forbidden line cooling of H II regions (and PNe)

Real nebulae contain other elements than just hydrogen. We must, therefore, relax our assumption of a pure H nebula. A special emission line formation process in "metals" can give emission line radiation which can escape from the gas and cool it. This is called forbidden line emission.

Bound electrons in atoms can be excited/de-excited in two ways: by photons (resulting in absorption/emission) or by collisions (non-radiative process, meaning there is no way to 'see' it) between atoms or atoms and electrons.

The Balmer ($n \to 2$) and Lyman ($n \to 1$) transitions in hydrogen are excited/de-excited by the absorption/emission of photons, and give rise to what are called allowed transitions in quantum mechanics. The downward (de-excitation) transitions of electrons within the atom occur spontaneously with a very high probability (typical transition probabilities are $\sim 10^9$ per sec).

However, some energy levels in metals (and their ions) are split into 'sub-levels' or 'fine structure states'. Downward transitions amongst these low-lying energy levels can have a very low probability of occurring spontaneously (typically ∼10–100 per sec – or 1 per 0.01–0.1 secs). Due to this they are often also called metastable states. Photons emitted in this way result in what are called Forbidden Lines.

The formation process of these Forbidden Lines provides the cooling of the nebula gas in the following way:

(a) In the ISM, the atom density is too low for *atom–atom* collisions to occur. However *atom–electron* collisions do occur. Bound electrons in the lower levels of e.g. O^+ and O^{++} are excited to higher levels by collisions with free electrons in the gas removing some of the kinetic energy of the colliding electrons. These free electrons come from the photoionisation of H.
Excitation potentials of these fine structure levels are ∼1 eV, which is approximately equal to the energy of the electrons ($kT \approx 1$ eV at $T_e = 10\,000$ K). *Note: all levels of H and He have much higher excitation potentials, meaning they cannot be excited by collisions with electrons of the same temperature.*

(b) Fine structure states are metastable, so the electron sits there for a relatively long time (0.01–0.1 secs). Remember, collisions with atoms are negligible. This gives the electron the chance to radiatively de-excite by itself (spontaneously), thus emitting a forbidden-line photon. The energy is now transferred to the photon.
Note: if the gas density is too high, the jump to the lower level can be induced by another

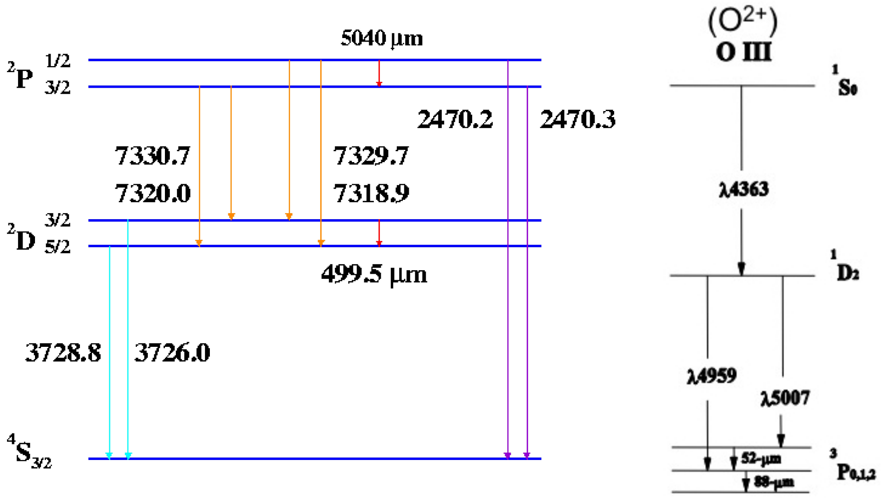

Figure 1: Energy level diagram for *left:* O II (O^+) and *right:* O III (O^{++}).

collision before spontaneous de-excitation has time to occur. This is called collisional de-excitation – no photon is emitted but kinetic energy is returned to the gas. For most H II regions the rate of radiative (spontaneous) de-excitation is greater that collisional de-excitation.

Every forbidden transition has a "critical density", above which the **collisional** de-excitation (by electron–atom collision) rate exceeds the **radiative** (spontaneous) de-excitation rate. In these circumstances we say the forbidden line becomes "quenched".

Despite the low abundance of metals (compared to H), forbidden transitions make a significant contribution because their excitation potentials (1–few eV) are so low. Only electrons with thermal (energy $\sim kT$) energies are needed, rather than ionising photons ($E > 13.6$ eV).

(c) because the photon emitted in the Forbidden Line transition has only a very weak probability of being absorbed the emitted photon easily escapes from the nebula gas, carrying away photon energy and this cools the gas.

By this process we have removed kinetic energy from the gas and transformed this to photon energy which escapes from the gas easily.

When this process is included in calculations of the heating and cooling rates it turns out that irrespective of the $T_{\rm eff}$ of the central OB star, we end up with electron temperatures of between 7 000–10 000 K, as observed.

The term 'forbidden' is really a misnomer! It was coined before we understood all the rules governing electron transitions. A more intuitive name for forbidden lines is "collisionally-excited lines" (CELs), since they are emitted following collisions of free electrons with ions in the nebular plasma.

For oxygen, the low energy states of singly ionised oxygen (O^+) and doubly ionised oxygen (O^{++}) are illustrated in Fig. 1. "Forbidden Transitions" amongst these levels include the lines at 3726, 3729 Å for O^+ and 5007, 4959, 4363 Å for O^{++}.

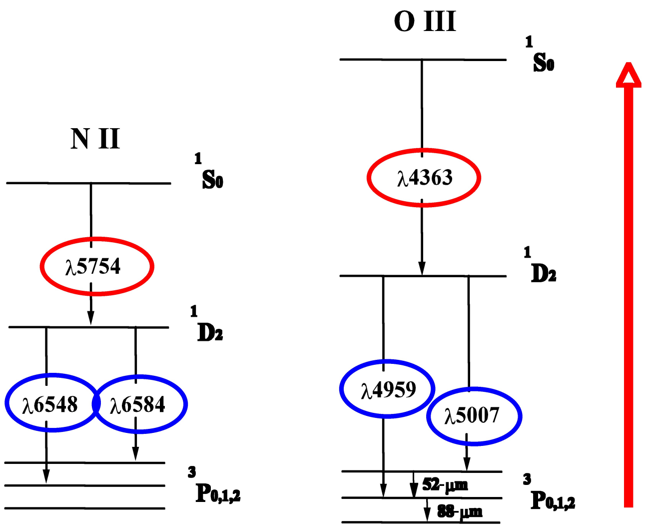

Figure 2: Energy level diagram for N II (N^+) and O III (O^{++}), showing the most commonly used temperature diagnostic ratios. The arrow indicates increasing energy.

Note: the ionisation potential energy of neutral O is about the same as for H (13.5 eV), whilst that of O^+ is about 35.1 eV. Thus only the hottest O stars can produce significant O^{++} in their H II regions.

Not all CELs result from forbidden transitions – it just happens that most optically observable CELs are forbidden. Moving only slightly into the UV, non-forbidden CELs can start to be seen (e.g. C IV $\lambda\lambda 1548, 1551$, or Si IV $\lambda\lambda 1394, 1403$).

2.4 How to derive temperatures, densities and the chemical composition of emission-line nebulae (H II regions, PNe)

It turns out that the observed relative intensities of forbidden lines in species like O^+, O^{++} and other forbidden lines in elements like sulphur, argon, chlorine provide good ways of measuring the electron temperatures and electron densities in H II regions.

The most widely used nebular 'thermometer' is the intensity ratio of the [O III] lines $\lambda 4363/\lambda 5007$, or equivalently $\lambda 4363/\lambda 4959$ [or even $(\lambda 4959{+}\lambda 5007)/\lambda 4363$] – see Figs 2 and 3. Intuitively, the temperature dependence can be seen from the fact that collisions with more energetic (hotter!) free photoelectrons are needed to push the bound electrons of the O^{++} ion in the upper state from where 4363 Å line is emitted than to populate the lower energy levels from where 4959 or 5007 Å are emitted. So the line strength ratio $\lambda 4363/\lambda 5007$ immediately tells us how hot the electron plasma in a nebula is. In a cool nebula the $\lambda 4363/\lambda 5007$ ratio will be lower than in a hotter nebula.

Similarly, the most commonly-used density measure is the intensity ratio of the [S II] lines $\lambda 6717/\lambda 6731$ or the [O II] lines $\lambda 3726/\lambda 3729$. Fig. 4 shows the energy level diagrams for these ions, where the transitions used are highlighted. Fig. 5 plots the density dependence of these line ratios for a given nebula temperature, T_e. It is immediately obvious that the usefulness of these indicators only spans a finite range in densities.

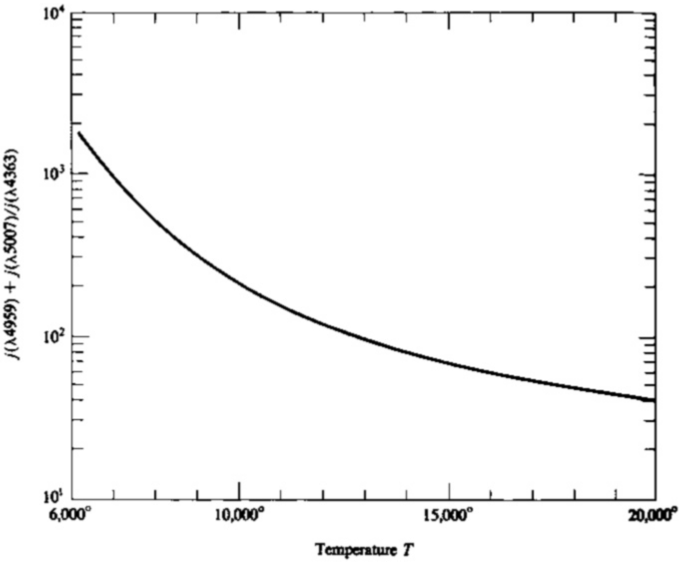

Figure 3: [O III](λ4959+λ5007)/λ4363 intensity ratio as a function of temperature. Reproduced from Osterbrock (1989).

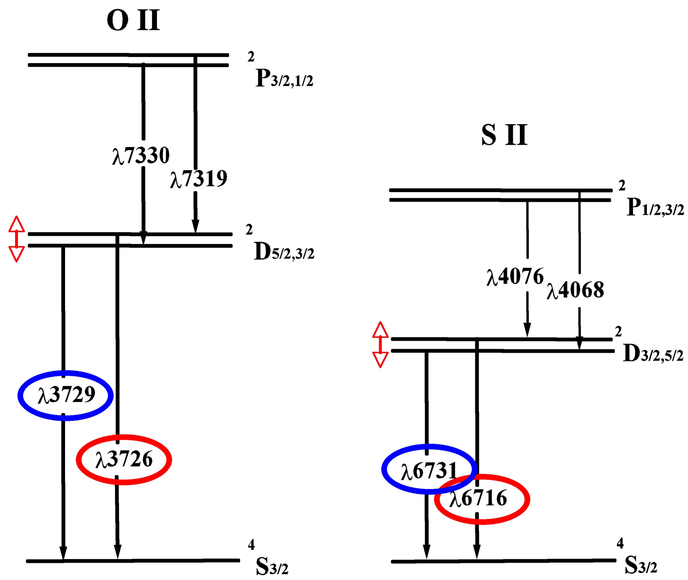

Figure 4: Energy level diagram for O II (O^+) and S II (S^+), showing the most commonly used density diagnostic ratios.

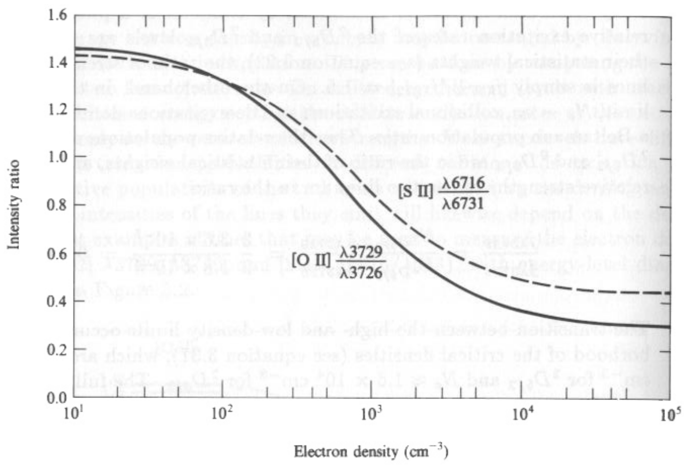

Figure 5: Variation of [O II] (solid line) and [S II] (dashed line) intensity ratios as a function of n_e at $T_e = 10000$ K. Reproduced from Osterbrock (1989).

The strengths of certain forbidden lines of heavy ions in an H II region or a planetary nebula, combined with knowledge of the electron temperature and density in the nebula, allows us to determine the abundance of these ions (and collectively of their respective element) *relative to hydrogen*. The abundance of O^{++}, for example, can be calculated using the following formula:

$$\frac{N(O^{++})}{N(H^+)} \sim \frac{n_e}{f(T_e)} \frac{I(\lambda 4959)}{I(H\beta)}$$

where n_e is the electron density; $f(T_e)$ is the fraction of O^{++} ions able to emit at 4959 Å (this has a strong dependence on nebular temperature – see Fig. 6) and $I(\lambda 4959)/I(H\beta)$ is the flux of the [O II] 4959 Å line relative to Hβ.

Now we can measure the strength of the forbidden lines from *all* the ionic stages of an element (e.g. O, O^+, O^{++}) and add up all the abundances to find the total abundance relative to hydrogen.

The technique is as follows: we first observe the nebula using a spectrograph coupled to a telescope. We measure the line strengths ("fluxes") relative to the strength of a representative hydrogen recombination line. The one typically used is Hβ 4861 Å (since it's strong and easily accessible in the optical part of the spectrum).

We use the observed ratios of the strengths of certain forbidden lines to determine the temperature and density of the nebular plasma.

We then use expressions such as that shown above to find out the abundance of ions relative to H^+, and add up the abundances of all the ionic stages to find the total abundance of a heavy element (e.g. C, N, O, Ne, S, Ar, Cl, etc.) relative to hydrogen in the nebula. Only with large telescopes can the weaker lines of the above elements be observed from nebulae in our Galaxy and beyond, allowing us to study the composition of distant galaxies. This information is vital for a proper understanding of the chemical evolution of the universe and the recycling of matter within galaxies (from stars to the ISM and vice-versa).

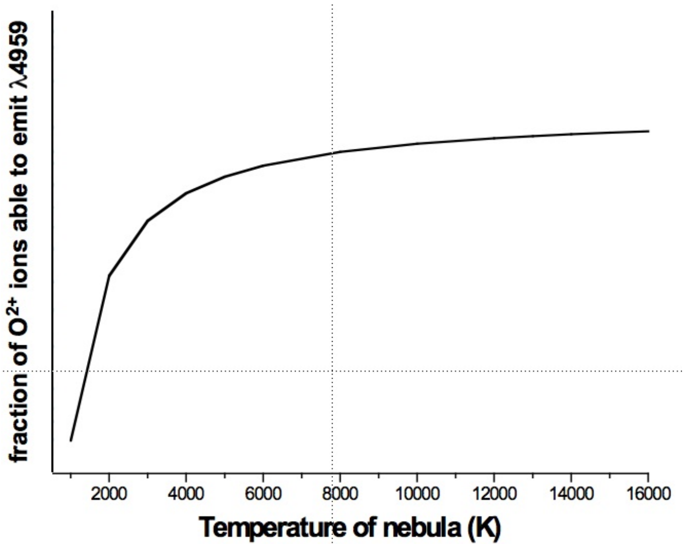

Figure 6: Fraction of O^{++} ions able to emit the 4959 Å line as a function of nebula temperature.

2.5 Additional information

UV forbidden lines (e.g. C III] 1908 Å; [O III] 1663 Å) are sampled from space (e.g. from the *IUE* or *FUSE* satellites). For IR lines (e.g. [O III] 52, 88 μm; [Ne III] 57 μm) we need a high altitude plane flying above most of water vapour in the atmosphere, since water absorbs in the IR (e.g. the Kuiper Airborne Observatory – KAO) or a satellite (e.g. *Spitzer*). For optical lines (e.g. [O III] 4363, 4959, 5007 Å, [O II] 3727, 3729 Å, [N II] 6548, 6584 Å, [Ar III] 7005 Å, [Ar IV] 4740 Å, [Ne III] 3968 Å, etc.), we can use ground based telescopes.

Ionised metal atoms can also emit recombination lines (allowed/permitted transitions), but are usually very weak (due to their low abundances). Like H recombination lines, they exhibit little or no dependence on nebular temperature or density so are not very useful.

Helium, which makes up about 10% (by number) of the gas in nebulae, does not emit Forbidden Lines, but recombination lines just like hydrogen does. The most prominent of these, and most accessible, are in the optical part of the spectrum: He I 4471 Å, 5876 Å, 6678 Å, and He II 4686 Å. Doubly-ionised helium however can only be found in planetary nebulae, as the cooler stars of H II regions do not emit photons energetic enough to ionise He^{+}. Hence the He II 4686 Å line (produced when He^{++} recombines with an electron to produce He^{+}) is NOT observed in H II regions.

+ Info

The abundance of helium in nebulae is derived from the observed strengths of the aforementioned helium recombination lines relative to Hβ. Sometimes He I recombination lines in the radio part of the spectrum are also used. Due to the way that helium and hydrogen recombination lines are created, the strengths of these lines do not depend very sensitively on the electron temperature in a nebula. Thus, even without a very accurate measurement of the nebular temperature, a rather accurate helium abundance (relative to H) can be derived.