2 Galactic radiation fields

- The physical state of the gas in the ISM is determined in large part by the interaction of the gas and the dust with the interstellar radiation field (ISRF). Several types of interaction are important:

  - The chemical state of the gas (the ionization fraction, the balance between atomic and molecular gas, etc.) depends on the rate at which molecules are photodissociated and atoms are photoionized by the radiation field.

  - The thermal state of the gas depends on the photoionization rate, and also on the rate of a process known as **photoelectric heating**: the ejection of an energetic electron from a dust grain due to the absorption of a UV photon by the grain.

  - The thermal state of the dust is almost entirely determined by the balance between the absorption by the grains of radiation from the ISRF and the re-emission of this energy in the form of thermal radiation.

  - The radiation field can even have dynamical consequences, if it is strong enough: radiation pressure can accelerate dust grains, or even move gas around

2.1 Solar neighbourhood

- In the Solar neighbourhood, the ISRF is dominated by six components:

  - Galactic synchrotron emission from relativistic electrons
  - The Cosmic Microwave Background (CMB)
  - Infrared and far-infrared emission from dust grains heated by starlight
  - Bound-bound, bound-free and free-free emission from 10^4 K ionized plasma (sometimes referred to as **nebular emission**)
  - Starlight
  - X-rays from hot (10^5–10^8 K) plasma

- We will briefly discuss each of these components in turn, along with an estimate of their local energy density. These energy densities are also summarized in Table 1.

Synchrotron emission

- Synchrotron emission is emission produced by relativistic charged particles (typically electrons) that are accelerated radially (i.e. at right-angles to their direction of flight). It is the relativistic analogue of **cyclotron radiation**.

- Synchrotron emission is generated in the ISM by the interaction between relativistic electrons and the ambient magnetic field: the electrons spiral along the field lines, emitting synchrotron radiation as they do so.

Table 1: ENERGY DENSITIES IN DIFFERENT COMPONENTS OF THE ISRF

Component of ISRF	Energy density (erg cm^{-3})
Synchrotron	2.7×10^{-18}
CMB	4.19×10^{-13}
Dust	5.0×10^{-13}
Nebular emission (b-f, f-f)	4.5×10^{-15}
Nebular emission (Hα)	8×10^{-16}
Nebular emission (other b-b)	10^{-15}
Starlight, $T_1 = 3000$ K	4.29×10^{-13}
Starlight, $T_2 = 4000$ K	3.19×10^{-13}
Starlight, $T_3 = 7000$ K	2.29×10^{-13}
Starlight, power-law	7.11×10^{-14}
Starlight, total	1.05×10^{-12}
Soft X-rays	10^{-17}

- The angular frequency of a relativistic electron spiralling along a magnetic field is

$$\omega_{\mathrm{B}} = \frac{eB}{\gamma m_e c},$$ (1)

where B is the magnetic field strength and γ is the Lorentz factor.

- The power radiated by the electron is given by the Larmor equation

$$\frac{\mathrm{d}E}{\mathrm{d}t} = \frac{2e^2}{3c^3} \left| \ddot{\vec{x}} \right|^2$$ (2)

- In the absence of an electric field, the acceleration of the electron is simply due to the Lorentz force and is given by

$$\ddot{\vec{x}} = \frac{\gamma e}{m_e c} (\dot{\vec{x}} \times \vec{B}).$$ (3)

- Therefore, for an electron moving at an angle α relative to the magnetic field, we have

$$\left| \ddot{\vec{x}} \right|^2 = \left(\frac{e}{m_e c} \right)^2 (\gamma v B \sin \alpha)^2,$$ (4)

from which it follows that the power radiated is given by

$$\frac{\mathrm{d}E}{\mathrm{d}t} = 2c\sigma_{\mathrm{T}} \frac{B^2}{8\pi} \gamma^2 \beta^2 \sin^2 \alpha,$$ (5)

where $\beta = v/c$ and σ_{T} is the Thomson scattering cross-section[1].

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[1]Recall that this is given by

$$\sigma_{\mathrm{T}} = \frac{8\pi}{3} \left(\frac{e^2}{m_e c^2} \right)^2.$$ (6)

- Averaging over all possible pitch angles α, we eventually find that

$$\frac{\mathrm{d}E}{\mathrm{d}t} = \frac{4}{3}U_{\mathrm{B}}c\sigma_{\mathrm{T}}\gamma^2\beta^2, \tag{7}$$

where $U_{\mathrm{B}} = B^2/8\pi$ is the energy density of the magnetic field.

- So far, this is for a single electron with Lorentz factor γ. To get the total power radiated by a collection of electrons, we would also need to know the velocity distribution function of the electrons. However, even without knowing this, one important point is clear: the strength of the synchrotron emission depends on two main factors, the magnetic field strength and the local number density of relativistic electrons.

- As we expect both of these to vary within the Galaxy, it follows that the amount of synchrotron emission is also strongly variable. In particular, it is strong in regions such as supernova remnants, where there is a high local energy density in relativistic electrons, and where the gas compression also tends to increase B.

- Observations of Galactic synchrotron emission show that it has a roughly power-law spectrum,

$$U_\nu \propto \nu^{-1}. \tag{8}$$

Its total energy density is small, but nevertheless it dominates over the CMB at wavelengths $\lambda > 30$ cm (or frequencies $\nu < 1\,\mathrm{GHz}$, and is the main form of Galactic emission in the long wavelength radio-wave regime.

CMB

- This should be familiar to everyone from Cosmology, but just in case it isn't...

- The CMB is an almost perfect Planck spectrum, with temperature $T = 2.725$ K. It dominates the energy budget of the ISRF between $\lambda = 600\mu$m and $\lambda = 30$ cm. However, since the gas and dust in the ISM is typically warmer than the CMB, it doesn't play much of a role in the overall energy balance of the ISM.

- In high-redshift galaxies, the CMB temperature is much higher ($T_{\mathrm{CMB}} \propto (1 + z)$), and it plays a much more central role in determining the energy balance of the gas and dust.

IR and FIR emission from dust

- Infrared emission from dust dominates the ISRF spectrum between $\lambda = 5\mu$m and $\lambda = 600\mu$m.

- About two-thirds of the total power is radiated in the mid and far-infrared, at $\lambda > 50\mu$m. This emission is largely in the form of thermal emission from dust grains: the spectrum is close to a black-body, and the corresponding mean dust temperature is around 17 K.

- The remaining one-third of the dust emission is largely concentrated in a series of distinct peaks at wavelengths $\lambda = 3.3, 6.2, 7.7, 8.6, 11.3$ and 12.7μm. These peaks correspond to vibrational emission bands produced by so-called **polycyclic aromatic hydrocarbons**, or PAHs for short: large organic molecules, containing one or more benzene rings (hence 'aromatic').

- PAH emission works rather differently than is the case for small grains. Because PAHs are so small, the rate at which they absorb photons is small, but each photon causes a significant change in the internal energy of the molecule. Their "temperature" therefore varies greatly with time – they are very hot (i.e. have lots of internal energy) immediately after they absorb a photon, but spend much of their time being very cold.

- Physically, what happens is actually a form of fluoresence – the PAHs absorb UV photons, putting them into a highly excited state, and then cascade back to the ground state via a large number of infra-red transitions.

- An important implication of this is that PAH emission dominates at short wavelengths (i.e. in the near and mid-IR) unless the other grains are also very hot. Since the strength of the PAH emission depends on the strength of the UV radiation field, it is a useful tracer of the formation of massive stars.

Nebular emission

- There are three main types of emission produced within ionized regions: free-free, free-bound and bound-bound emission.

- **Free-free** emission (aka **Bremsstrahlung**) is electromagnetic radiation produced by a charged particle (usually an electron) that is accelerated or decelerated by the electric field of another charged particle.

- The specific emissivity due to Bremsstrahlung of a thermal plasma composed of electrons and ions with charge Z can be written as

$$j_\nu = \frac{32\pi^2 Z^2 n^2 e^6}{3\sqrt{3}m_e^2 c^3}\sqrt{\frac{2m_e kT}{\pi}}g_{\text{ff}}(v,\nu)e^{-h\nu/kT}, \tag{9}$$

where g_{ff} is the **gaunt factor**, which is of order unity and which depends only logarithmically on the electron velocity v and the frequency ν.

- There are several important points to note here:

  - Because g_{ff} depends only weakly on ν, the Bremsstrahlung spectrum is largely flat at $h\nu \ll kT$, and falls off exponentially at $h\nu > kT$. The range of frequencies produced therefore depends on the temperature of the plasma. Gas which is photoionized by massive stars typically has a temperature $T \sim 10^4$ K (we will explore why in the next section), and so the most energetic Bremsstrahlung photons that it produces will typically have energies of around 1 eV, or wavelengths

of around 1μm. Hot, collisionally ionized plasma, on the other hand, may have a temperature as high as 10^8 K, and can easily produce Bremsstrahlung photons with X-ray energies.

- The total emissivity (i.e. j_ν integrated over all frequencies) scales as $n^2\sqrt{T}$, so hot dense plasmas produce much more emission than cool, diffuse plasmas.

- The emissivity also scales as Z^2, and so in very highly ionized gas, the contribution from metals can be significant even though they have small abundances relative to hydrogen or helium.

- **Bound-free** emission is produced by the recombination of ionized hydrogen and helium.

- When an H^+ ion and an electron recombine, the energy of the photon emitted depends upon the quantum state of the resulting hydrogen atom and the kinetic energy of the electron in the rest-frame of the ion. For recombination into a state with principle quantum number n, we have

$$h\nu = \left(1 - \frac{n^2 - 1}{n^2}\right) I_\mathrm{H} + E_\mathrm{kin,e}, \tag{10}$$

where $I_\mathrm{H} = 13.6$ eV is the ionization energy of hydrogen and $E_\mathrm{kin,e}$ is the kinetic energy of the electron. Similar equations can also be constructed for He^+ and He^{++} recombination.

- An important point to note here is that because of the $E_\mathrm{kin,e}$ term, the spectrum of bound-free emission consists of a continuum (albeit not a particularly flat one), rather than a series of distinct spectral lines.

- It is also worth noting that although the mean kinetic energy of the electrons in the plasma is simply $\frac{3}{2}kT$, the mean value of $E_\mathrm{kin,e}$ is slightly lower, because lower energy electrons are more likely to be involved in recombination.

- Finally, **bound-bound** emission is the radiation produced by transitions from one atomic bound state to another. There are two main effects that give rise to this emission. Firstly, these photons can be produced by collisional excitation (typically by collisions with electrons) followed by radiative de-excitation, i.e.

$$\mathrm{H} + \mathrm{e}^- \rightarrow \mathrm{H}^* \rightarrow \mathrm{H} + \gamma, \tag{11}$$

where H^* denotes an excited state.

- One important reason that electron collisions dominate here is that they have much larger velocities than protons or neutral hydrogen atoms:

$$\frac{v_\mathrm{e}}{v_\mathrm{H}} = \left(\frac{m_\mathrm{H}}{m_\mathrm{e}}\right)^{1/2} \simeq 43, \tag{12}$$

where m_e and m_H are the electron mass and the proton mass, respectively. Since the excitation rate coefficient is an average of the product σv over velocity, higher mean velocities lead to higher excitation rates, everything else being equal.

- The critical densities for the allowed transitions of H and He are very large, and so the vast majority of the atoms are in their ground state. Collisional excitation therefore produces Lyman series photons, plus some number of lower-energy photons corresponding to transitions between states with higher n.

- The other main source of bound-bound emission is recombination. When an ion and electron recombine, the final state of the resulting atom is typically not the ground state, and hence bound-bound emission is produced as the atoms cascade to the ground state.

- Detailed calculations of the relative number of different types of bound-bound photons are possible; the results typically depend on the gas temperature and the density. If you're interested in finding out more, then Hummer & Storey (1987, MNRAS, 224, 801) is a good place to start.

Starlight

- Stars produce energy primarily in the near infrared, visible and soft ultraviolet. However, in neutral regions of the ISM, stellar photons with energies greater than the ionization energy of hydrogen, 13.6 eV, are largely absent – they are absorbed by hydrogen atoms, ionizing them, and hence cannot penetrate deeply into neutral regions.

- Mathis et al. (1983, A&A, 128, 212) showed that in the solar neighbourhood, the starlight component of the ISRF could be represented at long wavelengths as the sum of three diluted black-body spectra. At wavelengths $\lambda > 2450$ Å, we have

$$\nu u_\nu = \sum_{i=1}^{3} \frac{8\pi h \nu^4}{c^3} \frac{W_i}{e^{h\nu/kT_i} - 1}, \tag{13}$$

where W_i and T_i are the dilution factor and temperature of each component, with

$$T_1 = 3000 \,\text{K}, \quad W_1 = 7.0 \times 10^{-13} \tag{14}$$
$$T_2 = 4000 \,\text{K}, \quad W_2 = 1.65 \times 10^{-13} \tag{15}$$
$$T_3 = 7500 \,\text{K}, \quad W_3 = 1.0 \times 10^{-14} \tag{16}$$

- At wavelengths $\lambda < 2450$ Å, the starlight contribution to the ISRF has been estimated by a number of authors. The earliest widely-cited estimate was made by Habing (1968). He estimated that $\nu u_\nu \simeq 4 \times 10^{-14} \,\text{erg cm}^{-3}$ at $\lambda = 1000$ Å, corresponding to a photon energy of 12.4 eV. It is often convenient to reference other estimates to this value, which we do via the dimensionless parameter

$$\chi \equiv \frac{(\nu u_\nu)_{1000\text{Å}}}{4 \times 10^{-14} \,\text{erg cm}^{-3}}. \tag{17}$$

- Alternatively, we can reference other estimates to the Habing field by comparing the total energy density in the range 6–13.6 eV. In this case, we define a different dimensionless parameter

$$G_0 \equiv \frac{u(6\text{--}13.6 \text{ eV})}{5.29 \times 10^{-14} \text{ erg cm}^{-3}}. \tag{18}$$

- If we are interested in e.g. the photodissociation of H_2 or CO, which requires photons with energies > 10 eV, then χ is the appropriate parameter to use. On the other hand, if we are interested in e.g. the photoelectric heating rate, which is sensitive to a wider range of photon energies, then G_0 is more appropriate.

- Two other estimates of the UV portion of the ISRF are in widespread use: one due to Draine (1978) and the other due to Mathis et al. (1983).

- Draine (1978) fit the field with a polynomial function:

$$\lambda u_\lambda = 6.84 \times 10^{-14} \lambda_3^{-5} \left(31.016 \lambda_3^2 - 49.913 \lambda_3 + 19.897 \right) \text{ erg cm}^{-3}, \tag{19}$$

where $\lambda_3 \equiv \lambda/1000$ Å. This field has a normalization, relative to the Habing field, of $\chi = 1.71$ and $G_0 = 1.69$.

- Mathis et al. (1983; henceforth MMP83) used instead a broken power-law fit:

$$\nu u_\nu = \begin{cases} 2.373 \times 10^{-14} \lambda^{-0.6678} & 1340\text{--}2450 \text{ Å} \\ 6.825 \times 10^{-13} \lambda & 1100\text{--}1340 \text{ Å} \\ 1.287 \times 10^{-9} \lambda^{4.4172} & 912\text{--}1100 \text{ Å}. \end{cases} \tag{20}$$

Here, all wavelengths are in μm, and the energy densities are in units of erg cm^{-3}. This estimate has $\chi = 1.23$ and $G_0 = 1.14$. The available observational evidence suggests that the MMP83 estimate is a better one than that of Draine (1978), but the latter is probably in wider use in models of the ISM.

Thermal X-rays

- Supernovae and other highly energetic events can produce pockets of very hot, ionized plasma in the ISM, with temperatures upwards of $T = 10^6$ K.

- This plasma cools both via line emission from highly ionized metals (H and He are fully ionized at these temperatures), and also Bremsstrahlung, which dominates at the highest temperatures.

- Much of this emission is in the form of X-ray photons. For example, in the case of Bremsstrahlung, the energy emitted per logarithmic frequency bin, νj_ν, is highest around the exponential cut-off in the Bremsstrahlung spectrum, i.e. around $h\nu \sim kT$. For $T = 10^6$ K, this corresponds to a photon energy of around 100 eV, in the extreme UV, while temperatures a factor of a few higher lead to emission that is primarily in the form of soft X-rays.[2]

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[2] "Soft" X-rays are conventionally those with energies $E < 2 \text{ keV}$.

- Below 2keV, only a relatively small column density of gas is required in order to absorb these X-rays, and so the local X-ray flux is highly variable in the ISM – high in diffuse gas near supernova remnants, or in the Galactic Centre, low in denser regions.

- X-rays contribute only a very small amount of the total energy density in radiation, but are important for their chemical effects – they can produce ionization in regions where UV photons cannot penetrate. Their role is rather similar to that of cosmic rays, discussed in more detail below.

2.2 Photodissociation regions

- Close to bright, massive stars, the UV radiation from the star dominates over the interstellar background. This radiation can ionize the gas, creating what is known as an **HII region**. (We will talk a lot more about the physics of these regions in later lectures).

- Photons with energies below the hydrogen ionization threshold can escape from these HII regions and photodissociate molecules such as H_2 and CO in the surrounding gas. Regions of the ISM thus affected are known as **photodissociation regions**, or PDRs for short.

- We can derive some rough estimates for the typical size of a PDR and the strength of the radiation field within it by considering a simple toy model for our massive star. If we assume that we can model the star as a black-body with effective temperature T_{eff}, then its integrated luminosity in the wavelength range $912 < \lambda < 2450$ Å is simply

$$L_{\text{UV}} = \frac{\int_{\lambda_1}^{\lambda_2} B_\lambda(T_{\text{eff}})\mathrm{d}\lambda}{\int_0^\infty B_\lambda(T_{\text{eff}})\mathrm{d}\lambda} L_{\text{bol}}, \tag{21}$$

where $\lambda_1 = 912$ Å, $\lambda_2 = 2450$ Å, L_{bol} is the bolometric luminosity of the star, B_λ is the Planck function, written in terms of the flux per unit wavelength, i.e.

$$B_\lambda = \frac{2hc^2}{\lambda^5} \frac{1}{\exp(hc/\lambda kT) - 1}. \tag{22}$$

- For a 20 $M_\odot$ O-type star, we have a bolometric luminosity $L_{\text{bol}} \simeq 54000\ L_\odot$, and an effective temperature $T_{\text{eff}} \simeq 30000$ K. Evaluating the ratio of integrals, we find that

$$L_{\text{UV}} \simeq 0.245 \times 54000\ L_\odot \simeq 13000\ L_\odot. \tag{23}$$

- The flux at a distance R from the star is therefore

$$F_{\text{UV}} = \frac{L_{\text{UV}}}{4\pi R^2} \simeq 0.4 \left(\frac{R}{1\ \text{pc}}\right)^{-2}\ \text{erg s}^{-1}\ \text{cm}^{-2}, \tag{24}$$

and the associated energy density is

$$U_{\text{rad}} = \frac{F_{\text{UV}}}{c} \simeq 1.4 \times 10^{-11}\ \text{erg cm}^{-3}. \tag{25}$$

- Comparing this with the radiation energy density quoted above for the ISRF between 912 Å and 2450 Å, we see that the radiation from our local massive star dominates at distances

$$R < \left(\frac{1.4 \times 10^{-11}}{7.11 \times 10^{-14}} \right)^2 \simeq 14 \, \text{pc.} \tag{26}$$

  Close to the star, the UV energy density can be hundreds or thousands times larger than the mean value in the ISRF.

- Our estimate for the energy density scales only weakly with T_{eff}, provided that the star is hot enough to produce significant amounts of radiation at $\lambda < 2450$ Å. On the other hand, it scales strongly with mass, via the stellar mass-luminosity relationship, which is approximately $L \propto M^{3.5}$ for stellar masses in the range $2 < M < 20 \, \text{M}_\odot$ and $L \propto M$ for stellar masses $M > 20 \, \text{M}_\odot$.

- It is important to note that so far we have assumed that the gas is optically thin to the star's radiation. In the low density, diffuse ISM, this is a reasonable assumption on the scales of interest, but this is not the case if the massive star remains embedded in the dense molecular cloud in which it formed. In this case, the size of the PDR around the star is much smaller, as most of the UV radiation is absorbed locally within the cloud.

3 Cosmic rays

- As well as photons, interstellar gas is also irradiated by so-called **cosmic rays**.

- These are high energy, relativistic particles – primarily protons, but with about 10% being alpha particles and $\sim 1\%$ being metal nuclei.

- Their energy spans a wide range, from 100 MeV up to more than 1 TeV. The total energy density in cosmic rays is approximately $2 \, \text{eV cm}^{-3}$, within a factor of a few of the mean thermal energy density of the ISM. Cosmic rays therefore play an important role in the overall energy balance of the gas.

- All but the most highly energetic cosmic rays are tied to the magnetic field of the galaxy and therefore scatter repeatedly within the disk, leading to the local energy density in CRs being relatively uniform.

- The spectrum of the cosmic rays (i.e. the flux per unit energy) decreases sharply with increasing energy, and so the majority of the heating and ionization that they provide comes from the least energetic, 100 MeV, cosmic rays.

- Unfortunately, it is precisely this part of the cosmic ray spectrum that we know the least about. At this energy, cosmic rays are unable to penetrate within the heliosphere, owing to interactions with the solar wind. Our determination of the cosmic ray ionization rate is therefore indirect, based on chemical constraints.

- As an example, consider the case of H_3^+, which is formed in the diffuse ISM via:

$$H_2 + CR \rightarrow H_2^+ + e^-, \tag{27}$$
$$H_2^+ + H_2 \rightarrow H_3^+ + H, \tag{28}$$

  where the first reaction is the rate-limiting step, and which is then destroyed by dissociative recombination:

$$H_3^+ + e^- \rightarrow products. \tag{29}$$

- In a diffuse cloud, with $n_{H_2} > n_H$, the equilibrium number density of H_3^+ that results from these reactions is given approximately by

$$n_{H_3^+} = \frac{\zeta_{H_2}}{k_{dr}} \frac{n_{H_2}}{n_e}, \tag{30}$$

  where k_{dr} is the dissociative recombination rate coefficient and ζ_{H_2} is the cosmic ray ionization rate of H_2.

- If we assume that the temperature and the values of the H_2-to-electron ratio do not vary greatly along the line of sight, then we can convert this expression to an expression relating the column densities of H_2, H_3^+ and electrons:

$$N_{H_3^+} = \frac{\zeta_{H_2}}{k_{dr}} \frac{N_{H_2}}{N_e}. \tag{31}$$

- Next we note that within diffuse molecular clouds, the main source of electrons is ionized carbon, C^+. We therefore assume that $N_{C^+} = N_e$, allowing use to write Equation 31 as:

$$N_{H_3^+} = \frac{\zeta_{H_2}}{k_{dr}} \frac{N_{H_2}}{N_{C^+}}. \tag{32}$$

  Finally, this can be rearranged to give:

$$\zeta_{H_2} = \frac{N_{H_3^+} N_{C^+} k_{dr}}{N_{H_2}}. \tag{33}$$

- All of the column densities on the right-hand side can be measured using absorption line techniques, which also constrain the kinetic temperature of the gas. Therefore, provided that the value of k_{dr} is known, the cosmic ray ionization rate of H_2 can be inferred.

- Applying this technique to the diffuse ISM, we find that $\zeta_{H_2} \sim 10^{-16}$ s^{-1} (or possibly larger along some sight-lines).

- The method outlined here only works in diffuse gas (as in dense, well-shielded gas, we can't assume that $N_e \sim N_{C^+}$), but similar techniques that relate the cosmic ray ionization rate to the column densities of two or more observable species can be constructed that work in dense, shielded gas. However, these techniques typically yield values of a

few times $10^{-17}\,\mathrm{s}^{-1}$ for the cosmic ray ionization rate in these regions. This difference appears to be real (i.e. it is not just a consequence of our simplifying assumptions), but the reason for it is currently not known. It may have something to do with **magnetic shielding** of denser gas, i.e. the reflection of cosmic rays out of dense regions owing to their interaction with the magnetic field there (which will generally be stronger than in the more diffuse gas).

- Finally, ionization of gas by cosmic rays is also associated with heating: the electrons produced during the ionization process typically have excess kinetic energy, which is then converted to heat. Typically, around 20 eV of heat is deposited for each ionization, and so the CR heating rate is given by

$$\Gamma_{\mathrm{CR}} \sim 3 \times 10^{-28}\,\mathrm{erg\,s}^{-1}. \tag{34}$$

This process is unimportant in comparison to photoelectric heating in low extinction regions, but can dominate the heating of dense, high extinction gas (particularly in regions with higher cosmic ray ionization rates).