INTRODUCTION

In the past, someone might join as a research student a group working on stars and using visible spectroscopy, and continue dealing with stars and using that wavelength range for the rest of his/her working life. Today the situation is very different. A spectrum of a faint point object presumed to be an 'ordinary' star can reveal something much more exotic, say with strong emission lines, whose analysis may take the researcher into very different regimes of temperature and density than those to which he or she is accustomed. Multi-wavelength programmes are increasingly common, and one individual can find him or herself in an extreme case using in one year X-ray, ultraviolet, visible, infrared, millimetre and radio observations. Not only are the techniques involved very different, but so is the terminology and, at least on the surface, the approach to interpretation. In some areas, like stellar atmospheres, there are several good textbooks at various levels, but in others the newcomer is reduced to searching for review articles or summer school lectures. There is a strong temptation to leave unfamiliar wavelength bands or temperature regimes to collaborators who, it is hoped, know what they are doing!

The aim of this book is to provide a consistent and coherent introduction to the interpretation of astronomical spectra from sources ranging from coronal densities of $10^8 \, \text{m}^{-3}$ and less and temperatures of $10^7 \, \text{K}$ to molecular clouds with temperatures of a few tens of degrees via stars with much higher density atmospheres but intermediate temperatures and gaseous nebulae (photoionized gases) with densities of the order of $10^9 \, \text{m}^{-3}$ and with temperatures of the order of $10^4 \, \text{K}$. It will, it is hoped, be found that many of the basic ideas applied to these different regimes are the same, even if they may appear to look somewhat different at first sight. The level of knowledge assumed is that of Physics undergraduates starting their final year, with a tolerant recognition of the fact that many such undergraduates seem to forget spectroscopic terminology as soon as they have finished their atomic or quantum physics courses. Most subjects are discussed first at a fairly general and qualitative level before they are considered in sufficient detail to enable the newcomer to the subject to be able to derive the

formed in LTE and some not; indeed, the contribution from deep levels in a stellar atmosphere may be in LTE but the contribution from high levels may deviate markedly from LTE, for the density is lower at higher levels and the optical depth to the edge of the atmosphere is small, so photons readily escape, reducing the level populations below their LTE values.

One has to try to estimate whether an LTE solution will be adequate for a particular purpose or what approximation to a full non-LTE solution might serve. In radio astronomy in particular one often uses the *excitation temperature* which is the temperature that would be needed to give the relative populations of the upper and lower levels of a transition *if* Boltzmann's equation held. If LTE holds, then the excitation temperature equals the kinetic temperature. A radiation field can often be characterized by a radiation temperature, which gives a Planck function approximating to the actual distribution of radiation. If the radiation field comes from a star at some distance, then the distribution of the field with frequency may be roughly Planckian at the effective temperature of the star, but the magnitude is reduced by a factor roughly proportional to the inverse square of the distance from the star called the *dilution factor*. If radiative processes are dominant and there is no dilution, then the excitation temperature will tend towards the radiation temperature.

If we now look particularly at lines, the concept of *saturation* is very important. Consider a thin cloud of gas with no continuous emission and some particular line transition which will, of course, give rise to an emission line. If we increase the number of atoms in the line of sight through the cloud, the line strength will at first increase in proportion to the number of atoms along the line of sight, but as the optical depth through the cloud increases some of the emission will be absorbed before it can escape from the cloud, and the line strength will increase more slowly than the number of atoms along the line of sight. Eventually increasing emission will be balanced by increasing absorption, the line strength will cease to grow, and we say that the line is saturated. Emission and absorption will balance to give an emission intensity equal to the Planck function evaluated at the line frequency and the excitation temperature of the line. The line is now optically thick. In general there will also be emission by continuum processes at all wavelengths, but at line wavelengths we have both continuum and line emission, so the line stands out above the continuum. In stars the continuum is optically thick, and if stellar atmospheres were isothermal both line plus continuum and continuum would radiate at the saturated strength and so the line would be invisible. However there is a temperature gradient in stellar atmospheres so the increased absorption coefficient at frequencies where we have both line and continuum opacity compared with that at neighbouring frequencies where we have continuum absorption alone means that optical depth 1 comes higher in the atmosphere at line frequencies. Since the temperature is lower higher up, the source function is usually lower, and the spectrum therefore looks less bright at the line frequency and we have an absorption line. At first the depth of such a line increases as the number of atoms capable of absorbing the line increases, but eventually we are looking at the top of the atmosphere and the line can become no deeper so again the line saturates.

We have ignored the fact that spectral lines are not monochromatic and have a natural breadth in frequency which is often swamped in astrophysical spectra by broadening due to motions in the gas (Doppler broadening) or by broadening due to interactions with other atoms, ions and electrons (collisional or pressure broadening). Not only is broadening of interest in itself, giving information on temperature, turbulence, rotation and pressure, but the line broadening affects saturation, for the centre of a line may be saturated while the wings are unsaturated and still growing with the number of atoms in the line of sight. Thus the frequency or wavelength integrated strength of the line depends on the line profile and the broadening mechanism.

In the last seven chapters we will apply some of these ideas to various astrophysical environments.

Thus the radiation field is given by *Planck's equation* which for frequency ν and temperature T is

$$I_\nu = \frac{2h\nu^3}{c^2} \frac{1}{e^{h\nu/kT} - 1}$$

(1.1)

where h is Planck's constant, c is the speed of light and k is Boltzmann's constant, here and throughout this book. The radiation field is, of course, isotropic. The source function in complete thermodynamic equilibrium is also given by Planck's equation. In complete equilibrium, the number of particles of mass M travelling with velocities between V and $V + dV$, is given by *Maxwell's equation*:

$$N(V)dV = 4\pi V^2 \left(\frac{M}{2\pi kT}\right)^{3/2} e^{-MV^2/2kT} dV$$

(1.2)

which is needed in estimating collisional rates. The fraction of particles of a particular atomic, ionic or molecular species excited to level n with excitation energy E_n above the ground state (the energy difference between the level and the ground state) is given by *Boltzmann's equation*:

$$\frac{N_n}{N} = \frac{g_n}{U} e^{-E_n/kT}$$

(1.3)

with

$$U = \sum_i g_i e^{-E_i/kT}$$

where g_n is the statistical weight of the nth level, i.e. the number of degenerate sub-levels in the nth level and U is called the partition function and serves to ensure that the populations N_n of all the levels add up to the total population N of the species concerned. Finally in complete equilibrium the ratio of the number of particles in a particular stage of ionization to the number in the next lower stage of ionization is given for an *ionization potential I* (the energy needed to remove an electron from the lower stage of ionization) by *Saha's equation*:

$$\frac{N(\text{higher stage})N_e}{N(\text{lower stage})} = \left(\frac{2\pi m_e kT}{h^3}\right)^{3/2} \frac{2U(\text{higher stage})}{U(\text{lower stage})} e^{-I/kT}$$

(1.4)

where N_e is the number of electrons per unit volume and m_e is the mass of the electron. A similar equation holds for the dissociation of a molecule, where for a diatomic molecule AB one replaces $N(\text{lower})$ by $N(\text{AB})$ and $N(\text{higher})$ and N_e by $N(\text{A})$ and $N(\text{B})$, m_e becomes the reduced mass of the molecule $m_A m_B/m_{AB}$, and I becomes the dissociation energy D, with appropriate changes to the partition functions.

Complete thermodynamic equilibrium cannot occur in a stellar atmosphere because it would require the temperature to be the same everywhere with no temperature gradient to drive an outward flow of radiation, and because it would

require an isotropic radiation field which again would imply no net flow of radiation. However, one can sometimes treat conditions locally as being characterized by a single local temperature which in Planck's equation gives the source function, in Saha's equation gives the ionization and in Boltzmann's equation gives the excitation, although the radiation field will not be precisely Planckian at the local temperature, and the 'single' temperature at which thermodynamic equilibrium holds will vary from layer to layer. In other words, what a given layer *does* to a radiation field which in general is not a thermodynamic equilibrium field is given by the equations of complete equilibrium characterized by the local kinetic temperature of the gas. This is called the approximation of *local thermodynamic equilibrium*, or LTE, as it is universally abbreviated, and its use enormously simplifies the solution of the equations of radiative transfer.

Elastic collisions in which particles exchange kinetic energy are much more frequent than collisions that produce excitation or ionization. Hence in the vast majority of astronomical circumstances Maxwell's equation for the velocity distribution holds, and in this book we shall assume that Maxwell's equation at the local kinetic temperature always gives the velocity distribution. It will be shown later that if Boltzmann's equation holds for the relative upper and lower level populations, then the source function reduces to the Planck function. Inelastic collisions producing excitation and de-excitation and ionization drive the degree of excitation towards that given by Boltzmann's equation and the degree of ionization towards that given by Saha's equation, so LTE is more likely in high density conditions. Equally if the optical depth is very large, then photons travel small distances so even if radiative processes are dominant over collisional processes, local conditions have a dominating effect and the local kinetic temperature will control the level populations via the occasional collisional process.

LTE will clearly hold in the interior of a star where the density is high, the average distance travelled by a photon is small, and the net outwards flux represents a small deviation from a nearly isotropic radiation field. Equally clearly we cannot expect LTE to hold at all in the thinnest parts of the interstellar medium where the density represents a good terrestrial vacuum, but the radiation field has major non-local contributions from the microwave background and the general star background. On the other hand it often turns out that in the lower density parts of the interstellar medium the equations of statistical equilibrium are particularly simple with only a few low-lying levels being significantly populated and the gas being optically thin so that any photon emitted is sure to escape without being reabsorbed. The latter point means that we do not need to solve an equation of radiative transfer at all. Similar considerations apply to gaseous nebulae, which also often have low densities and optical thinness at many wavelengths, with the radiation field supplied by the illuminating star. Unfortunately, most environments in stellar atmospheres and dense molecular clouds lie somewhere in between these extremes, the gas being of moderate density but optically thick and many levels being significantly populated. Some lines may be

basic equations and to obtain approximate results from observations. No claim is made here to have incorporated all the latest corrections and all the latest physical data such as transition probabilities—for these the reader is referred to the review articles. Units are S.I. (mks) throughout except for a few stray references to parsecs, and if a formula appears without units, S.I. units should be assumed. No discussion has been attempted here of observational techniques—there are other books dealing with these topics—but of course it is very important that the observer is fully aware of the observational constraints and problems when interpreting the data.

The first four chapters deal with fundamentals common to most regimes such as radiative transfer, statistical equilibrium and line formation and broadening. Subsequent chapters look at applications to stars, photoionized gases, the cold interstellar medium (which includes a discussion of masers), winds and circumstellar shells, ending with a brief discussion of coronal gases. We shall finish this introduction with a brief consideration of some of the concepts found in the opening chapters.

Radiative transfer is concerned with how the radiative energy flow changes as one moves through an emitting and absorbing gas. One wants in the end to be able to predict the spectrum that emerges from a given gas cloud or star as a function of direction and frequency, because this is what we observe and comparison of predictions with observations enables the composition, temperature and density of the cloud or stellar atmosphere to be ascertained. In general, the temperature and density will vary as one moves deeper into the object observed, particularly in the case of stars, so one would like to be able to model the cloud or stellar atmosphere—that is, for a given energy input to be able to predict the run of temperature through the cloud or stellar atmosphere. Any region in a cloud or stellar atmosphere must obey energy conservation with total energy input equal to total energy output. In the case of a non-extended stellar atmosphere, since there are no sources or sinks of energy in the absence of convection or other mechanical means of energy transport, the total flux (which measures the radiative energy flow) passing through successive layers of the atmosphere must be a constant, although the distribution of monochromatic flux with frequency will change. This condition of *flux constancy* expresses the law of energy conservation, and calculating the temperature gradient needed to drive a given constant flux (fixed by the luminosity and radius of the star) against the opposition of the opaqueness of the gas, together with the condition of hydrostatic equilibrium (with *g* fixed by the radius and mass of the star), enables a *model stellar atmosphere* to be produced giving temperature and pressure as a function of depth. In contrast, the temperature gradient is of much less importance in the case of an interstellar molecular cloud and the net radiative flow may be small with considerable non-radiative energy input in the form of cosmic rays, so that the simplest interpretations of molecular cloud spectra often assume a constant temperature. However, in more detailed work it is necessary to produce a model of the temperature distribution in molecular clouds as well.

One still has to determine the emissive and absorbing powers of the gas which enter into the equations of radiative transfer, to be more precise the *emission coefficient* and the *absorption coefficient*. These coefficients often appear in radiative transfer in the form of the ratio of emission to absorption coefficient, the *source function*, and of the *optical depth*, the integral of the absorption coefficient with respect to distance along a particular line of sight. The absorption coefficient at a particular frequency will be the sum of the absorption coefficients of all the atomic and molecular radiative processes that can take place at that frequency. The absorption coefficient for each process will be proportional to the product of the transition probability and the number of atoms or ions or molecules that are in the appropriate quantum state (lower energy level) for the transition concerned. Similarly, the emission coefficient will be proportional to the product of the transition probability and the number of atoms or ions or molecules in the appropriate upper energy level. In turn, the number of atoms or ions or molecules in the correct initial state depends not only on the abundance of the element or elements concerned, but also on the degree of excitation, ionization, and (for molecules) dissociation.

Now if we have a steady state (and in astrophysics we often do, but there are important exceptions such as shock waves and solar flares), and are given the composition, density and temperature of the gas under consideration and all the appropriate atomic and molecular constants, we can determine the degree of excitation, ionization and dissociation by using the equations of *statistical equilibrium*. For each quantum level, the rate of departure to all other levels by all processes (collisional as well as radiative) equals the rate of arrival from all other levels by all processes. An equation expressing this balance is written for each level, and the set of equations is solved simultaneously subject to the given total abundances of the elements involved and the given pressure and temperature, to obtain the populations of all the levels. In general this cannot be done because of the enormous number of levels involved which makes the matrix to be solved impossibly large. A second problem is that the rates of some of the radiative processes filling or emptying a level depend on the radiation field, which was what we were trying to find in the first place! Mathematically we need to solve simultaneously a large matrix and a set of differential equations.

The situation is made worse by the fact that many of the atomic constants involved, like transition probabilities and collisional cross-sections, are hard to measure experimentally and can only be estimated theoretically for the simplest atomic structures.

Under conditions of complete thermodynamic equilibrium the situation is greatly simplified because many of the quantities involved are related by the equations of statistical mechanics.