

## VI. The First Three Minutes (Big Bang Nucleosynthesis)

Reading: Sections 9.0, 9.2, 9.3. Chapter 10. We will come back to Chapter 9, but these sections include background material for Chapter 10.

Using General Relativity, the assumption of a homogeneous and isotropic universe, and the standard model of nuclear and particle physics, we can predict what should happen during the first three minutes of cosmic history.

While we do not observe this epoch directly, we do observe its residue, the cosmic abundances of  $^4\text{He}$ ,  $\text{D}$ ,  $^3\text{He}$ , and  $^7\text{Li}$ .

These allow us to test important aspects of this story.

### The baryon-to-photon ratio

At temperature  $T$ , the number density of photons in a blackbody distribution is

$$n_\gamma \approx 0.244 \left( \frac{kT}{\hbar c} \right)^3.$$

With the present day temperature  $T_{\text{CMB}} = 2.73 \text{ K}$ , one can show that the ratio of baryons to photons is

$$\eta \equiv n_b/n_\gamma = 5.4 \times 10^{-10} \left( \frac{\Omega_b h^2}{0.02} \right),$$

where  $\Omega_b$  is the ratio of the mean density of baryons to the critical density,  $h \equiv H_0/100 \text{ km s}^{-1} \text{ Mpc}^{-1}$ , and I have scaled  $\Omega_b h^2$  to the current observational estimate.

The photon density itself is  $413 \text{ photons cm}^{-3}$ .

Roughly speaking, there is one baryon (proton or neutron) for every billion CMB photons.

### Age vs. Temperature

At redshift  $z$ , the temperature of the photon background is

$$T = 2.73 \times (1 + z) \text{ K}, \quad kT = 2.39 \times 10^{-4} \times (1 + z) \text{ eV}.$$

Before  $z = z_{\text{eq}} \sim 10^4$ , the dominant form of energy in the universe was radiation: photons and neutrinos.

Curvature was negligible, so the Friedmann equation can be used to relate the Hubble parameter to the radiation temperature

$$H^2 = \frac{8\pi G}{3} \frac{\epsilon(T)}{c^2}, \quad \epsilon(T) = a_{\text{SB}} T^4 \times 1.68.$$

The extra 0.68 comes from the neutrino contribution. The  $T$  in this equation is actually the photon temperature; the neutrino temperature is somewhat lower.

In a radiation-dominated universe,  $t = 1/(2H) \propto 1/T^2$ . With the appropriate constants

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

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

### Particles and anti-particles in the early universe

If the rates of reactions that exchange energy between particles are fast compared to the expansion rate, the populations should relax to thermal equilibrium.

Consider a particle of mass  $m_K$ . If  $kT > 2m_K$  and the rate of reactions that can create  $K\bar{K}$  pairs from other particles (e.g., photons) are fast compared to the expansion rate, maximizing entropy  $\implies$  an abundance of  $K$  and  $\bar{K}$  particles roughly equal to the photon abundance (up to statistical weight factors).

More quantitatively, the number density of a coupled species relative to the number density of photons is suppressed by a factor  $\exp(-m_K c^2/kT)$ .

If the  $K\bar{K}$  annihilation rate drops below the expansion rate while  $kT > 2m_K$ , the particle decouples and redshifts thereafter independent of other species. This happens to neutrinos, and it may happen to other weakly interacting particles.

If the temperature falls below  $2m_K$  while the  $K\bar{K}$  annihilation rate is high, the particles will annihilate and dump their energy into the background of still-coupled particles. This happens to electrons and positrons.

In one of the leading scenarios for dark matter, WIMPS (Weakly Interacting Massive Particles) are first suppressed in number as the temperature falls below their rest mass, then decouple once the typical annihilation time (determined by the weak interaction rate) exceeds the age of the universe. The dark matter would then consist of equal numbers of WIMPS and anti-WIMPS.

### Baryon asymmetry

The local universe has baryons but almost no anti-baryons.

This asymmetry is small in the sense that the baryon-to-photon ratio is  $n_b/n_\gamma \sim 10^{-9}$ . Since the particle density in the relativistic era was  $\sim n_\gamma$ , *almost* all relativistic baryons were annihilated by anti-baryons.

The baryon asymmetry (and corresponding  $e^+/e^-$  asymmetry) may be an initial condition of the universe.

Alternatively, the baryon asymmetry may have arisen after the Big Bang.

A number of specific models of “baryogenesis” have been proposed, none entirely compelling. Some place baryogenesis at the strong-electroweak (“GUT”) symmetry breaking, others at electroweak symmetry breaking. Since the level of CP violation observed in the weak interaction is  $\sim 10^{-9}$ , it is not unreasonable to imagine creating a baryon asymmetry of the observed order.

### Some early transitions

At  $t \sim 10^{-11}$  s,  $kT \sim M_{W,Z}$ , electroweak unification breaks.

Before this time, the  $W$  and  $Z$  bosons are effectively “massless” (because their rest masses are small compared to the characteristic particle energy), and the electromagnetic and weak forces are of essentially equal strength.

After this time, weak interactions are much weaker (lower cross section) than electromagnetic interactions, and they are short range (nuclear scale only).

Details of this transition are poorly understood.

It is possible that baryogenesis occurs at this phase transition.

At  $t \sim 10^{-6}$  s,  $kT \sim$  nucleon binding energy, quarks combine into hadrons.

Before this time, quarks behave as free particles.

After this time they appear only bound into hadrons, principally protons and neutrons (the only hadrons with long lifetimes).

Details of this transition are poorly understood.

This is also the epoch when  $kT \sim m_p c^2$ .

$p\bar{p}$  and  $n\bar{n}$  annihilation leaves behind a small residue of protons and neutrons,  $n_p \sim n_n \sim \eta n_\gamma$ .

We still have  $kT \gg m_e c^2$ , so  $n_{e^-} \sim n_{e^+} \sim n_\gamma$ .

At  $t \sim 1$  s,  $\sigma_\nu n_\nu c \sim 1$ , neutrinos decouple.

(Here  $\sigma_\nu$  is the typical neutrino interaction cross section,  $n_\nu$  is the number density of neutrinos, and  $c$  is the speed of light.)

Before this time, weak interactions are fast enough to keep the number and energy of neutrinos in equilibrium with other species.

After this time, a typical neutrino has no interaction with any other particles, and the neutrinos evolve in isolation, redshifting in energy as the universe expands.

### The neutron-to-proton ratio

After electron-positron annihilation, the constituents of the universe are photons, neutrinos, and (at an abundance smaller by a factor  $\sim 10^9$ ) protons, neutrons, and electrons.

Weak interactions involving neutrinos can still convert neutrons to protons and vice versa.

Since neutrons are more massive than protons, they are less abundant — conversion of a neutron to a proton is less probable than conversion of a proton to a neutron.

In thermal equilibrium

$$\frac{n_n}{n_p} = e^{-Q/kT}, \quad Q \equiv (m_n - m_p)c^2 = 1.2934 \text{ MeV}.$$

The conversion reactions become slow compared to the age of the universe at  $t \sim 3$  seconds,  $kT \sim 0.7$  MeV.

Since the interaction rate is dropping quickly as the density and temperature of the universe decline, there are no subsequent conversions.

The neutron-to-proton ratio “freezes in” at

$$\frac{n_n}{n_p} \approx e^{-1.2934/0.7} \approx 1/6.$$

## Deuterium synthesis

Synthesis of elements heavier than hydrogen has to start with deuterium formation ( $n+p \rightarrow \text{D} + \gamma$ ).

The binding energy of deuterium is  $B_D = 2.22 \text{ MeV}$ .

Naively one expects deuterium synthesis to begin when the temperature falls to  $kT \sim B_D$ .

However, the baryon-to-photon ratio is  $\eta \sim 5 \times 10^{-10}$ , so the exponential tail of the blackbody distribution can still dissociate deuterium even when  $kT$  is significantly below  $B_D$ .

Synthesis of deuterium actually begins when  $kT \sim B_D / -\ln \eta \sim 0.1$  MeV, at time  $t \sim 2$  minutes.

The decay time for free neutrons is  $\sim 900$  sec, so in two minutes a small but non-negligible fraction of the neutrons left over from “freeze-out” have decayed.

The neutron-to-proton ratio at the time of deuterium synthesis is  $n/p \sim 1/7$ .

## The ${}^4\text{He}$ fraction

The reaction rate for  $n + p \rightarrow \text{D} + \gamma$  is fast, so when the temperature falls below 0.1 MeV, all neutrons are quickly processed into deuterium.

The reactions that process D into  ${}^4\text{He}$  are also fast, and to first order all neutrons go into  ${}^4\text{He}$ .

A robust prediction of the standard big bang model is that  $24 \pm 1\%$  of the baryonic mass of the universe is in  ${}^4\text{He}$ , with almost all of the rest in hydrogen.

There is a weak dependence of the predicted helium fraction on  $\eta$ , and uncertainty in  $\eta$  is the main contribution to the  $\pm 1\%$  uncertainty.

Just as the mass fraction of heavy elements is usually written  $Z$ , the mass fraction of helium is usually written  $Y$  (and hydrogen  $X$ ).

The *primordial* helium abundance, which is the mass fraction of helium produced in the big bang, not including later contributions from stars, is written  $Y_P$ .

*The Observations:* The helium fraction in the interstellar gas of the most metal-poor galaxies is  $Y = 0.24 \pm 0.01$ . It is difficult to reduce the systematic uncertainties in the observations below 0.01.

## Deuterium

A small fraction ( $\sim 10^{-4} - 10^{-5}$ ) of D “escapes” and is never processed into heavier nuclei.

This fraction is sensitive to  $\eta$ .

The primordial deuterium abundance is therefore a good way to determine the mean baryon density of the universe.

Stars can destroy deuterium by fusing it into heavier elements, but as far as we know they cannot make it because there are no conditions under which they will make deuterium and not fuse it into heavier elements.

*The Observations:* The measured deuterium-to-hydrogen ratio in the interstellar medium of distant, metal-poor gas absorbing the light of background quasars is  $(D/H) = (2.8 \pm 0.5) \times 10^{-5}$ . Taking this to be the primordial ratio implies  $\Omega_b h^2 = 0.021 \pm 0.002$ .

## $^3\text{He}$ and $^7\text{Li}$

Incomplete processing leaves a  $^3\text{He}$  fraction similar to that of  $D$  ( $\sim 10^{-5}$ ).

A small amount ( $\sim 10^{-10}$ ) of  $^7\text{Li}$  is produced by  $^4\text{He} + ^3\text{He} \rightarrow ^7\text{Li}$ .

$^3\text{He}$  and  $^7\text{Li}$  can both be destroyed and produced in stars, which makes it challenging to determine the primordial value.

*The Observations:* Accounting for the uncertainties in stellar production and destruction, which are a factor of several, measurements are consistent with the BBN (Big Bang Nucleosynthesis) predictions.

## The Roadblock

There are no stable elements of atomic number 5 or 8.

Starting with protons and neutrons, there is no way to bridge the gap at atomic number 8 to build heavier nuclei.

Stars do it by the “triple- $\alpha$ ” reaction:  $^4\text{He} + ^4\text{He} + ^4\text{He} \rightarrow ^{12}\text{C}^*$ , but this requires high temperature *and* density.

Consequence: only light elements are made in the early universe. All elements heavier than  ${}^7\text{Li}$  are made in stars.

### The BBN Bottom Line

The observationally inferred primordial abundances of  ${}^4\text{He}$ , D,  ${}^3\text{He}$ , and  ${}^7\text{Li}$  are consistent with the predictions of the standard big bang model, which assumes GR, a homogeneous and isotropic universe, and the standard model of nuclear and particle physics.

The  ${}^4\text{He}$  test has a precision of  $\sim 10\%$ , while the others have a precision of a factor of several.

This agreement between predictions and observations is strong evidence that the standard big bang model applies back to at least  $t \sim 1$  s and  $T \sim 10^{10}\text{K}$ .

This agreement rules out many possible variations on the standard model of particle physics, such as extra light neutrino species, or time-variation of the fundamental constants that control the strength of the nuclear, electromagnetic, and weak forces.

The primordial D /H ratio constrains the baryon density to  $\Omega_b h^2 = 0.021 \pm 0.002$ , assuming the standard big bang model to be correct.

Recently, CMB anisotropies have provided an entirely independent way to constrain the baryon density, yielding  $\Omega_b h^2 = 0.022 \pm 0.001$ .

The agreement of these two independent determinations is further strong evidence for the standard model.