Chapter 2

The Physics and Chemistry

To appreciate the following story, we need to be familiar with the setting, the cast, their observational appearance and their physical interactions. Here, each member of the cast – gas, dust, cosmic rays, magnetic fields and radiation – is introduced and their potential behaviour is studied.

Stars form in clouds. Like atmospheric clouds, they are largely molecular and opaque. However, star-forming clouds are made more complex by a quite stunning range of physical, chemical and dynamical processes. The processes work in synthesis to provide triggers, regulators and blockers during the collapse stages. Energy is transferred from large scales down to small scales. Simultaneously, feedback occurs from small scales back onto large scales. As we shall see, the progress of a parcel of gas is far from systematic – at each stage in the development there is a chance that it will be rejected. It will, therefore, prove rewarding to first acquire a broad overview.

2.1 Scales and Ranges

At first sight, when interstellar clouds and atmospheric clouds are compared, they appear to have little in common. According to Table 2.1, they could hardly be more different. Yet they do resemble each other in other respects. Besides both being molecular and opaque, they both produce magnificent reflection nebula when illuminated by a nearby star, as illustrated in Fig. 2.1. Most relevant, however, is that they are both ephemeral: they are transient with lifetimes as short as their dynamical times (as given by their size divided by their typical internal speed, as listed in Table 2.1). In other words, their own swirling, turbulent motions disperse the clouds rapidly.

(a) (b)

Fig. 2.1 Molecular clouds and atmospheric clouds. Similar hydrodynamic processes shape (a) interstellar clouds and (b) cumulus clouds in the sky despite contrasting scales. Both are illuminated by stars. Image (a) is a detail from the Eagle Nebula (M16) displayed fully in Fig. 2.2 (Credit: (a) J. Hester & P. Scowen (Arizona State U.), HST & NASA) and (b) Shawn Wall.

Table 2.1 A comparison of scales between typical molecular and atmospheric clouds.

	Molecular Cloud	Atmospheric Cloud
Size	10^{14} km	1 km
Mass	10^{36} gm	10^{11} gm
Particle density	10^{3} cm^{-3}	10^{19} cm^{-3}
Temperature	20 K	260 K
Mol./atomic weight	2.3	29
Speed of sound	0.3 km s^{-1}	0.3 km s^{-1}
Turbulent speed	3 km s^{-1}	0.003 km s^{-1}
Dynamical time	Million years	Five minutes

Why hasn't there been a consensus on the star formation mechanism, given that there are only a few components to interact? The answer lies in the range of physical, chemical and dynamical processes which link them. This results in a variation of the dominant laws from stage to stage. In addition, we meet extreme regimes way beyond our experience, as summarised in Table 2.2.

High compressibility provides the major conceptual distinction of an interstellar cloud from a cumulus cloud. A parcel of gas from the hot interstellar medium in our Galaxy would decrease its volume by a factor of 10^{26} on its way to becoming a star, passing through the states shown in Table 2.2. That is, its number density increases from $0.01\,\mathrm{cm^{-3}}$ to $10^{24}\,\mathrm{cm^{-3}}$ (compare to the molecular density of air of $10^{19}\,\mathrm{cm^{-3}}$), and so reaches a mass density of almost $1\,\mathrm{g\,cm^{-3}}$.

We will encounter interstellar gas with a wide range in temperatures. Molecular clouds are cold (5–50 K) but protostars send shock waves into the clouds, capable of temporarily raising the temperature to over 10,000 K without destroying the molecules (§10.5). Atoms can be raised to temperatures in excess of a million Kelvin through faster shock waves and in the highly active coronae of young stars (§9.7).

Table 2.2 Major star formation scales. The temperature, T, is in Kelvin and the final column lists the dynamical time scale in seconds.

Phase	Size (cm)	Density $\mathrm{g\,cm^{-3}}$	T (K)	Time (s)
Atomic ISM	10^{21}–10^{20}	10^{-26}–10^{-22}	10^{6}–10^{2}	10^{15}
Molecular cloud	10^{20}–10^{18}	10^{-22}–10^{-18}	10^{2}–10^{1}	10^{14}
Protostar collapse	10^{18}–10^{12}	10^{-18}–10^{-3}	10^{1}–10^{6}	10^{13}
Pre-main-seq. contraction	10^{12}–10^{11}	10^{-3} –10^{0}	10^{6} – 10^{7}	10^{15}

Lengths are measured in three units according to how best to appreciate the scale in discussion. Thus, cloud sizes are measured in terms of parsecs ($1\,\mathrm{pc} = 3.09 \times 10^{18}\,\mathrm{cm}$), cores, stellar and planetary systems in terms of Astronomical Units ($1\,\mathrm{AU} = 1.50 \times 10^{13}\,\mathrm{cm}$) and single stars in terms of the solar radius ($1\,\mathrm{R_\odot} = 6.96 \times 10^{10}\,\mathrm{cm}$). Hot diffuse gas occupies galactic kiloparsec scales while gas accretes to within a few solar radii of the growing young star. This is a difference in scale of ten magnitudes of ten from 10^{21} cm to 10^{11} cm.

Time is usually given in years, that is 3.15×10^{7} seconds. Speeds, however, are expressed in centimetres or kilometres per second, whichever is appropriate for the phenomena studied.

Mass is usually expressed in solar units rather than grams ($1\,\mathrm{M_\odot} = 1.99 \times 10^{33}\,\mathrm{g}$) and gravitational acceleration is given by GM/R^2

for a distance R from a mass M, where $G = 6.67 \times 10^{-8}\,cm^3\,g^{-1}\,s^{-2}$ is the gravitational constant.

While the mass density of gas is given in $gm\,cm^{-3}$, it is often useful to measure the mass per cubic parsec for the stellar content where $1\,M_\odot\,pc^{-3} = 6.8 \times 10^{-23}$ g cm^{-3}. A particle density is often more appropriate but can often be confusing since we employ the hydrogen atomic density, n(H), hydrogen molecular density $n(H_2)$, hydrogen nucleon density ($n = n(H) + 2n(H_2)$), and the total particle density, n_p including an extra $\sim 10\%$ of helium atoms (above the hydrogen nucleonic number) which contribute 40% more by mass. The abundances of other elements are small and can be neglected in the overall mass budget of a cloud.

Luminosity and power are expressed in solar units ($1\,L_\odot = 3.83 \times 10^{33}\,erg\,s^{-1}$) where the erg is the CGS abbreviation for $1\,g\,cm^2\,s^{-2}$ of energy. The KMS unit of Watt ($1\,W_\odot = 10^7\,erg\,s^{-1}$) is also often employed. Observers find that the traditional 'magnitude' unit provides a convenient logarithmic scaling of luminosity. The magnitude system is particularly convenient when extinction is in discussion since it is linearly related to the amount of obscuring material, whereas the luminosity falls off exponentially. Magnitudes will be defined and employed in §2.4.1.

Expressions for energy and temperature present the greatest variety. Astrochemists and X-ray astronomers often discuss in terms of electron Volts where $1\,eV = 1.60 \times 10^{-12}$ erg (see Table 2.3). The frequency of radiation, ν, also converts to an energy $h\nu$ where h is the Planck constant, $h = 6.63 \times 10^{-27}$ erg s. As an example, ionisation of cold H_2 requires an energy exceeding 15.4 eV. So, photons with $\nu > 3.72 \times 10^{15}$ Hz are required which, according to Table 2.3, are ultraviolet photons.

The excitation energies of atoms and molecules are often expressed as a temperature using $E = kT$ where $k = 1.38 \times 10^{-16}\,erg\,K^{-1}$ is the Boltzmann constant.

Wavelengths are often more convenient than frequencies since the length of a wave can be directly compared to the size of atoms, molecules and dust particles. Thus, the Ångstrom, $1\ \text{Å} = 10^{-8}$ cm, and the micron, $1\ \mu m = 10^{-4}$ cm, are often useful. Units for radiation and wavelengths are summarised for reference in Table 2.3.

Table 2.3 Wavelengths and energies relevant to star formation studies.

Regime	Wavelength	Energy	Frequency (Hz)
Radio/millimetre	0.1–1000 cm		3×10^7–3×10^{11}
Sub-millimetre	300–1000 μm		3×10^{11}–1×10^{12}
Far-infrared	10–300 μm		1×10^{12}–3×10^{13}
Near-infrared	0.8–10 μm		3×10^{13}–4×10^{14}
Optical	4000–8000 Å		4×10^{14}–7×10^{14}
Ultraviolet	3000–4000 Å		7×10^{14}–1×10^{15}
Far UV	912–3000 Å	4–13.6 eV	1×10^{15}–3×10^{15}
Extreme UV	100–912 Å	13.6–100 eV	3×10^{15}–2×10^{16}
Soft X-ray	-	0.1–2 keV	2×10^{16}–4×10^{17}
X-ray	-	2–1000 keV	4×10^{17}–2×10^{20}
Gamma ray	-	1–1000 MeV	2×10^{20}–2×10^{23}

2.2 The Ingredients

2.2.1 *Atoms, molecules and dust*

The interstellar medium, or ISM, is a broad name for all that exists between the stars within galaxies. It includes diverse clouds of gas, which are composed mainly of atoms, molecules and ions of hydrogen, (H), electrons (e), as well as small amounts of other heavier elements in atomic and molecular form. There is a constant composition by number of about 90% hydrogen and 9% helium. The abundances of other atoms, mainly carbon, oxygen or nitrogen, vary depending on the enrichment due to stellar nucleosynthesis (the creation of new atoms by fusion in stars) and the removal due to condensation onto solid particles called dust grains (see below).

Molecules have been found concentrated in dense aggregates called molecular clouds. These are cold and dark regions in which hydrogen molecules outnumber other molecules by 1000 to 1 on average. This remains true until the heavy elements held in dust grains (see below) drift to the midplane of a disk surrounding a young star, the location where planets may eventually form.

Before 1970 there was little evidence for interstellar molecules. This all changed when millimetre, infrared and ultraviolet astronomy started. Now, more than 120 molecular species have been detected and identified in space. In molecular clouds, besides H_2, these include OH, H_2O, NH_3, CO and many more complex organic (carbon based) ones including formaldehyde, ethyl alcohol, methylamine and formic acid. Recent observations seem to indicate that the amino acid glycine may even be present in these clouds. At least, the molecular precursors (e.g. HCN and H_2O) are known to inhabit these clouds. There are obviously many more molecules to be discovered but detection is more difficult for molecules of greater complexity.

The ISM also includes vast numbers of microscopic solid particles known collectively as interstellar *dust*. They consist mainly of the elements carbon (C) or silicon (Si) with H, O, Mg and Fe in the form of ices, silicates, graphite, metals and organic compounds. The Milky Way contains vast lanes of dust which, being dark, were originally thought to be due to the absence of stars. In fact, dust forms about 1% by mass of all interstellar matter. Most of the dust mass is contained in the larger grains of size exceeding 1000Å, and contain 10^9 atoms. Others are more like large molecules, such as the 'polycyclic aromatic hydrocarbons' (PAHs), consisting of perhaps 100 atoms. By number, most of the grains are actually small (50 Å).

Dust is produced when heavy elements condense out of the gaseous phase at temperatures less than 2000 K. To each element belongs a condensation temperature T_c, at which 50% of the atoms condense into the the solid phase when in thermodynamic equilibrium. For the refractory (rocky) elements (e.g. Mg, Si, Fe, Al, Ca) $T_c \sim 1200$–1600 K; for the volatile elements (O, N, H and C), all critical to life, $T_c < 200$ K.

The origin of the dust is the cool expanding outer layers of evolved red giant stars. These winds are conducive to the condensation of grains from the refractories. The grains are ejected along with gases to contaminate ISM material at a rate of $10^{-7}\,M_\odot\,yr^{-1}$ per star. The dust is subsequently widely distributed throughout the interstellar medium by the blast waves from supernovae. The dust will cycle several times through diffuse and dense clouds and so becomes well mixed and and processed. In the cold dense clouds, the condensation of other (even volatile) molecules (water, methane etc.) takes place. In molecular clouds, the grains act as nucleation sites for the condensation of even volatile molecules (e.g. water, methane) and the growth of the mantles.

We have discovered numerous inorganic and complex organic molecules in the dense molecular clouds. These molecules may survive in comets and asteroids which could have bombarded the youthful Earth (and other planets) to provide an injection of organic molecules and volatiles (e.g. water) needed for their formation. Some speculation exists that life may even evolve in these clouds and that the Earth was 'seeded' by such a cloud (panspermia).

2.2.2 *Cosmic rays, ions and magnetic field*

Extremely energetic nuclei known as *cosmic rays* penetrate everywhere, processing the ISM through collisions. They mainly consist of protons and electrons with energies that can exceed 10^{20}eV. On collision with hydrogen nuclei, they produce particles called π-mesons which decay into gamma rays. Thus, by measuring the gamma-ray flux, we can constrain the density of hydrogen nuclei. Since we are usually forced to measure cloud mass through trace ingredients, such as CO, cosmic rays provide an important means of corroboration of the H_2 density.

Cosmic rays also penetrate deep into clouds where ionising UV radiation is excluded. In practice, this is at depths which exceed 4 magnitudes of visual extinction (as defined in §2.4.1). The result is that cosmic rays provide a minimum degree of ionisation even in very cold optically thick clouds. This is crucial since the ions interact strongly with both the magnetic field *and* the neutral molecules, binding the field to the fluid. As a result, stars might not form due to the resisting pressure of the magnetic field unless the ions and magnetic field can drift together through the molecular fluid (see §7.5). The ion density n_i is fixed by balancing the ionisation rate in unit volume,

$$\frac{dn_i}{dt} = \zeta \times n(H_2) \quad \mathrm{cm}^{-3}\mathrm{s}^{-1}, \tag{2.1}$$

simply proportional to the number density, with the recombination rate, $5 \times 10^{-7}\,\mathrm{n}_i^2\,\mathrm{cm}^{-3}\,\mathrm{s}^{-1}$. Here, the coefficient $\zeta = 3 \times 10^{-17}\,\mathrm{s}^{-1}$ is taken. Equating formation and destruction yields an equilibrium ion fraction

$$\frac{n_i}{n(H_2)} \sim 2.4 \times 10^{-7} \left(\frac{n(H_2)}{10^3\ \mathrm{cm}^{-3}} \right)^{-1/2}. \tag{2.2}$$

Therefore, an extremely low fractional ionisation is predicted and, indeed, found. Incidentally, the ionisation process also leads to the production of H atoms, which we quantify below in §2.4.3.

An all-pervading but invisible magnetic field exerts a force on electrically-charged particles and so, indirectly, influences clouds of gas and dust. Another effect of the field is to align elongated dust grains. The spin axes of the grains tend towards the magnetic field direction while the long axes orient perpendicular to the field. This effect allows the field direction projected onto the plane of the sky to be deduced. There are two methods. The first method is to measure the polarisation of background starlight. The second method studies the linear polarisation of thermal dust emission. The field direction is transverse to the direction of polarisation, although there are other flow and field effects which may confuse the results. This method often yields a combination of a uniform field and a chaotic structure. Hourglass shapes and toroidal fields have also been reported. The second method is to measure the polarisation of background starlight.

The field strength has proven extraordinarily difficult to measure, mainly relying upon the quantum effect of Zeeman splitting. Molecular lines which have been successfully utilised are CN at 0.3 cm, H_2O at 1.3 cm and OH at 2 cm and 18 cm. The strength and role of the field will be explored in §6.6.

Indirect estimates of the field strength involve large modelling assumptions and, to date, have been applied only to a few, perhaps not typical, cloud regions. These methods involve modelling locations where molecules are excited and compressed in shock waves (as C-shocks – see §10.5), or relate the field strength to the resistance of the field to being twisted or bent in a turbulent medium.

Finally, the ISM is awash with numerous photons of electromagnetic radiation originating from nearby stars as well as a general background from the galaxy as a whole. This radiation is not so pervasive as the cosmic rays or magnetic field but dominates the physics of the cloud edges and provides the illumination for the silhouettes of the dark clouds such as shown in Fig. 2.2.

The cosmic microwave background, the heat of the cooling, expanding Universe, has a temperature $T_{bg} = 2.73\,\mathrm{K}$. Molecular clouds may indeed get this cold. In the early Universe, at high redshifts (see §13.1.1), the temperature was much higher and microwave background photons inhibited H_2 formation and so probably delayed primordial star formation (see §13.1.1).

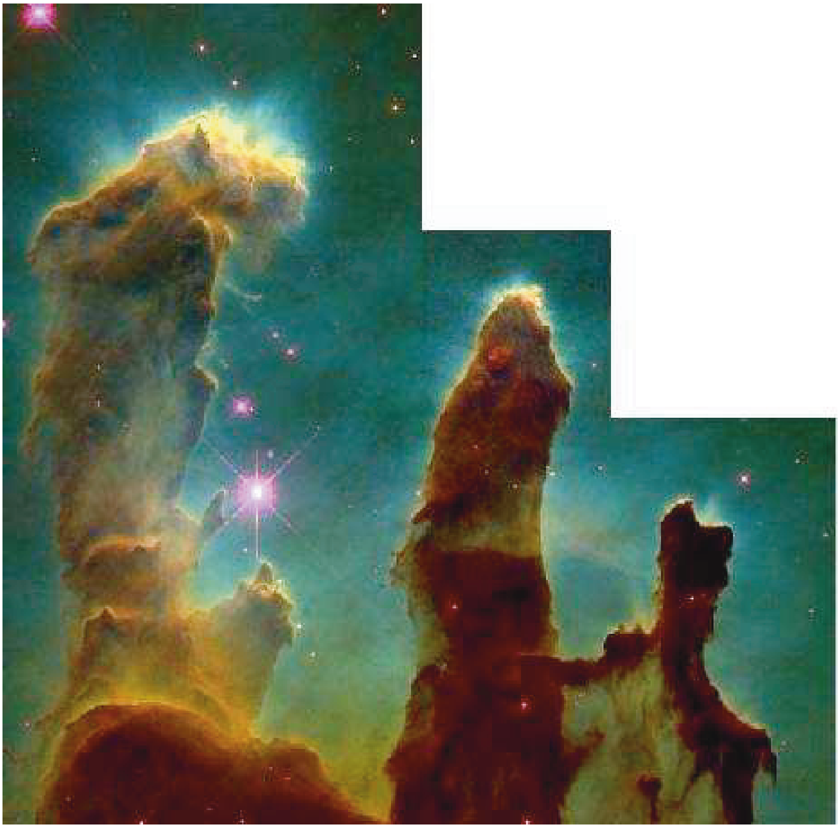

Fig. 2.2 EGGs are evaporating gaseous globules emerging from pillars of molecular hydrogen gas and dust. This image of the Eagle Nebula in the constellation Serpens demonstrates the effects of ultraviolet light on the surfaces of molecular clouds, evaporating gas and scattering off the cloud to produce the bright reflection nebula. Extinction by the dust in the remaining dark pillars, resistant to the radiation, blot out all light. Stellar nurseries of dense EGGs are exposed near the tips of the pillars. The Eagle Nebula, associated with the open star cluster M16 from the Messier catalogue, lies about 2 kiloparsecs away. (Credit: J. Hester & P. Scowen (Arizona State U.), HST & NASA).

2.3 Observations

2.3.1 *Radio*

Gas can be excited between specific quantum energy levels and so emit or absorb light at discrete frequencies. Line emission at radio wavelengths can arise from both atoms and molecules. The 21 cm (1420 MHz) line of neutral atomic hydrogen has been used to map out the large scale distribution of atomic clouds. The emission is produced from the ground state of neutral hydrogen where the electron spin axis can be aligned parallel or anti-parallel to the proton's spin axis. When the electron is parallel it is in a higher energy state than when anti-parallel. Mild inter-atomic impacts knock an atom into the higher energy state so that it may then emit upon returning.

Atomic lines are also generated in the radio following a cascade down energy levels after an ion recombines with an electron. Whereas the optical Hα line arises from an energetic transition deep within the potential well of the hydrogen atom, other recombination lines originate from very high levels of H as well as other atoms, producing lower energy photons.

Maser emission may be generated when the populations of two energy levels of a molecule become inverted when in a steady state. That is, when the higher energy level also contains the higher population. In this case, a photon with energy corresponding to the transition energy is likely to stimulate further photons moving in the same direction, rather than be absorbed. It follows that the emission can grow exponentially rather than being damped exponentially by absorption. Amplification must be limited by saturation, i.e. it is limited by the rate at which the molecule is pumped into the higher level. Maser spots are so produced, representing those locations where the narrow cone of emission points towards us. H_2O and OH masers are often observed in star formation regions. Other inversions occur in methanol (CH_3OH), ammonia (NH_3) and formaldehyde (H_2CO) and also prove useful tracers of conditions and structure.

Finally, continuum emission is produced as a result of the deceleration of electrons during collisions with ions. This is called free-free emission and is particularly observable at radio wavelengths although the spectrum is much broader.

2.3.2 *Millimetre, submillimetre and far-infrared*

The main constituents, H_2 and He, cannot radiate at the low temperatures of molecular clouds. We, therefore, rely on the ability as collision partners to excite heavier trace molecules such as CO, NH_3 and HCN. These molecules are detected through emission in spectral lines due to transitions in rotational, vibrational and electronic energy states. Downward transitions tend to give radiation in the infrared (vibrational) and submillimetre and millimetre (rotational), whereas electronic changes emit photons in the ultraviolet (UV) and visible range.

The CO molecule is the most abundant tracer with an abundance 10^{-4} times that of the H_2. The CO lines can be employed to determine the temperature. For example, the first three transitions on the rotational (J) ladder emit at 115 GHz (J=1–0), 230 GHz (J = 2–1) and 345 GHz (J = 3–2) and are ideal for tracing cold gas with temperature in the range 5–40 K. The molecule is a linear rotor and is detectable at millimetre wavelengths not only in ^{12}CO but in the ^{13}CO and C^{18}O isotopes also. The energy levels for such diatomic molecules are approximately given by a rigid rotator with E = B(J+1)J, where B is a constant, with a selection rule ΔJ = $\pm$1. This means that the higher-J transitions generate lines in the infrared and highlight warm gas, e.g. the ^{12}CO J = 19–18 transition at 140 μm will produce most CO emission for gas at 1000 K while the J = 6–5 transition at 434 μm represents the peak for 100 K gas. In this manner, temperatures of clouds can be probed through the ratios of low-J and mid-J rotational CO intensities as well as numerous other molecular line ratios.

The density can be constrained by measuring emission from different molecules. This is because some molecules are more prone to collisional transitions and others to radiative transitions. We demonstrate this by taking a highly simplistic model. The average distance that a molecule will travel before colliding with another molecule is called the mean free path, λ_p. This distance will be shorter in denser regions or where the molecules present larger cross-sectional areas. Therefore, we expect

$$\lambda_p = \frac{1}{\sigma_p \; n(H_2)} \tag{2.3}$$

where σ_p, the collision cross-section, is related to the impact parameter (approximately the molecule size) of order of 10^{-15} cm^2. The rate of collisions of a molecule with the H_2 is then the inverse of the collision timescale t_C

$$C = \frac{1}{t_C} \sim \sigma_p n(H_2) v_{th} \tag{2.4}$$

where the mean velocity can be approximated as $v_{th} \sim 10^4 \sqrt{\mathrm{T}}$ cm s^{-1}. Hence the collision or excitation rate is roughly $C = 10^{-11}\, n(H2)\, \sqrt{T}\,\mathrm{s}^{-1}$ for all molecules. We now compare this to the known de-excitation rate through spontaneous emission, A, which is the probability per unit time for radiative decay. If $C << A$ then excitations are relatively inefficient. The cloud density at which the molecule is likely to radiate most efficiently and, hence, the critical density at which it serves as a tracer is where $C \sim A$. Taking a cold cloud of temperature 10 K then gives the following.

The CO 1–0 line at 115 GHz (with $A = 7 \times 10^{-8}$ s^{-1}) traces low density molecular clouds with $n = 10^2 - 10^3$ cm^{-3} and is, therefore, the choice for mapping galactic-scale molecular distributions. On the other hand, CS 2–1 at 98 GHz (with $A = 2 \times 10^{-5}$) traces high density molecular clumps with $n \sim 10^6$ cm^{-3}.

Although CO is easily excited, radiative transfer will affect the detectability. If the opacity is high, the emitted photon will be reabsorbed, reducing the value of A. Self-absorption due to lower excitation foreground gas will also dramatically distort observed line profiles. To measure the optical depth, we utilise lines from different isotopes. For example, if the 1–0 lines from both ^{12}CO and ^{13}CO are optically thin, then the intensity ratio is given by the abundance ratio of *sim* 90. If we find the measured ratio is considerably lower, then we can conclude that the ^{12}CO line is optically thick.

Certain atoms and ions possess fine structure in their electronic ground state. This results from coupling between orbital and spin angular momenta. The splitting is typically 0.01–0.1 eV and the upper levels are relatively easily reached through collisions in quite warm dense gas. Hence, provided the atoms are not bound onto dust or molecules, they can trace physical conditions in clouds. In particular, the C I line at 609 μm, the singly-ionised C II line at 158 μm and the O I line at 63 μm are often prominent in far-infrared spectra.

2.3.3 *Infrared observations*

Molecular hydrogen can be directly observed in the infrared. Unfortunately, it can only be observed when in a warm state or exposed to a strong UV radiation field. Being a symmetric molecule, dipole radiation is forbidden and only quadrupole transitions with rotational jumps $\Delta J = 2$ (or 0, of course, for a vibrational transition) are allowed. This implies that there are two types or modification of H_2: ortho (odd J) and para (even J). The difference lies in the nuclear spins which are only changed in certain types of collision. For ortho H_2, the nuclear spins are aligned to yield a spin quantum number of $S = 1$ while for para H_2 the spins are anti-parallel and $S = 0$. Quantum mechanics then yields the number of different discrete states that the molecule can exist in as proportional to $(2J+1)(2S+1)$. For this reason, ortho lines are often found to be a few times stronger than para lines when a sufficient range of levels are occupied. On the other hand, the least energetic transition is in para-H_2 between $J = 0$ and $J = 2$, with an energy of 510 K. For this reason, para H_2 will be dominant in cold gas in which the molecule approaches equilibrium.

The most commonly observed H_2 lines lie in the near-infrared. Water vapour in the atmosphere contributes substantially to the opacity leaving specific wavelength windows open to ground-based telescopes. These are centred at wavelengths of 1.25 μm (J-band), 1.65 μm (H-band) and 2.2 μm (K-band). These transitions are ro-vibrational which means that changes in both vibrational and rotational levels are permitted. The most commonly observed spectral line is the 1–0 S(1) transition at 2.12 mum, used to image hot molecular gas where protostellar outflows impact with their molecular cloud. The 1–0 denotes the vibrational change from the first to the ground level, while the S(1) denotes the change in rotational state with S denoting $\Delta \mathrm{J} = +2$ and the final state is $\mathrm{J} = 1$ (the letters Q and O traditionally denote the other permitted transitions $\Delta \mathrm{J} = 0$ and -2, respectively).

Warm dust particles emit light across a broad continuous range of infrared frequencies, producing a 'continuum' given by Kirchoff's laws. The spectrum will peak at frequencies proportional to the temperature of the dust, as quantified in §6.1. Most bright adult stars tend to emit in the visible and above because of their higher temperature. Hence dust generally stands out in the infrared bands.

2.3.4 *Optical, ultraviolet and X-rays*

Dense atomic gas which is ionised either by radiation or by collisions in shock waves will subsequently recombine. The recombination spectrum consists of numerous well-studied lines, generated as the recombining electron cascades down the energy ladder. In particular, the Hα Balmer line of H (from the $\mathrm{n} = 3$ to the $\mathrm{n} = 2$ orbit, where n is the principle quantum number) at 6563 Å is mainly responsible for the red colour of many so-called reflection nebula. The Lyman-alpha line Lyα, at 1216 Å (from $\mathrm{n} = 2$ to the ground $\mathrm{n} = 1$), is significant as a coolant on top of contributing significantly to the ultraviolet.

Cold interstellar gas can be detected through atomic hydrogen absorption lines (in the visible), and to a lesser extent the molecular hydrogen absorption lines (in the UV which is efficiently absorbed). Of course, we require the nebula to be back-lit by a star and for some of the stellar light to survive the passage through the cloud. These lines can be distinguished from the star's lines by the fact that they will be quite sharp since the gas is much cooler.

X-ray emission is not observed from molecular clouds. Hot gas in the interstellar medium and in the vicinity of young stars emits Bremsstrahlung emission, which is a form of free-free emission as described in §2.3.1.

2.4 Processes

2.4.1 *Interstellar extinction and reddening*

Dust is recognised as the substance which makes Dark Clouds, such as the Horsehead Nebula, dark. The dust obscures light from objects in the background as well as from internal objects. Dust can be detected due to four primary effects that it has on starlight. These are Extinction, Reddening, Polarisation and Infrared Emission.

Dust extinction is the dimming of starlight caused by absorption and scattering as it travels through the dust. The ability for photons to penetrate is proportional to the column of dust in a cloud, which is proportional to the column density of gas, the number of hydrogen atoms per unit area, in its path, for a given ratio of dust to gas particles. The extinction grows exponentially with the column: we find that a layer of gas with a column of $\mathrm{N}_H = 2 \times 10^{21}\,\mathrm{cm}^{-2}$ usually contains enough dust to decrease the number of visual photons by a factor of 2.5. We term this factor a *magnitude* of extinction. This is evident from decades of observations of the extinction cross-section, summarised in Fig. 2.3. A further column of gas thus reduces the remaining radiation by another factor of 2.5. Therefore, a column of gas of $10^{23}\,\mathrm{cm}^{-2}$ reduces the number of photons by fifty 'magnitudes' of 2.5, or a factor of 10^{20}. Such columns are often encountered in star-forming cores, obliterating all visual and UV radiation from entering or escaping. To summarise, the visual extinction typically indicates the column of interstellar gas according to

$$A_V = \frac{N_H}{2 \times 10^{21}\ \mathrm{cm}^{-2}}. \tag{2.5}$$

Fortunately for us, extinction is not equal at all wavelengths – it is now often possible to see through the dust thanks to advances in radio and infrared astronomy. This is illustrated in Fig. 2.3. For example, the near-infrared extinction at 2.2μm in the K-band is related to the visual extinction by $\mathrm{A}_K \sim 0.11\,\mathrm{A}_V$. An infrared extinction law of the form $\mathrm{A}_\lambda \propto \lambda^{-1.85}$ is often assumed. Dust emission and absorption bands generate prominent features in the infrared spectra. A well-known bump at $9.7\,\mu$m is attributed

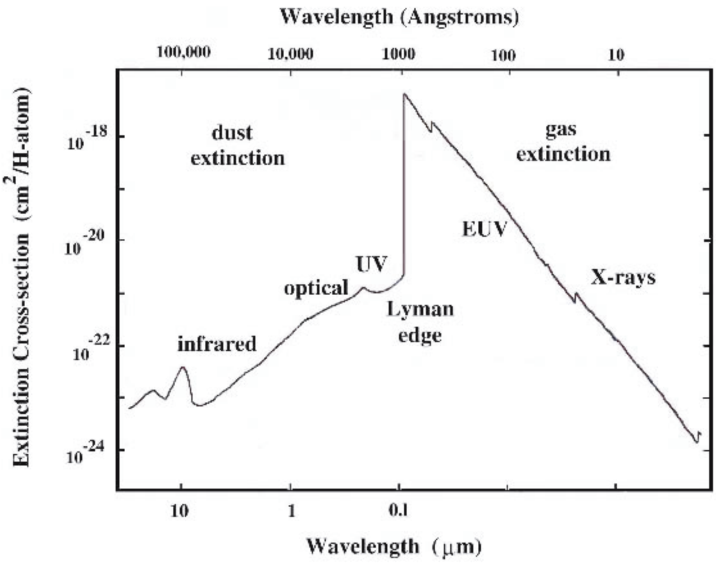

Fig. 2.3 A schematic diagram displaying interstellar extinction across the spectrum. Note the wide range in extinctions and the advantage of infrared observations over optical and ultraviolet (from original data accumulated by Ch. Ryter).

to a silicate feature from Si–O stretching. A 3.07μm feature is accredited to water ice. Other ice features include CO and NH_3. It appears that the volatiles condense out in dense regions onto the resilient silicate and graphite cores. Other previously 'unidentified bands' appear to be caused by Polycyclic Aromatic Hydrocarbons (PAHs), discussed in §2.2.1.

Reddening occurs because blue light is more strongly scattered and absorbed than red. This is quantified by the colour excess produced between any two wavelengths. Measured in magnitudes, the excess is proportional to the extinction provided the dust properties in the ISM are uniform. For example, we can derive the selective extinction $\mathrm{R} = \mathrm{A}_V/\mathrm{E}_{B-V} = 3.1$ as characteristic of the diffuse ISM where E_{B-V} is the relative excess produced by extinction between the optical filters B (0 .44μm) and V (0.55μm). In dark clouds such as ρ Ophiuchus, $\mathrm{R} \sim 4.2$, indicating the presence of larger grains, due to coagulation or to the accretion of volatiles onto grain mantles.

Starlight can become polarised when passing through a dust cloud. Elongated grains tend to have their spin axes aligned with the magnetic field, producing polarisation along the magnetic field direction. The reduced extinction in the infrared allows the magnetic field direction to be traced in dense clouds. Scattering even off spherical grains can also polarise light. In this case the polarisation is orthogonal to the local magnetic field.

Finally, the dust particles are responsible for much of the observed infrared radiation. Each grain will absorb visible and UV light from nearby stars, heat up and emit in the infrared as much energy as it absorbs. The IRAS and COBE satellites showed that infrared emission was strongest in regions where there is a high concentration of interstellar gas.

2.4.2 *Photo-dissociation*

For molecular clouds to survive, the molecules need to be protected from dissociation by high energy photons. Extreme UV (EUV) photons from hot O and B stars are unlikely to reach the molecules because they ionise atomic hydrogen. The EUV is also called the Lyman continuum, with energies hν $> 13.6\,\mathrm{eV}$, equivalent to wavelengths $\lambda < 912\,\text{Å}$, sufficient to photo-ionise even cold atomic hydrogen from its ground state. The wavelength $\lambda = 912\,\text{Å}$ is defined as the Lyman limit or 'edge'. The dramatic effect of this edge is evident in Fig. 2.3. The extreme UV creates H II regions around massive stars which completely absorb the EUV and re-radiate it at longer wavelengths. This shields the molecules from the EUV.

Molecules directly exposed to far-UV radiation, however, are also rapidly destroyed. Far UV radiation with energy in the range 5–13.6 eV (912–2000 A) is able to penetrate the skin of a cloud. The photons ionise atoms such as carbon and iron and photo-dissociate hydrogen molecules (but do not ionise). The transition layer between the exposed dissociated gas and the molecular interior is not smooth but probably very wrinkled and clumpy, allowing radiation to penetrate quite deep. We call this thick skin a *Photo-dissociation Region* or PDR.

A molecule is dissociated in two steps. It first undergoes electronic excitation by absorption of a resonant UV photon. Then, there is about a 10% probability that the molecule radiates into a state in the vibrational continuum, in which the two atoms are not bound and so fly apart. Alternatively, and most probably, dissociation does not result but the excited molecule returns to the ground electronic state where it cascades down through the vibrational and rotational energy states. This generates a 'fluorescence spectrum' of H_2 emission lines, characterised by many strong lines originating from high levels of vibration.

For molecules to build up they require their absorption lines to become optically thick, thus forming a self-protecting layer to the far UV. This is called self-shielding. In the absence of self-shielding, the radiation will still be attenuated by dust, and molecules will form when sufficient dust lies in the path of the radiation.

For H_2 self-shielding to be effective, however, we require only the equivalent of about $A_V \sim 0.1$. The CO molecule, being less abundant and so less effective at self-shielding, requires $A_V \sim 0.7$. Thus a dense cloud exposed to UV radiation will possess several layers of skin. The full list of layers is as follows:

- an outer ionised hydrogen region,
- a thin neutral atomic H layer,
- an outer H_2 layer containing C and O atoms,
- an inner H_2 layer containing CO, and
- a dark core containing H_2, CO and O_2.

Most of the PDR radiation originates from the warmed C and O atoms through their fine-structure emission lines (see §2.3.3).

2.4.3 *Hydrogen chemistry*

Molecules do not form readily in interstellar gas. Collisions between atoms are too fast for them to lose energy and become bound: a dynamical interaction time is typically just 10^{-13} s whereas the time scale to lose energy via radiation is at least 10^{-8} s. Furthermore, three-body collisions in which the third body could carry away the excess energy are extremely rare. Besides these problems, the mean UV photon flux in the Galaxy is of order 10^7 photons cm^{-2} s^{-1} and a molecule presents a cross-section of order 10^{-17} cm^2 to the passing photons. Hence, exposed molecules survive for periods of just 10^{10}s, just 300 yr.

Molecular hydrogen forms much more efficiently on the surfaces of grains. The main requirement is that one atom is retained on the grain surface until a second atom arrives and locates it. At low gas and grain temperatures and with a typical distribution of grain sizes, we usually take a rate of formation per unit volume of $3 \times 10^{-18}\, \mathrm{n\, n(H)\, T^{1/2}\, cm^{-3}\, s^{-1}}$. Hence, the time scale for H_2 formation, normalised to a typical molecular clump, is

$$t_F = 3 \times 10^6 \left(\frac{n}{10^3\ \mathrm{cm}^{-3}} \right)^{-1} \left(\frac{T}{10\ \mathrm{K}} \right)^{-1/2}\ \mathrm{yr}. \tag{2.6}$$

We will later find that this is also a typical cloud age.

High uncertainties, however, concern the nature of the dust and the limitations of laboratory experiments. In particular, the dust temperature, T_d, is a critical factor. If T_d is less than about 20 K, then H_2 formation might proceed at the above rapid rate. Above 20 K, the formation rate appears to decrease enormously. In comparison, observational estimates of dust temperatures in the diffuse unshielded ISM range from 13 K to 22 K. In star formation regions, however, the stellar radiation field provides considerable dust heating. An estimate of how rapidly the dust temperature is reduced with depth through a cloud surface is given in terms of the visual extinction A_V from the formula $T_d^5 \propto \exp(-1.66\,A_V)$.

The above H_2 formation time is still clearly insufficient to form molecular clouds until the cloud is shielded from the UV radiation. When sufficient, however, the cloud will become fully molecular with cosmic rays maintaining a tiny fraction of H atoms: balancing molecule destruction at the rate $\xi\, n(H_2)$ with the above formation rate yields a density of hydrogen atoms of $0.1\,\mathrm{cm}^{-3}$ within dark clouds, independent of the cloud density. Molecules will also be dissociated by UV radiation produced internally by collisions with the cosmic ray-induced electrons.

Molecular ions play a key part in cloud chemistry since they react fast even at low temperatures. H_3^+ is a stable ion produced by the cosmic rays. Along with CO^+, it is often important in the subsequent ion-molecule chemistry.

What if no dust is present, such as in primordial gas (§13.1.2) and probably also in protostellar winds and jets (§10.7)? Despite the above arguments, molecular hydrogen will still form in the gas phase under certain conditions. It is a two-step process. We first require H^- to be produced via the reaction $H + e \rightarrow H^- + h\nu$. Then, some of the H^- can follow the path $H^- + H \rightarrow H_2 + e$, although the H^- may be, in the meantime, neutralised by reactions with protons or the radiation field. A complete analysis shows that a rather high electron fraction or a paucity of dust is necessary for these reactions to assume importance. It should also be noted that in collimated protostellar winds, the density may be sufficient for three-body formation: $3H \rightarrow H_2 + H$.

Besides photo-dissociation discussed above, H_2 is destroyed in collisions with H_2 or other molecules and atoms. Each collisional dissociation removes 4.48 eV, the molecular binding energy. Hence dissociation also cools the gas while reformation may then heat the gas (although the reformed excited molecule may prefer to radiate away much of this energy before it can be redistributed).

Under what condition will hydrogen return to the atomic state simply through energetic collisions? During the collapse of a protostellar core, the gas warms up and molecule dissociation becomes the most important factor (§6.6). Approximately, the dissociation rate (probability of dissociation per second) takes the form $5 \times 10^{-9}\, \mathrm{n(H_2)} \exp(-52{,}000\,\mathrm{K/T})\ \mathrm{s}^{-1}$ (note that 4.48 eV has converted to 52,000 K). This equation holds for high density gas since it assumes that all energy levels are occupied according to collisions (thermal equilibrium) and not determined by radiation. Using this formula, if gas at a density of $10^6\,\mathrm{cm}^{-3}$ is heated to 1,400 K, it would still take 3×10^{18} s, to dissociate. That is much longer than the age of the Universe. At 2,000 K, however, the time required is down to just 3×10^{13} s, or a million years – comparable to the collapse time. Therefore, if energy released during a collapse is trapped in the gas, a critical point may be reached where molecule destruction is triggered. This is exactly what we believe occurs (see §6.6).

2.4.4 *Chemistry*

Astrochemistry provides a means to connect the structure and evolution of molecular clouds to the distributions and abundances of numerous molecular species. It usually requires the computational treatment of a complex network of chemical paths. Chemical models have now been developed to predict the existence and abundance of various molecules. The predicted abundances of some main species are in agreement with observed values but for many other molecules, especially heavy atom-bearing molecules and large polyatomic molecules, anomalies persist. Revisions to the models are ongoing, including improved treatments of the neutral-neutral reactions and grain-surface molecular depletion and desorption. As a result, the relative column densities of observed molecules can be employed to infer the chemical evolution and may serve as an indicator of cloud history.

We summarise here the facts most relevant to an overview of star formation. The chemistry can be broken up into three categories.

Shielded cloud chemistry applies if the gas is cold and photodissociation is low. This should be relevant to the dense inner regions of dark clouds. Cosmic rays still penetrate and they provide the catalyst for the chemical processing. Almost all the hydrogen is in molecular form so the reaction network begins with the cosmic ray production of H_2^+. At the rate of ionisation of $10^{-17}\ \zeta\ \mathrm{s}^{-1}$ from §2.2.2, the average H_2 is ionised roughly once every 10^9 yr. Once ionised, the H_2^+ quickly reacts with another H_2 to form H_3^+ and an H atom. This ion-neutral reaction proceeds swiftly, according to a so-called Langevin rate constant of 2×10^{-9} cm^3 s^{-1}. This implies that in a medium of density 1000 cm^{-3}, the average wait is several days to react to form H_3^+. Clearly, the initial ionisation is the rate-limiting step.

The chemistry now all turns on the H_3^+. The many processes include proton transfer with carbon and oxygen atoms (e.g. $C + H_3^+ \rightarrow CH^+ + H_2$), dissociative recombination (e.g. $HCO^+ + e \rightarrow CO^+ + H_2$, generating CO) and radiative association (e.g. $C^+ + H_2 \rightarrow CH_2^+ + \nu$, where e represents an electron and ν a photon. Heavy elements are also removed from the gas phase by condensation onto dust grains.

Exposed cloud chemistry applies to molecular regions of low extinction. Here, CO is photo-dissociated and the carbon is photo-ionised by starlight since it has an ionisation potential of 11.26 eV, below the Lyman limit. Therefore, C^+ is the most abundant ion with a fraction $n(C^+) \sim 10^{-4}\, n(H_2)$, which far exceeds the dense cloud level given by Eq. 2.2. The C^+ is converted to CO^+, which then is converted to HCO^+, which produces CO through dissociative recombination as noted above. The CO will take up all the available carbon when self-shielded to the ultraviolet radiation. Otherwise, the CO is rapidly dissociated through $CO + h\nu \rightarrow C + O$.

Warm molecular chemistry occurs when the gas is excited by collisions to temperatures of order 1,000 K. The sequence $O + H_2 \rightarrow OH + H$ - 4480 K and then $C + OH \rightarrow CO + H$ is very rapid in hot gas, usually converting C to CO before cooling reduces the temperature which would slow down the reactions. Note that the process is endothermic: collisional energy is required to make the OH.

Water molecules also follow from the OH through another endothermic reaction $OH + H_2 \rightarrow H_2O + H$ - 2100 K. In cold dense cores or collapsing cores, many molecules including H_2O accrete onto grains, forming icy mantles. Surface chemistry and radiation processes modify the composition.

The carbon monoxide molecule is the best tracer of star-forming gas on large scales for several reasons. First, it is rotationally excited even in gas of a few Kelvin. Second, it is formed very efficiently. Thirdly, it is not easily destroyed with a relatively high dissociation energy of 11.09 eV (electron Volts). Nevertheless, CO is also often depleted in dense regions, forming CO ice. It is quite commonly depleted by an order of magnitude or more, along with other species such as CS and HCO^+.

In the presence of a young interacting star, the ices are heated and molecules evaporated from the dust. This occurs in a sequence, according to their sublimation temperatures. Shock waves and turbulence will also strongly heat the cloud on larger scales, returning volatiles and radicals back into the gas phase through processes such as high-energy ion impacts, termed sputtering. Shattering also occurs through grain collisions. This tends to modify the distribution of grain masses.

It should also be mentioned that large enhancements of minor isotopes have been recorded in dense regions near young stars. Examples are DCN and DCO^+. These are explained as due to deuterium fractionation in which the deuterium initially in HD molecules undergoes a path of reactions.

2.4.5 *Cooling and heating*

The capacity of molecular clouds to cool is absolutely crucial to the mass of stars which form (see §4.3.4 on the Jeans Mass). Radiation is thought to be the chief carrier of the energy from a cloud (rather than conduction, for example). Photons are emitted from atoms and molecules as the result of spontaneous de-excitation. The particles are first excited by collisions, mainly through impacts with H_2, H, dust or electrons. If the density is high, however, the de-excitation will also be through an impact rather than through a photon, as explicitly shown in §2.3.2). This implies that radiative cooling is efficient at low densities but impacts redistribute the energy at high densities. We define a critical density to separate these two regimes (referred to in the literature as local thermodynamic equilibrium, or LTE, above, and non-LTE below). This concept is probably vital to primordial star formation (§13.1.3).

An important coolant is CO through rotational excitation as well as trace atoms C II at low densities and neutral C and O fine-structure excitation at high density. Although dominant by number, H_2 is not usually efficient at radiating away energy since it is a symmetric molecule in which a quadrupole moment is not effective.

In warm gas, the molecules CO, OH and H_2O also make considerable contributions to the cooling, and vibrational H_2 emission cools molecular gas after it has been strongly heated in shock waves to over 1,000 K. Rotational excitation of H_2 will also cool gas heated above 200 K.

A second important cooling path is available through gas-grain collisions. Provided the grains are cooler than the gas, collisions transfer energy to the dust. The dust grains are efficient radiators in the long wavelength continuum (in the infrared and submillimetre) and so the radiated energy escapes the cloud. Clearly, the dust is also a potential heating source through collisions if the dust can be kept warmer than the molecules by background radiation.

Cloud heating is provided by far ultraviolet photons. They eject electrons at high speed from grains. The excess energy is then thermalised in the gas. This process is appropriately called *photoelectric heating*. Deep inside clouds where the FUV cannot penetrate, heating is provided through the cosmic ray ionisations discussed above with perhaps 3.4 eV of heat deposited per ionisation.

Gas motions also heat the gas. The energy in subsonic turbulent motions is eventually channelled by viscosity into thermal energy after being broken down into small scale vortices (as discussed in §4.4). The energy of supersonic turbulence is dissipated more directly into heat by the creation of shock waves, described in §4.4. In addition, the gravitational energy released by cloud contraction produces compressive heating (or cooling in expanding regions).

There are many other potential heating mechanisms that we have already come across, depending on the state of the gas. These include direct photo-ionisation, collisional de-excitation of H_2 after UV pumping and H_2 formation.

2.5 Summary

This chapter has served three purposes: to provide backgrounds to the materials, to the means of observation and to the physical processes. Particular regard has been given to the processes relevant to molecular clouds. In fact, it is the large number and variety of physical processes which creates a fascinating story. While the cast of characters is small, they can take on many roles. Each molecule and atom has its own peculiarities and regime of importance, either physically, observationally or both.

Having explored the means by which light is emitted, absorbed and scattered, we are ready to appreciate and interpret the observations. The following chapters will then focus on the dominant interactions which describe the conditions suitable for star formation.