Astro 735: Cosmology
Lecture 9: The early universe and primordial nucleosynthesis

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The early universe was hot. We have seen that the temperature decreases as $T \propto a^{-1}$, and that matter-radiation equality occurred at $a_{r,m} = 3 \times 10^{-4}$, corresponding to a redshift $z_{r.,m} = 3270$. Before this, radiation dominates over matter, and $a \propto t^{1/2}$. We can write down simple expressions for the temperature in this era:

$$T(t) \approx 10^{10}\ \mathrm{K} \left(\frac{t}{1\ \mathrm{s}}\right)^{-1/2} \tag{1}$$

or

$$kT(t) \approx 1\ \mathrm{MeV} \left(\frac{t}{1\ \mathrm{s}}\right)^{-1/2}. \tag{2}$$

The mean energy per photon is

$$E_{\mathrm{mean}}(t) \approx 2.7kT(t) \approx 3\ \mathrm{MeV} \left(\frac{t}{1\ \mathrm{s}}\right)^{-1/2}. \tag{3}$$

(Dividing the energy density by the number density of blackbody radiation gives a mean photon energy $E_{\mathrm{mean}} \simeq 2.70kT$.)

- Abundances and distributions of many types of particles depend on how the reactions that produce them were affected by the expansion and cooling of the universe.
- The Hubble time $t_H = 1/H_0$ sets the timescale for temperature and density changes due to expansion, so for particles to be in thermal equilibrium, their reaction rate per particle Γ must be greater than the expansion rate,
$$\Gamma \gg \frac{\dot{a}}{a}. \tag{4}$$
- When the reaction rate becomes lower than the expansion rate, particles can no longer interact, and their abundances and distributions are "frozen;" this process is also called "freeze out." We will see how this determines the abundances of the light elements.
- Energy scale for nuclear fusion and fission is that of the typical binding energy per nucleon, ~ 8 MeV.
- A plot of the binding energy per nucleon as a function of the number of nucleons is shown in Figure 1. Note that deuterium (^{2}H or D) is not very tightly bound, and that ^{4}He is tightly bound compared to other light nuclei. (The peak of the curve is ^{56}Fe, which accounts for the fact that massive stars cannot get energy from fusing elements heavier than iron, resulting in Type II core-collapse supernovae.)
- The binding energy of deuterium is 2.22 MeV, and when the temperature of the universe dropped sufficiently far below this value protons and neutrons were able to fuse to form deuterium, beginning the process of Big Bang nucleosynthesis. Primordial nucleosynthesis occurred in the first few minutes of the universe.
- The first thing we can observe is that it was very inefficient. From an energy point of view, we would expect most nucleons to be in their most bound states, in which case the baryonic matter would be an iron-nickel alloy. This obviously isn't the case; about 75% of the baryons are unbound protons (^{1}H),

and most of the remainder is ^{4}He. The primordial helium fraction of the universe can be written as the mass density of ^{4}He divided by the total density of baryonic matter,

$$Y_p \equiv \frac{\rho(^4\mathrm{He})}{\rho_b}. \tag{5}$$

(Elemental abundances are commonly written as X, Y and Z, where X is the hydrogen fraction by mass, Y is the helium fraction, and Z is the fraction of all heavier elements.) The minimum observed value of Y in astronomical objects is $Y = 0.24$.

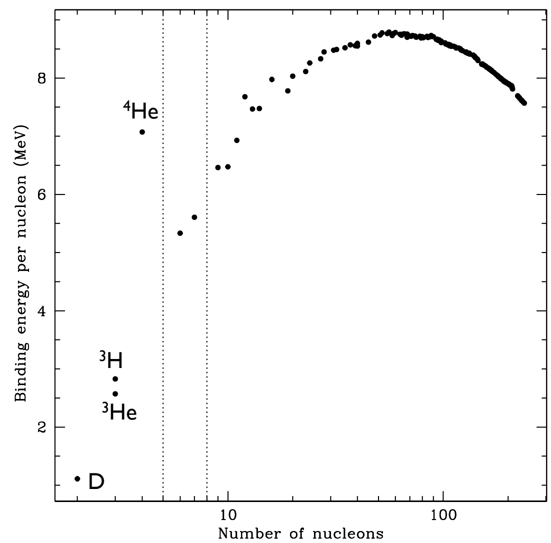

Figure 1: Binding energy per nucleon as a function of the number of nucleons (A). Nuclei important in the early universe are marked, and note the lack of stable nuclei with $A = 5$ and $A = 8$.

1 Neutrons and protons

The basic building blocks for nucleosynthesis are neutrons and protons. The rest energy of a neutron is greater than that of a proton by a factor

$$Q_n = m_n c^2 - m_p c^2 = 1.29 \text{ MeV} \tag{6}$$

A free neutron is unstable, decaying via the reaction

$$n \to p + e^- + \bar{\nu}_e. \tag{7}$$

- Decay time for a free neutron is $\tau_n = 890$ s (if you start out with a population of free neutrons, after a time t, a fraction $f = \exp(-t/\tau_n)$ will remain). Since the energy Q_n released by the decay of a neutron into a proton is greater than the rest energy of an electron ($m_e c^2 = 0.51$ MeV), the remainder of the energy is carried away by the kinetic energy of the electron and the energy of the electron anti-neutrino.

- This decay time is short; once the universe was several hours old, it contained essentially no free neutrons. However, a neutron which is bound into a stable atomic nucleus won't decay, which is why there are still neutrons around today.

- Consider the state of the universe when its age is $t = 0.1$ s. At that time, the temperature was $T \approx 3 \times 10^{10}$ K, and the mean energy per photon was $E_{\rm mean} \approx 10$ MeV. This energy is much greater than the rest energy of a electron or positron, so there were positrons as well as electrons present, created by pair production:

$$\gamma + \gamma \rightleftharpoons e^- + e^+ \tag{8}$$

- At $t = 0.1$ s, neutrons and protons were in equilibrium with each other, via the interactions

$$n + \nu_e \rightleftharpoons p + e^- \tag{9}$$

and

$$n + e^+ \rightleftharpoons p + \bar{\nu}_e \tag{10}$$

- As long as neutrons and protons are kept in equilibrium by these reactions, their number density is given by the **Maxwell-Boltzmann equation**. So the number density of neutrons is

$$n_n = g_n \left(\frac{m_n kT}{2\pi\hbar^2}\right)^{3/2} \exp\left(-\frac{m_n c^2}{kT}\right) \tag{11}$$

and the number density of protons is

$$n_p = g_p \left(\frac{m_p kT}{2\pi\hbar^2}\right)^{3/2} \exp\left(-\frac{m_p c^2}{kT}\right), \tag{12}$$

where g_n and g_p are the statistical weights. The statistical weights are equal, with $g_p = g_n = 2$, so the neutron-to-proton ratio is

$$\frac{n_n}{n_p} = \left(\frac{m_n}{m_p}\right)^{3/2} \exp\left(-\frac{(m_n - m_p)c^2}{kT}\right). \tag{13}$$

- We can simplify this. First, $(m_n/m_p)^{3/2} = 1.002$, and we won't lose much in accuracy if we set this factor equal to one. Second, the difference in rest energy of the neutron and proton is $(m_n - m_p)c^2 = Q_n = 1.29$ MeV. Thus, the equilibrium neutron-to-proton ratio is just

$$\frac{n_n}{n_p} = \exp\left(-\frac{Q_n}{kT}\right). \tag{14}$$

- At temperatures $kT \gg Q_n = 1.29$ MeV, corresponding to $T \gg 1.5 \times 10^{10}$ K and $t \ll 1$ s, the number of neutrons is nearly equal to the number of protons. However, as the temperature starts to drop below 1.5×10^{10} K, protons begin to be strongly favored, and the neutron-to-proton ratio drops exponentially.

- If the neutrons and protons remained in equilibrium, then by the time the universe was six minutes old, there would be only one neutron for every million protons. However, neutrons and protons do not remain in equilibrium for nearly that long.

- Interactions between neutrons and protons in the early universe (Equations 9 and 10) involve the interaction of a baryon with a neutrino (or anti-neutrino). Neutrinos interact with baryons via the weak nuclear force. The cross-sections for weak interactions have the temperature dependence $\sigma_w \propto T^2$;

at the temperatures we are considering, the cross-sections are small. A typical cross-section for the interaction of a neutrino with any other particle via the weak nuclear force is

$$\sigma_w \sim 10^{-47}\ \mathrm{m}^2 \left(\frac{kT}{1\ \mathrm{MeV}}\right)^2 . \tag{15}$$

(Compare this to the Thomson cross-section for the interaction of electrons via the electromagnetic force, $\sigma_e = 6.65 \times 10^{-29}\ \mathrm{m}^2$.)

- In the radiation-dominated universe, the temperature falls at the rate $T \propto a(t)^{-1} \propto t^{-1/2}$, and so the cross-sections for weak interactions diminish at the rate $\sigma_w \propto t^{-1}$. The number density of neutrinos falls at the rate $n_\nu \propto a(t)^{-3} \propto t^{-3/2}$, so the rate Γ with which neutrons and protons interact with neutrinos via the weak force falls rapidly:

$$\Gamma = n_\nu c \sigma_w \propto t^{-5/2}. \tag{16}$$

- Meanwhile, the Hubble parameter is only decreasing at the rate $H \propto t^{-1}$. When $\Gamma \approx H$, the neutrinos decouple from the neutrons and protons, and the ratio of neutrons to protons is frozen (at least until the neutrons start to decay, at times $t \sim \tau_n$).

- An exact calculation of the temperature T_{freeze} at which $\Gamma = H$ requires a knowledge of the exact cross-section of the proton and neutron for weak interactions. Using the best available laboratory information, the freezeout temperature turns out to be $kT_{\mathrm{freeze}} = 0.8$ MeV, or $T_{\mathrm{freeze}} = 9 \times 10^9$ K. The universe reaches this temperature when its age is $t_{\mathrm{freeze}} \sim 1$ s. The neutron-to-proton ratio, once the temperature drops below T_{freeze}, is frozen at the value

$$\frac{n_n}{n_p} = \exp\left(-\frac{Q_n}{kT_{\mathrm{freeze}}}\right) \approx \exp\left(-\frac{1.29\ \mathrm{MeV}}{0.8\ \mathrm{MeV}}\right) \approx 0.2. \tag{17}$$

- So at times $t_{\mathrm{freeze}} < t \ll \tau_n$, there was one neutron for every five protons in the universe.

- The scarcity of neutrons relative to protons explains why Big Bang Nucleosynthesis was so incomplete, leaving three-fourths of the baryons in the form of unfused protons.

- A neutron will fuse with a proton much more readily than a proton will fuse with another proton. When two protons fuse to form a deuterium nucleus, a positron must be emitted (to conserve charge); this means that an electron neutrino must also be emitted (to conserve electron quantum number). The proton-proton fusion reaction can be written as

$$p + p \rightarrow \mathrm{D} + e^+ + \nu_e. \tag{18}$$

- The involvement of a neutrino in this reaction tells us that it involves the weak nuclear force, and thus has a minuscule cross-section, of order σ_w.

- By contrast, the neutron-proton fusion reaction is

$$p + n \leftrightharpoons \mathrm{D} + \gamma. \tag{19}$$

No neutrinos are involved; this is a strong interaction (one involving the strong nuclear force). The cross-section for interactions involving the strong nuclear force are much larger than for those involving the weak nuclear force.

- The rate of proton-proton fusion is much slower than the rate of neutron-proton fusion, for two reasons. First, the cross-section for proton-proton fusion, since it is a weak interaction, is extremely small compared to the cross-section for neutron-proton fusion. Second, since protons are all positively charged, they must surmount the Coulomb barrier between them in order to fuse.

- Two protons can fuse, of course, as in the Sun. However, fusion in the Sun is a very slow process. If you pick out any particular proton in the Sun's core, it has only one chance in ten billion of undergoing fusion during the next year. Only exceptionally fast protons have any chance of undergoing fusion, and even those high-speed protons have only a tiny probability of quantum tunneling through the Coulomb barrier of another proton and fusing with it. The core of the Sun, though, is a stable environment; it's in hydrostatic equilibrium, and its temperature and density change only slowly with time.

- In the early universe, by strong contrast, the temperature drops as $T \propto t^{-1/2}$ and the density of baryons drops as $n_b \propto t^{-3/2}$. Big Bang Nucleosynthesis is a race against time. After less than an hour, the temperature and density have dropped too low for fusion to occur.

- Note also that the rate of neutron-neutron fusion in the early universe is negligibly small compared to the rate of neutron-proton fusion. The reaction governing neutron-neutron fusion is

$$n + n \rightarrow \mathrm{D} + e^{+} + \bar{\nu}_e. \tag{20}$$

- Again, the presence of a neutrino (an electron anti-neutrino, in this case) tells us this is an interaction involving the weak nuclear force. Thus, although there is no Coulomb barrier between neutrons, the neutron-neutron fusion rate is tiny. In part, this is because of the scarcity of neutrons relative to protons, but mainly it is because of the small cross-section for neutron-neutron fusion.

- Given the speed of neutron-proton fusion when compared to the leisurely rate of proton-proton and neutron-neutron fusion, we can state, as a lowest order approximation, that BBN proceeds until every free neutron is bonded into an atomic nucleus, with the leftover protons remaining solitary. In this approximation, we can compute the maximum possible value of Y_p, the fraction of the baryon mass in the form of ^{4}He.

- To compute the maximum possible value of Y_p, suppose that every neutron present after the proton-neutron freezeout is incorporated into a ^{4}He nucleus. Given a neutron-to-proton ratio of $n_n/n_p = 1/5$, we can consider a representative group of 2 neutrons and 10 protons. The 2 neutrons can fuse with 2 of the protons to form a single ^{4}He nucleus. The remaining 8 protons, though, will remain unfused. The mass fraction of ^{4}He will then be $Y_{\max} = 4/12 = 1/3$. More generally, if $f = n_n/n_p$, with $0 \leq f \leq 1$, then the maximum possible value of Y_p is $Y_{\max} = 2f/(1+f)$.

- If the observed value of $Y_p = 0.24$ were greater than the predicted $Y_{\max}$, that would be a cause for worry; it might mean, for example, that we didn't really understand the process of proton-neutron freezeout.

- However, the fact that the observed value of Y_p is less than $Y_{\max}$ is not worrisome; there are various factors which act to reduce the actual value of Y_p below its theoretical maximum.
  - If nucleosynthesis didn't take place immediately after freezeout at $t \approx 1$ s, then the spontaneous decay of neutrons would inevitably lower the neutron-to-proton ratio, and thus reduce the amount of ^{4}He produced.
  - If some neutrons escape fusion altogether, or end in nuclei lighter than ^{4}He (such as D or ^{3}He), they will not contribute to Y_p.

  - If nucleosynthesis goes on long enough to produce nuclei heavier than ^{4}He, that too will reduce Y_p.

- In order to compute Y_p accurately, as well as the abundances of other isotopes, it will be necessary to consider the process of nuclear fusion in more detail.

2 Deuterium synthesis

Let's move on to the next stage of Big Bang Nucleosynthesis, just after proton-neutron freezeout is complete.

- The time is $t \approx 2$ s. The neutron-to-proton ratio is $n_n/n_p = 0.2$.
- The neutrinos, which ceased to interact with electrons about the same time they stopped interacting with neutrons and protons, are now decoupled from the rest of the universe.
- The photons, however, are still strongly coupled to the protons and neutrons. Big Bang Nucleosynthesis takes place through a series of two-body reactions, building heavier nuclei step by step.
- The essential first step in BBN is the fusion of a proton and a neutron to form a deuterium nucleus:

$$p + n \leftrightharpoons \mathrm{D} + \gamma. \tag{21}$$

  The energy released in this reaction and carried away by the gamma ray is the binding energy of a deuterium nucleus,

$$B_D = (m_n + m_p - m_D)c^2 = 2.22\,\mathrm{MeV}. \tag{22}$$

- From the Boltzmann equations for the number densities of free protons, free neutrons and deuterium nuclei, we can derive an expression for their relative numbers:

$$\frac{n_D}{n_p n_n} = \frac{g_D}{g_p g_n}\left(\frac{m_D}{m_p m_n}\right)^{3/2}\left(\frac{kT}{2\pi\hbar^2}\right)^{-3/2}\exp\left(\frac{[m_p + m_n - m_D]c^2}{kT}\right) \tag{23}$$

- Make the substitution $[m_p + m_n - m_D]c^2 = B_D$. The statistical weight factor of the deuterium nucleus is $g_D = 3$, in comparison to $g_p = g_n = 2$ for a proton or neutron. To acceptable accuracy, we can also write $m_p = m_n = m_D/2$. We then have

$$\frac{n_D}{n_p n_n} = 6\left(\frac{m_n kT}{\pi\hbar^2}\right)^{-3/2}\exp\left(\frac{B_D}{kT}\right). \tag{24}$$

- This tells us that deuterium is favored in the limit $kT \to 0$, and that free protons and neutrons are favored in the limit $kT \to \infty$.
- Define T_{nuc} as the temperature at which $n_D/n_n = 1$; that is, the temperature at which half the free neutrons have been fused into deuterium nuclei.
- As long as equation 24 holds true, the deuterium-to-neutron ratio can be written as

$$\frac{n_D}{n_n} = 6n_p\left(\frac{m_n kT}{\pi\hbar^2}\right)^{-3/2}\exp\left(\frac{B_D}{kT}\right). \tag{25}$$

- We can write the deuterium-to-neutron ratio as a function of T and the baryon-to-photon ratio η if we make some simplifying assumptions. Even today, we know that $\sim$ 75% of all the baryons in the universe are in the form of unbound protons. Before the start of deuterium synthesis, 5 out of 6 baryons (or $\sim$ 83%) were in the form of unbound protons. So we can write

$$n_p \approx 0.8 n_b = 0.8\eta n_\gamma = 0.8\eta \left[0.243 \left(\frac{kT}{\hbar c} \right)^3 \right] \tag{26}$$

  where the factor in square brackets is the photon number density of blackbody radiation.

- Substituting this into Equation 25, we find that the deuterium-to-proton ratio is a relatively simple function of temperature:

$$\frac{n_D}{n_n} \approx 6.5\eta \left(\frac{kT}{m_n c^2} \right)^{3/2} \exp\left(\frac{B_D}{kT} \right). \tag{27}$$

- The temperature $T_{\rm nuc}$ of deuterium nucleosynthesis can be found by solving this equation with $n_D/n_n = 1$. With $m_n c^2 = 939.6$ MeV, $B_D = 2.22$ MeV, and $\eta = 5.5 \times 10^{-10}$ (more on that later), the temperature of deuterium synthesis is $kT_{\rm nuc} \approx 0.066$ MeV, corresponding to $T_{\rm nuc} \approx 7.6 \times 10^8$ K.

- The temperature drops to this value when the age of the universe is $t_{\rm nuc} \approx 200$ s.

- Note that the time delay until the start of nucleosynthesis, $t_{\rm nuc} \approx 200$ s, is not negligible compared to the decay time of the neutron, $\tau_n = 890$ s. By the time nucleosynthesis actually gets underway, neutron decay has slightly decreased the neutron-to-proton ratio from $n_n/n_p = 1/5$ to $n_n/n_p \approx 0.15$, which in turn lowers the maximum possible ^{4}He mass fraction from $Y_{\rm max} \approx 0.33$ to $Y_{\rm max} \approx 0.27$.

3 Nucleosynthesis beyond deuterium

- The deuterium-to-neutron ratio n_D/n_n does not remain indefinitely at the equilibrium value given by Equation 25.

- Once a significant amount of deuterium forms, there are many possible nuclear reactions available. For instance, a deuterium nucleus can fuse with a proton to form ^{3}He:

$$\mathrm{D} + p \leftrightharpoons {}^3\mathrm{He} + \gamma \tag{28}$$

- Alternatively, it can fuse with a neutron to form ^{3}H (tritium):

$$\mathrm{D} + n \leftrightharpoons {}^3\mathrm{H} + \gamma \tag{29}$$

- Tritium is unstable; it spontaneously decays to ^{3}He, emitting an electron and an electron anti-neutrino in the process. However, the decay time of tritium is approximately 18 years; during the brief time that Big Bang Nucleosynthesis lasts, tritium can be regarded as effectively stable.

- Deuterium nuclei can also fuse with each other to form ^{4}He:

$$\mathrm{D} + \mathrm{D} \leftrightharpoons {}^4\mathrm{He} + \gamma \tag{30}$$

- However, it is more likely that the interaction of two deuterium nuclei will end in the formation of a tritium nucleus (with the emission of a proton),

$$\mathrm{D} + \mathrm{D} \leftrightharpoons {}^{3}\mathrm{H} + p \tag{31}$$

or the formation of a 3He nucleus (with the emission of a neutron),

$$\mathrm{D} + \mathrm{D} \leftrightharpoons {}^{3}\mathrm{He} + n. \tag{32}$$

- There is never a large amount of ^{3}H or ^{3}He present during the time of nucleosynthesis. Soon after they are formed, they are converted to ^{4}He by reactions such as

$$\begin{aligned} {}^{3}\mathrm{H} + p &\leftrightharpoons {}^{4}\mathrm{He} + \gamma & (33)\\ {}^{3}\mathrm{He} + n &\leftrightharpoons {}^{4}\mathrm{He} + \gamma & (34)\\ {}^{3}\mathrm{H} + \mathrm{D} &\leftrightharpoons {}^{4}\mathrm{He} + \mathrm{n} & (35)\\ {}^{3}\mathrm{He} + \mathrm{D} &\leftrightharpoons {}^{4}\mathrm{He} + \mathrm{p}. & (36) \end{aligned}$$

- None of these post-deuterium reactions involves neutrinos; they all involve the strong nuclear force, and have large cross-sections and fast reaction rates. This means that once nucleosynthesis begins, D, ^{3}H, and ^{3}He are all efficiently converted to ^{4}He.

- Once ^{4}He is reached, however, we run into a problem creating heavier elements.
  - For such a light nucleus, ^{4}He is exceptionally tightly bound, as illustrated in Figure 1.
  - By contrast, there are no stable nuclei with $A = 5$. If you try to fuse a proton or neutron to ^{4}He, it won't work; ^{5}He and ^{5}Li are not stable nuclei. Thus, ^{4}He is resistant to fusion with protons and neutrons.
  - Small amounts of ^{6}Li and ^{7}Li, the two stable isotopes of lithium, are made by reactions such as

    $$^{4}\mathrm{He} + \mathrm{D} \leftrightharpoons {}^{6}\mathrm{Li} + \gamma \tag{37}$$

    and

    $$^{4}\mathrm{He} +^{3}\mathrm{H} \leftrightharpoons {}^{7}\mathrm{Li} + \gamma \tag{38}$$

  - In addition, small amounts of ^{7}Be are made by reactions such as

    $$^{4}\mathrm{He} +^{3}\mathrm{He} \leftrightharpoons {}^{7}\mathrm{Be} + \gamma. \tag{39}$$

  - The synthesis of nuclei with $A > 7$ is hindered by the absence of stable nuclei with $A = 8$. For instance, if ^{8}Be is made by the reaction

    $$^{4}\mathrm{He} +^{4}\mathrm{He} \rightarrow^{8}\mathrm{Be} \tag{40}$$

    the ^{8}Be nucleus falls back apart into a pair of ^{4}He nuclei with a decay time of only $\tau = 3 \times 10^{-16}$ s.

- Bottom line: once deuterium begins to be formed, fusion up to the tightly bound ^{4}He nucleus proceeds very rapidly.

- Fusion of heavier nuclei occurs much less rapidly.

- The precise yields of the different isotopes involved in BBN are customarily calculated using a fairly complex computer code. The complexity is necessary because of the large number of possible reactions which can occur once deuterium has been formed, all of which have temperature-dependent cross-sections. Thus, there's a good deal of bookkeeping involved.

- The results of a typical BBN code, which follows the mass fraction of different isotopes as the universe expands and cools, is shown in Figure 2.

  - Initially, at $T \gg 10^9$ K, almost all the baryonic matter is in the form of free protons and neutrons.
  - As the deuterium density climbs upward, however, the point is eventually reached where significant amounts of ^{3}H, ^{3}He, and ^{4}He are formed.
  - By the time the temperature has dropped to $T \sim 4 \times 10^8$ K, at $t \sim 10$ min, Big Bang Nucleosynthesis is essentially over. Nearly all the baryons are in the form of free protons or ^{4}He nuclei. The small residue of free neutrons decays into protons. Small amounts of D, ^{3}H, and ^{3}He are left over, a tribute to the incomplete nature of Big Bang Nucleosynthesis. (The ^{3}H later decays to ^{3}He.) Very small amounts of ^{6}Li, ^{7}Li, and ^{7}Be are made, though the ^{7}Be is later converted to ^{7}Li by electron capture: $^7\mathrm{Be} + e^- \rightarrow\ ^7\mathrm{Li} + \nu_e$.)

- Yields of D, ^{3}He, ^{4}He, ^{6}Li, and ^{7}Li depend on various physical parameters.

- Most importantly, they depend on the baryon-to-photon ratio η.

- A high baryon-to-photon ratio increases the temperature $T_{\rm nuc}$ at which deuterium synthesis occurs, and hence gives an earlier start to Big Bang Nucleosynthesis. Since BBN is a race against the clock as the density and temperature of the universe drop, getting an earlier start means that nucleosynthesis is more efficient at producing ^{4}He, leaving less D and ^{3}He as leftovers.

- A plot of the mass fraction of various elements produced by Big Bang Nucleosynthesis is shown in Figure 3. Note that larger values of η produce larger values for the ^{4}He mass fraction Y_p and smaller values for the deuterium density, as explained above.

- The dependence of the ^{7}Li density on η is more complicated. Within the range of η plotted in Figure 3, the direct production of ^{7}Li by the fusion of ^{4}He and ^{3}H is a decreasing function of η, while the indirect production of ^{7}Li by ^{7}Be electron capture is an increasing function of η. The net result is a minimum in the predicted density of ^{7}Li at $\eta \approx 3 \times 10^{-10}$.

- Broadly speaking, we know immediately that the baryon-to-photon ratio can't be as small as $\eta \sim 10^{-12}$. If it were, BBN would be extremely inefficient, and we would expect only tiny amounts of helium to be produced ($Y_p < 0.01$).

- Conversely, we know that the baryon-to-photon ratio can't be as large as $\eta \sim 10^{-7}$. If it were, nucleosynthesis would have taken place very early (before neutrons had a chance to decay), the universe would be essentially deuterium-free, and Y_p would be near its maximum permissible value of $Y_{\rm max} \approx 0.33$.

- Pinning down the value of η more accurately requires making accurate observations of the primordial densities of the light elements; that is, the densities before nucleosynthesis in stars started to alter the chemical composition of the universe. In determining the value of η, it is most useful to determine the primordial abundance of deuterium. This is because the deuterium abundance is strongly dependent on η in the range of interest. Thus, determining the deuterium abundance with only modest accuracy will enable us to determine η fairly well.

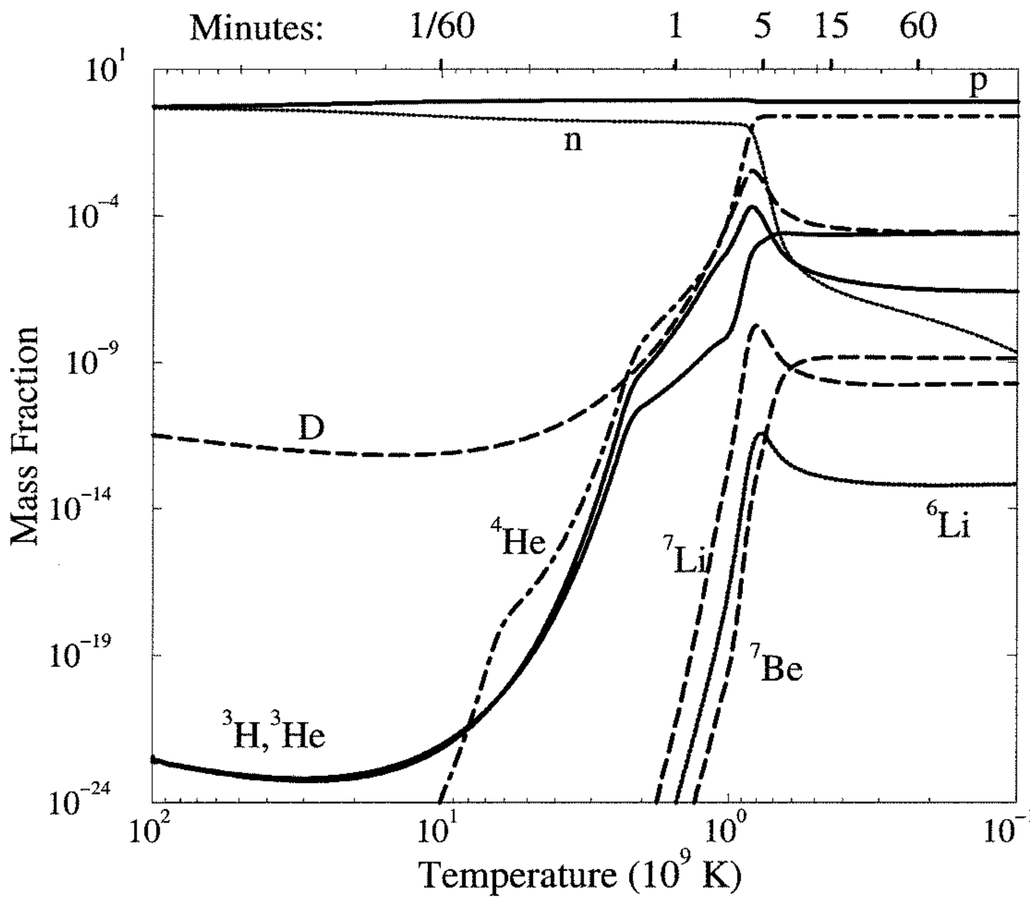

Figure 2: Mass fraction of nuclei as a function of time during the epoch of nucleosynthesis. A baryon-to-photon ratio of $\eta = 5.1 \times 10^{-10}$ is assumed (from Tytler et al., 2000. Physica Scripta,T85, 12).

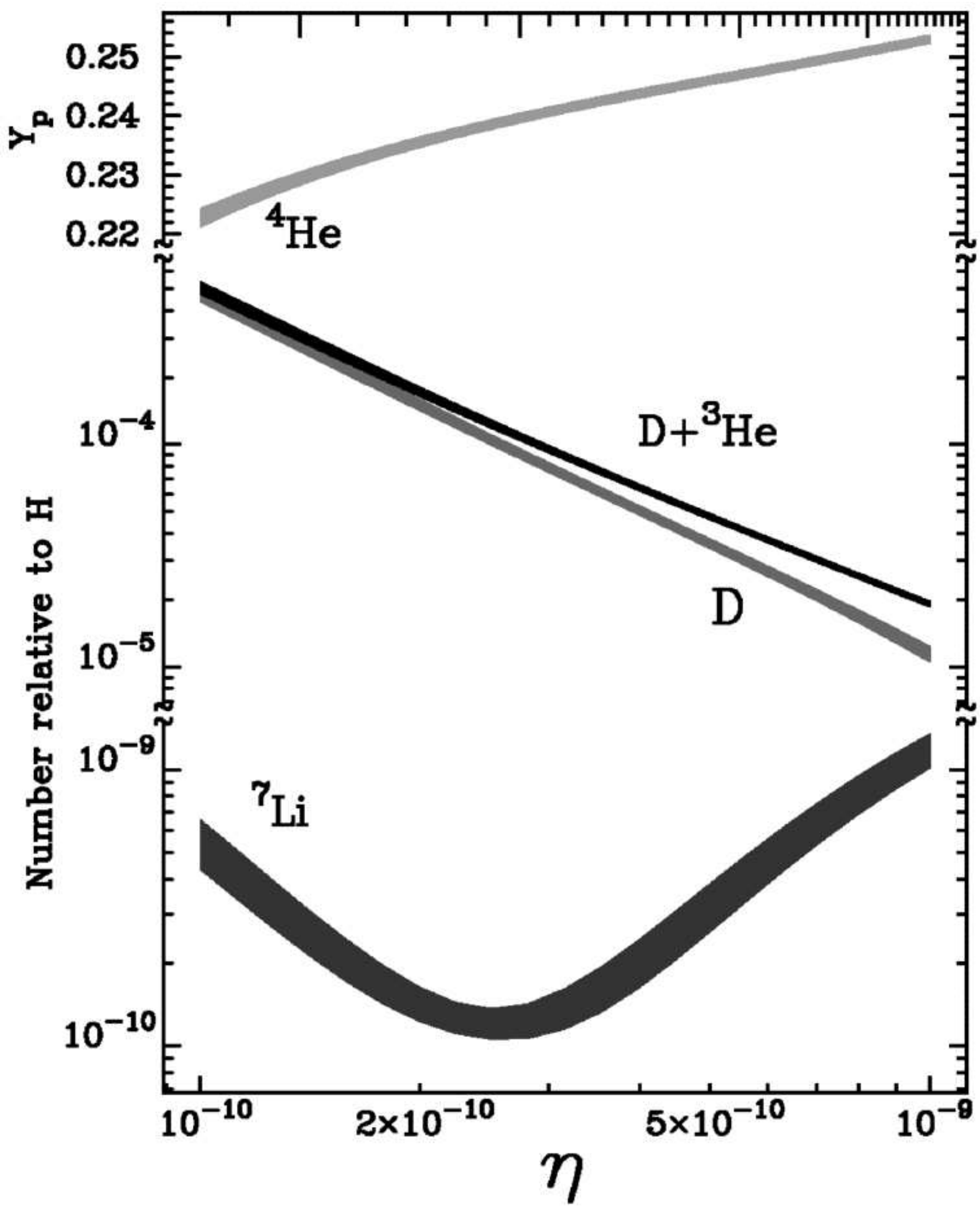

Figure 3: The mass fraction of ^{4}He, and the number densities of D, D+^{3}He, and ^{7}Li expressed as a fraction of the H number density. The width of each line represents the 95% confidence interval in the density (from Burles, Nollett, and Turner, 2001, *ApJ*, 552, L1).

- By contrast, the primordial helium fraction Y_p has only a weak dependence on η for the range of interest, as shown in Figure 3. Thus, determining η with a fair degree of accuracy would require an extremely precise measurement of Y_p.

- A major problem in determining the primordial deuterium abundance is that deuterium is very easily destroyed in stars. When an interstellar gas cloud collapses gravitationally to form a star, the first fusion reactions that occur involve the fusion of deuterium into helium. Thus, the abundance of deuterium in the universe tends to decrease with time.

- Deuterium abundances are customarily given as the ratio of the number of deuterium atoms to the number of hydrogen atoms (D/H). For instance, in the local interstellar medium within our Galaxy, astronomers find an average value D/H $\approx 1.6 \times 10^{-5}$; that is, there's one deuterium atom for every 60,000 ordinary hydrogen atoms. However, the Sun and the interstellar medium are contaminated with material that has been cycled through stellar interiors. Thus, we expect the primordial deuterium-to-hydrogen value to have been (D/H)$_p > 1.6 \times 10^{-5}$.

- Currently, the most promising way to find the primordial value of D/H is to look at the spectra of distant quasars. We don't care about the deuterium within the quasar itself; instead, we just want to use the quasar as a flashlight to illuminate the intergalactic gas clouds which lie between it and us. If an intergalactic gas cloud contains no detectable stars, and has very low levels of elements heavier than lithium, we can hope that its D/H value is close to the primordial value, and hasn't been driven downward by the effects of fusion within stars.

- Neutral hydrogen atoms within these intergalactic clouds will absorb photons whose energy corresponds to the Lyman-α transition (the transition of the atom's electron from the ground state $n = 1$ to the next higher energy level $n = 2$). In an ordinary hydrogen atom (1H), the Lyman-α transition corresponds to a wavelength λ_H = 121.57 nm. In a deuterium atom, the greater mass of the nucleus causes a small isotopic shift in the electron's energy levels. As a consequence, the Lyman-α transition in deuterium corresponds to a slightly shorter wavelength, λ_D = 121.54 nm.

- When we look at light from a quasar which has passed through an intergalactic cloud at redshift $z_{\rm cl}$, we will see a strong absorption line at $\lambda_H(1 + z_{\rm cl})$, due to absorption from ordinary hydrogen, and a much weaker absorption line at $\lambda_D(1 + z_{\rm cl})$, due to absorption from deuterium.

- An example of this is shown in Figure 4.

- In principle the measurement is straightforward; we can estimate the quantity of hydrogen from the strength of the hydrogen line and the quantity of deuterium from the strength of the deuterium line. In practice it's more complicated, however.

- The small wavelength spacing and weakness of the deuterium line mean that high resolution, high signal-to-noise spectra are required, meaning that the number of quasars for which the measurement can be performed is limited.

- In addition the measurement can only be made for clouds in a narrow range of optical depth (the fraction of photons absorbed by the cloud). Too much absorption, and the hydrogen line will be too wide and cover the deuterium line; too little, and the deuterium line will be too weak.

- We must also hope that the deuterium line isn't due to a second, smaller cloud at a slightly different redshift, which would cause us to overestimate the amount of deuterium, and that stellar processes in the cloud haven't destroyed some of the deuterium, which would mean that our measurement is an underestimate and the deuterium is not primordial.

- The six best examples give (D/H) $= (2.68 \pm 0.08) \times 10^{-5}$.
- If we assume that this reflects the primordial abundance of deuterium, we then get a precise measurement of the photon-to-baryon ratio

$$\eta = (6.0 \pm 0.4) \times 10^{-10}. \tag{41}$$

- This has been confirmed by WMAP measurements of the CMB, as we'll discuss later. We can also use η to determine the current baryon density $n_{b,0} = \eta n_{\gamma,0}$; the result is again consistent with measurements from WMAP.

4 Baryon-antibaryon asymmetry

- The results of Big Bang Nucleosynthesis tell us what the universe was like when it was relatively hot ($T_{\rm nuc} \approx 7 \times 10^8$ K) and dense ($u_{\rm nuc} \approx 10^{33}$ MeV m^{-3}), with the energy density almost entirely in the form of radiation.
- The mean photon energy and mean mass density weren't extremely high; the physics of BBN is well-understood, as demonstrated by the excellent agreement of predictions from BBN with WMAP measurements.
- Some of the initial conditions for Big Bang Nucleosynthesis, however, are rather puzzling.
  - The baryon-to-photon ratio, $\eta \approx 6 \times 10^{-10}$, is a remarkably small number; the universe seems to have a strong preference for photons over baryons.
  - It's also worthy of remark that the universe seems to have a strong preference for baryons over antibaryons. The laws of physics demand the presence of antiprotons ($\bar{p}$), containing two anti-up quarks and one anti-down quark apiece, as well as antineutrons ($\bar{n}$), containing one anti- up quark and two anti-down quarks apiece. In practice, though, it is found that the universe has an extremely large excess of protons and neutrons over antiprotons and antineutrons (and hence an excess of quarks over antiquarks).
  - At the time of Big Bang Nucleosynthesis, the number density of antibaryons ($\bar{n}$ and $\bar{p}$) was tiny compared to the number density of baryons, which in turn was tiny compared to the number density of photons. This imbalance, $n_{\rm antib} \ll n_b \ll n_\gamma$, has its origin in the physics of the very early universe.
- When the temperature of the early universe was greater than $kT \approx 150$ MeV, the quarks which it contained were not confined within baryons and other particles, as they are today, but formed a sea of free quarks (sometimes called quark soup).
- During the first few microseconds of the universe, when the quark soup was piping hot, quarks and antiquarks were constantly being created by pair production and destroyed by mutual annihilation:

$$\gamma + \gamma \leftrightharpoons q + \bar{q} \tag{42}$$

  where q and $\bar{q}$ represent any quark-antiquark pair.

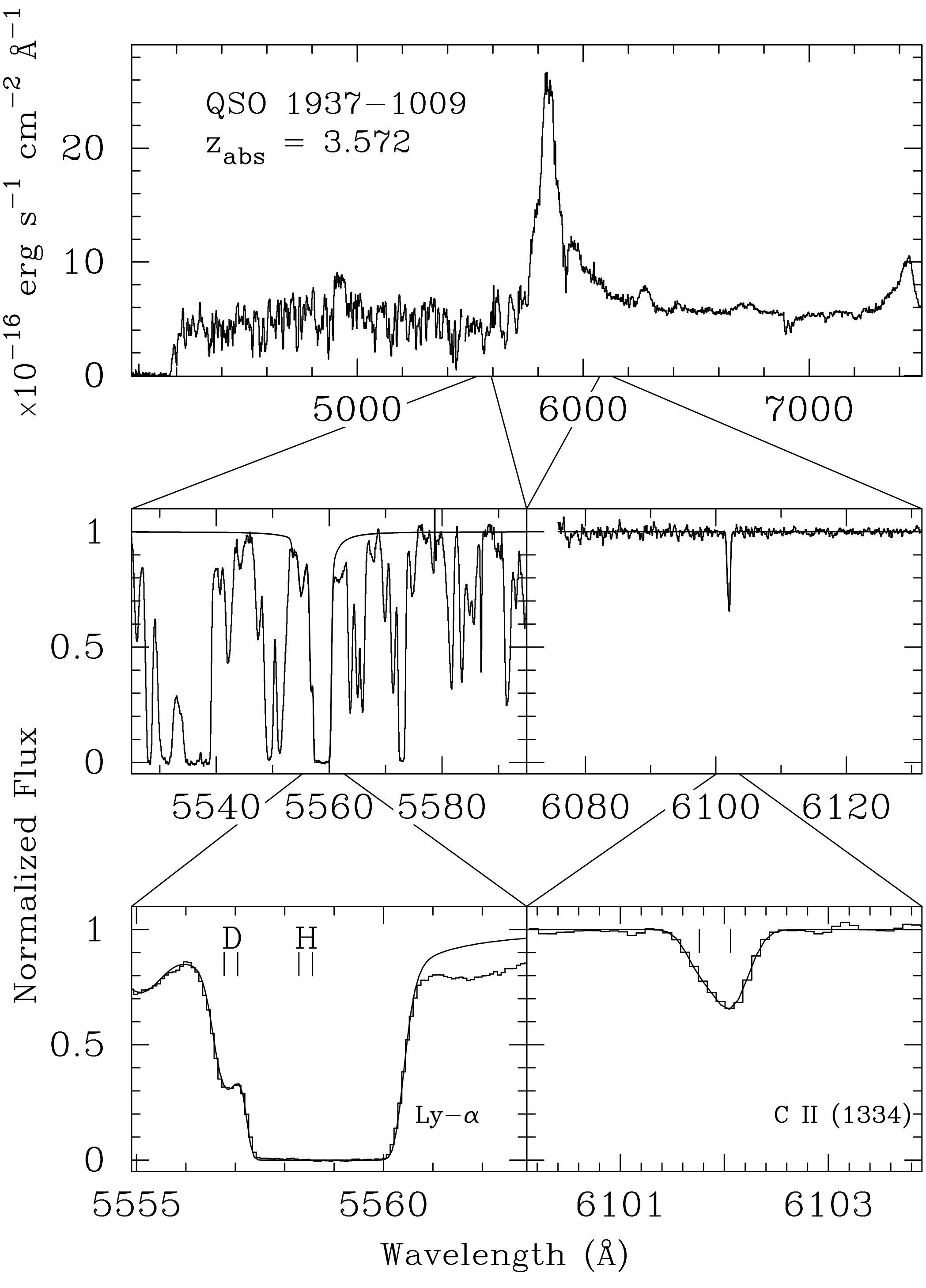

Figure 4: Absorption from hydrogen and deuterium along the line of sight to a quasar. From Burles & Tytler, Stellar Evolution, Stellar Explosions and Galactic Chemical Evolution, Proceedings of the 2nd Oak Ridge Symposium on Atomic and Nuclear Astrophysics, Oak Ridge, Tennessee, 2-6 December 1997 (arXiv:astro-ph/9803071). See also Burles & Tytler, 1998, *ApJ*, 499, 699.

- During this period of quark pair production, the numbers of up quarks, anti-up quarks, down quarks, anti-down quarks, and photons were nearly equal to each other. However, suppose there were a very tiny asymmetry between quarks and antiquarks, such that

$$\delta_q \equiv \frac{n_q - n_{\bar{q}}}{n_q + n_{\bar{q}}} \ll 1. \tag{43}$$

- As the universe expanded and the quark soup cooled, quark-antiquark pairs would no longer be produced. The existing antiquarks would then annihilate with the quarks. However, because of the small excess of quarks over antiquarks, there would be a residue of quarks with number density

$$\frac{n_q}{n_\gamma} \sim \delta_q. \tag{44}$$

- Thus, if there were 1,000,000,000 quarks for every 999,999,997 antiquarks in the early universe, three lucky quarks in a billion would be left over after the others encountered antiquarks and were annihilated. The leftover quarks, however, would be surrounded by roughly 2 billion photons, the product of the annihilations.
- After the three quarks were bound together into a baryon at $kT \approx 150$ MeV, the resulting baryon-to-photon ratio would be $\eta \sim 5 \times 10^{-10}$.
- Thus, the very strong asymmetry between baryons and antibaryons today and the large number of photons per baryon are both products of a tiny asymmetry between quarks and antiquarks in the early universe.
- The exact origin of the quark-antiquark asymmetry in the early universe is still not exactly known. The physicist Andrei Sakharov, as far back as the year 1967, was the first to outline the necessary physical conditions for producing a small asymmetry; however, the precise mechanism by which the quarks first developed their few-parts-per-billion advantage over antiquarks still remains to be found.