SETI AT PLANCK ENERGY: WHEN PARTICLE PHYSICISTS BECOME COSMIC ENGINEERS

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ABSTRACT

What is the meaning of the Fermi Paradox – are we alone or is starfaring rare? Can general relativity be united with quantum mechanics? The searches for answers to these questions could intersect. It is known that an accelerator capable of energizing particles to the Planck scale requires cosmic proportions. The energy required to run a Planck accelerator is also cosmic, of order $100 M_{\odot} c^2$ for a hadron collider, because the natural cross section for Planck physics is so tiny. If aliens are interested in fundamental physics, they could resort to cosmic engineering for their experiments. These colliders are detectable through the vast amount of "pollution" they produce, motivating a YeV SETI program. I investigate what kinds of radiation they would emit in a fireball scenario, and the feasibility of detecting YeV radiation at Earth, particularly YeV neutrinos. Although current limits on YeV neutrinos are weak, Kardashev 3 YeV neutrino sources appear to be at least 30–100 Mpc apart on average, if they are long-lived and emit isotropically. I consider the feasibility of much larger YeV neutrino detectors, including an acoustic detection experiment that spans all of Earth’s oceans, and instrumenting the entire Kuiper Belt. Any detection of YeV neutrinos implies an extraordinary phenomenon at work, whether artificial and natural. Searches for YeV neutrinos from any source are naturally commensal, so a YeV neutrino SETI program has value beyond SETI itself, particularly in limiting topological defects. I note that the Universe is very faint in all kinds of nonthermal radiation, indicating that cosmic engineering is extremely rare.

Subject headings: extraterrestrial intelligence — astroparticle physics — neutrinos

1. INTRODUCTION

We are here. They are not. These appear to be the basic observations that we have so far in the Search for Extraterrestrial Intelligence (SETI). On the one hand, we evolved on Earth, demonstrating that beings with the technology to communicate with other stars arise with nonzero probability. The vastness of the Universe in space and time then suggests that ETs have had ample time to alter the Galaxy around them. If they do not, cosmic engineering is relatively easy to see, across intergalactic distances for the largest projects (Akamatsu & Elvis 2011). Others are motivated by consumption, in which aliens wish to maximize their population, available energy, or computational power. But whatever the underlying motivation, cosmic engineering is relatively easy to see, across intergalactic distances for the largest projects (Kardashev 1985). Nothing seems to forbid cosmic engineering, and ETs have had ample time to alter the Galaxy around us. They apparently should not only be here, but they practically wouldn’t even have to look for them.

Yet there are no credible and unambiguous signs of ETs anywhere so far, whether in the Solar System, elsewhere in the Galaxy, or in other galaxies. This is despite a number of surveys being carried out with a variety of methods. Radio searches are the primary tool of SETI, particularly focusing on narrowband signals (Cocconi & Morrison 1959; Tarter 2001). Although these turned up a few intriguing candidate artificial signals, none of them have ever been detected again (Gray 1994; Gray & Marvel 2001; Lazio et al. 2002; Gray & Ellingsen 2002), and overall no compelling candidate ETs have been found (Blair et al. 1992; Horowitz & Sagan 1993; Shostak et al. 1996; Rampadarath et al. 2012). Optical and near-infrared laser pulses are a plausible method of interstellar communication, even with our current technology (Schwartz & Townes 1961), and optical SETI programs aim...
to detect them (Eichler & Beskin 2001; Howard et al. 2004; Hanna et al. 2009). There have been a few attempts to search for artifacts in the Solar System as well, to check whether they are in fact here (Freitas & Valdes 1980; Freitas 1983; Steel 1993). Finally, cosmic engineering is thought to be detectable through its waste heat (Dyson 1960; Kardashev 1964; Sagan & Walker 1966), so far-to-mid-infrared surveys constrain its prevalence (Slysh 1983; Carrigan 2009; Wright et al. 2014b).

Other channels for SETI have been proposed, including: looking for artificial planetary transits (Arnold 2005), deliberate stellar variability (Learned et al. 2008), or industry (Whitmire & Wright 1980; Loeb & Turner 2012; Lin et al. 2014); more thorough searches for artifacts on planetary surfaces, interplanetary space, or starships (Bracewell 1960; Rose & Wright 2004; Harris 2002; Davies 2012); and searching for messages sent as high energy neutrinos (Subotowicz 1979; Learned 1994; Silagadze 2008; Learned et al. 2009) or cosmic rays (CRs; Swan 2006). Neutrino communication in particular has been demonstrated on Earth (Stancil et al. 2012). Nothing artificial has stood out yet with these methods either.

So, are we alone in the Galaxy, or even the observable Universe? Or does something always prevent ETs from starfaring and cosmic engineering – with the ominous implication that humanity has no chance of achieving starfaring either? The Fermi Paradox, with its pessimistic implications for SETI and its frightening insinuations about us, provokes a lot of controversy (Cirkovic 2009). The first issue is whether interstellar flight and settlement is feasible. But the ETs may not even have to physically send astronauts to other stars. As the notorious counterargument goes, they might launch replicating machines in their stead – yet we do not see signs of these machines either (Tipler 1980). The other main issue is whether ETs decide to settle other stars and begin cosmic engineering. As often phrased, the Fermi Paradox seems to assume that ETs are interested in consuming the Galaxy for the mere sake of consumption. Might ETs simply have different values that discourage this kind of expansion (c.f. Anderson 1960; Sagan & Newman 1983; Haqq-Misra & Baum 2009)?

The problem is that everyone must agree not to go starfaring. Even a single species of ETs may be composed of a multitude of societies, themselves composed of a multitude of groups, with varying goals over many historical epochs (Hart 1975; Wright et al. 2014a). After all, our own societies include a wide range of opinions on whether ETs exist, how we might find them, and particularly whether we should contact them. Arguments from large numbers are turned on their head: if it defies belief that we are alone among millions of stars, it defies belief that there are no cosmic engineers or starfarers among millions of societies. The Fermi Paradox highlights an instability; as with a hydrodynamic instability, only one unstable mode can bring everything crashing down. But is there a good reason to cosmic engineer, something to trigger the instability?

I propose curiosity as a motivation for cosmic engineering. The other great quest in physics and astronomy of the past few decades has been the search for a unified theory of physics including gravity. The most natural energy scale for unification is the Planck energy, $E_{\text{Planck}} = \sqrt{h\kappa G}$, which evaluates to $1.22 \times 10^{28}$ eV = 12.2 XeV – far, far beyond the reach of any accelerator possible with our technology. While various extensions to the Standard Model of particle physics have been proposed, there seems to be no evidence of any of them in the LHC (Chatrchyan et al. 2012; Aad et al. 2013) or in other precision tests (many searches are listed in Olive & Particle Data Group 2014). This leads to speculations that we are living in the “Nightmare Scenario”, in which there is no new physics until the Planck scale itself (Cho 2007). Of course, we know that there is physics beyond the Standard Model in the form of dark matter and dark energy. But we have no guarantee that they can be solved without Planck-scale physics. Or – maybe just as bad – they might turn out to be just another set of fields, providing no revolutionary insights on how to unify fundamental physics. We appear to be facing a physics “desert” between 1 TeV and 10 XeV, with no observations telling us how to get across it.

In discussions of the Nightmare Scenario and the physics desert, the idea of probing Planck scale physics directly is generally dismissed because it would require a particle accelerator that is “as big as the Solar System”, “the size of the Galaxy”, or “the size of the whole Universe” (Akahito 1989; Green 1999; Davies 2003; Susskind 2008; Kakui 2010; Adler 2010). But starfaring ETs conceivably could build an accelerator the size of a galaxy over a few million years. If ETs are interested in fundamental physics, they may resort to such engineering. So why not look for artificial particle accelerators that are literally the size of a galaxy?

Aside from resolving the accelerators themselves, or seeing their waste heat, we could search for YeV–XeV radiation generated within the accelerator. Neutrinos could be particularly effective in escaping the accelerator and reaching the Earth (Jackson 1998, p. 777). The natural background of ultra-high energy cosmic rays (UHECRs) falls off above $\sim 40$ EeV, and none have been observed with an energy greater than $\sim 1$ ZeV (Abbasi et al. 2008; Abraham et al. 2008; 2010). Although YeV–XeV cosmic rays (CRs) might be emitted by Planck-scale relics of the early Universe like cosmic strings (e.g., Hill et al. 1987; Birke & Sarkad 1998; Berezinsky et al. 2009), these “top down” models are strongly constrained (Rubtsov et al. 2006). Detection of YeV–XeV radiation would be a revolutionary discovery, implying either a completely new class of accelerators or new physics (Thompson & Lacki 2011).

Because the paper is about individual relativistic particles, I use Gaussian centimeter-gram-second units (Jackson 1998, p. 777). The symbols for the constants of nature appear in Table 1. I also rate ET artifacts’ power output $L$ with the Kardashev scale as quantified by Kardashev (1964):

$$K = 2.0 + 0.1 \log_{10} \left( \frac{L}{L_{\odot}} \right) .$$

So a power of $1 L_{\odot}$ is given as Kardashev 2.0 (K2.0 or K2), a power of $10^3 L_{\odot}$ is Kardashev 2.5 (K2.5), a power of $10^{10} L_{\odot}$ is Kardashev 3.0 (K3.0 or K3), and so on.

1 The SI prefixes for very large numbers are E ($10^{18}$; exa), Z ($10^{21}$; zetta), and Y ($10^{24}$; yotta). No larger SI prefixes have been decreed, but I adopt symbols that continue in reverse alphabetical order: X ($10^{15}$), W ($10^{18}$), and V ($10^{20}$), as proposed (sometimes in jest) by Jeff Aronson, Jim Blower, and Shbis Saiban. The history of these systems is described by Saiban (2015).

2 When I refer to “neutrinos” in this paper, I actually mean both neutrinos ($\nu$) and anti-neutrinos ($\bar{\nu}$), and I include all three known flavors (electron, muon, and tau). For YeV–XeV neutrino experiments, the differences between these types do not matter much.

3 For $K$ different than 2, these values are different than those used by Kardashev (1964).
from synchrotron and curvature emission sets an upper limit of just 0.1 YeV (Thompson & Lacki 2011). The magnetic fields in actual astrophysical objects limit nucleon energies to ~ 1 ZeV, as the magnetic fields are far too small to create a black hole (Hillas 1984).

Casher & Nussinov (1995, 1997) surmised that the laws of physics forbid us from seeing a Planck particle, arguing no accelerator can be built that reaches those energies. A necessary assumption is that macroscopic amounts of matter cannot reach ultrarelativistic speeds. Otherwise, a particle that “merely” reaches ZeV to YeV energies in some ultrarelativistic flow can reach Planck energy in our frame (or vice-versa). But since then, we have strong evidence that gamma-ray bursts (GRBs) launch flows with bulk Lorentz factors $\Gamma \approx 100$–1000 (Lithwick & Sari 2001; Abdo et al. 2009a). There appears to be no fundamental law against accelerating YeV particles, which are boosted to Planck energies in the flow’s frame. As long as Lorentz invariance holds to Planck scale, reaching Planck energy is technically allowed if extremely difficult. Relativistic boosting also is a possible way around the bounds in Kardashev (1995) and Thompson & Lacki (2011), if the entire central engine itself is boosted to large Lorentz factors. (This is how cosmic strings might produce Planck energy particles, for example; Berezinsky et al. 2009).

A subtlety for colliders built to investigate Planck scale physics is that the relevant quantity is the Mandelstam $s$ (center of mass energy squared), not just the particle energy $E$ (chapter 46 of Olive & Particle Data Group 2014). Because of special relativity, if a relativistic Planck energy particle hits a target particle with mass $m$ at rest, $\sqrt{s}$ is only about $\sqrt{E_{\text{Planck}} m}$. Two particles with Planck scale energy must hit each other in the lab frame to reach $\sqrt{s} \approx E_{\text{Planck}}$. This may be done by accelerating two beams of particles and aiming them at each other.

In any case, it appears that (bottom-up) Planck accelerators require a lot of fine-tuning, and are unlikely to appear in nature although ETs might find ways to engineer them. Of course, even if Planck energy is unachievable, ETs could still be interested in physics at smaller YeV grand unified scales. The basic arguments of this paper still apply, though with weaker constraints on the size and luminosity of accelerators.

### 3. WHY THEY SHOULD BE BIG

Electric fields $E$ accelerate charged particles to higher energies, whereas magnetic fields $B$ merely deflect them without doing work. The dot product of $E$ and $B$ is Lorentz invariant; the component of $E$ parallel to the magnetic field is the same in all frames, but there is some frame where there is no electric field perpendicular to the magnetic field. If $|E \cdot B| > 0$, then the accelerator is basically electrostatic – the particle accelerates along the electric field, gaining energy as it does so. Otherwise, the scattering of the particle by the electromagnetic field is completely elastic in some frame, with no work done on it; the relativistic boost of $B$ is fundamentally the reason for energy transfer to the particle.

The theory of quantum electrodynamics (QED) defines a characteristic electromagnetic field $B_{\text{QED}}^2 = m_e^2 c^3 / (e h) = 4.4 \times 10^{13}$ G (Harding & Liv 2008). The characteristic size of a particle accelerator with electromagnetic fields $B$ that can accelerate a particle of charge $Ze$ to energy $E$ is

$$l_{\text{min}} = \frac{E}{Ze B_{\text{QED}}^2} \approx 9.3 \times 10^{11} \text{ cm} \left( \frac{E}{E_{\text{Planck}}} \right) \left( \frac{B}{B_{\text{QED}}} \right)^{-1}.$$
The characteristic size of a Planck-scale accelerator is greater than 10 R_⊙.

For an electrostatic accelerator (like a linear accelerator), this is truly a fundamental limit as far as we know. The QED vacuum becomes unstable if \( E > B_0^2 \) unless there is a suitable magnetic field. Any attempt to build the electric field to \( B_0^2 \) would simply result in the creation of electron-positron pairs instead (Fedotov et al. 2010). An ET that wished to build an electrostatic accelerator would necessarily be space-faring because the accelerator is too big to fit onto any planet.

With a magnetic accelerator (like a synchrotron), equation \[^2\] corresponds to the Hillas criterion (Hillas 1984). This says that the particle’s gyropериod must be smaller than the light-crossing time of the accelerator region. In principle, the accelerator could be smaller than \( \sim 10 R_⊙ \) if \( B \gg B_0^2 \), a condition reached in magnetars.

No Planck accelerator made of normal atomic matter could be as small as equation \[^2\]. Atoms deform in magnetic fields of \( \sim 10^9 \) G (Harding & Lai 2000). A Planck accelerator with gigagauss electromagnetic fields would be several thousand AUs wide. A minimum electromagnetic field of \( \sim 4 \) mG within the accelerator is set by the size of the Universe, \( \sim 3 \) Gpc \( \approx 10^{27} \) cm. This kind of magnetic field is far bigger than those within the intergalactic medium (IGM; Dolag et al. 2005; Neronov & Vovk 2010; Yoon et al. 2014) or within most galaxies (e.g., Lacki et al. 2010; Beck 2012). Based on the required electromagnetic energy density, a Planck accelerator is probably close to collapsing into a black hole, and black holes might be harnessed (c.f., Kardashev 1995; Vidal 2011; Thompson & Lacki 2011).

4. WHY THEY SHOULD BE BRIGHT

Doing experiments with particle accelerators is not just a matter of reaching the highest energies possible. The other central consideration is the particle luminosity of the accelerator, especially the number flux of particle bunches as they pass through each other. Accelerators are built to find events, particle interactions regulated by physics. Each type of event has a cross section \( \sigma \). The rate at which a class of events occur in the accelerator is proportional to the product of the cross section and the particle luminosity. Thus, the integrated particle luminosity over time is basically the reciprocal of a cross section; the longer the accelerator runs, the more the effective particle luminosity.

The “natural” cross section for quantum processes occurring at an energy \( E \) is

\[
\sigma_{\text{natural}} \sim 4\pi \left( \prod \alpha_i \right) \frac{\hbar c}{E} \left( \frac{\hbar c}{E} \right)^2 \approx \frac{1}{\pi} \left( \prod \alpha_i \right) \frac{\hbar c}{E} \left( \frac{\hbar c}{E} \right)^2
\]

(3)

where the \( \alpha \) are coupling constants of each force involved, and \( n \) is the order of each coupling, the number of each kind of vertex in the Feynman diagram (e.g., Halzen & Martin 1984). When \( E \) is the mass energy of some force carrier, then \( (\hbar c/E) \) is its Compton wavelength. At Planck energy and \( \alpha \approx 1 \), the natural cross section is approximately the Planck area:

\[
\sigma_{\text{Planck}} \equiv 2G \frac{c^3}{n^2} = 3.3 \times 10^{-55} \text{ cm}^2
\]

(4)

Note that this is approximately the surface area of a black hole

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[^1]: I will consistently refer to this as “particle luminosity”. “Luminosity” on its own refers to the astrophysical meaning of energy luminosity in this work.

[^2]: Extrapolating to Planck energy, \( \alpha_e = 1 \).
energies. If ETs wish to get good statistics on Planck scale events, they may want to study thousands of events, requiring that much more energy.

The brightness of the accelerator, its waste heat, and its YeV–XeV radiation depends on how long it runs. I assume the experiments would last $t_{\text{run}}= 4.4 \times 10^{17}$ s at longest. Galaxies evolve on Gyr timescales, and the supplies of power and mass might vanish around the accelerator if the ETs took much longer. A more natural timescale might be the “Fermi” timescale to cross a galaxy and establish a K3 society, about 1 Myr. The luminosities for these timescales are also listed in Table 2. A hadronic Planck accelerator would need to be K3 for a run time of 1 Myr, and a neutrino collider would still be K2 over that duration.

At the opposite extreme, the accelerator lasts at least one light crossing time, and the smallest an accelerator could be is a Schwarzschild radius. The minimum run time is then $\sim \sqrt{\Psi \Gamma / c}$, with the accelerator reaching the maximal Einstein luminosity of $L_E = c^3 / G = 3.63 \times 10^{59}$ erg s$^{-1}$ (K4.6; Hartle 2003). These kinds of accelerators are brief transients lasting between a fraction of a millisecond and a few hours (c.f., Thompson & Lacki 2011).

5. HOW THEY SHOULD SHINE

5.1. Conditions in the accelerator region

Which kinds of radiation the accelerator could emit depends on the conditions in the accelerator region. Any YeV particles generated in collisions could radiatively cool to low energies if the accelerator is filled with magnetic fields, radiation, or baryons. If the densities are high enough, all of the energy could be thermalized. But if the accelerator is optically thin to at least some YeV particles, it shines directly in YeV–XeV radiation from relativistic transients like Gamma Ray Bursts (Rees & Meszaros 1992; Waxman 1995; Thompson & Lacki 2011). The outflow is powered by a central engine with luminosity $L$ and a variability timescale $\delta t$. It emits an outflow that expands with bulk Lorentz factor $\Gamma$ that covers $4\pi \Psi$ steradians (so $\Psi = 1$ is isotropic). The outflow is unsteady, with different parts having slightly varying speeds. The irregularities crash into one another and generate internal shocks when the outflow has expanded to a characteristic size $r = \Gamma^2 c \delta t$. It is in these internal shocks that acceleration occurs.

The accelerator region moves outward with the bulk flow. I define the accelerator frame to be comoving with the bulk flow. Quantities in the accelerator frame are marked with an apostrophe ('). The characteristic cov-ering fraction is $\psi_T = 1 / (2 \Gamma^2)$, and the energy density is

$$u' = \frac{L}{\psi_T \Gamma^4 c^2} = \frac{L}{\psi_T \Gamma^6 c^3 \delta t^2}.$$  

The typical energy density in the flow is

$$u = 1 \times 10^7 \text{ erg cm}^{-3} \Psi^{-1} \psi_T^{-1} \delta t^{-2} \left( \frac{L}{L_E} \right),$$

where $\Gamma = \Gamma_1 / 1000$ and $\delta t = \delta t_s / (1 \text{ s})$. For an isotropic outflow, the highest proton energies are achieved in outflows with $\Gamma \approx 2000$ (Thompson & Lacki 2011).

Instead of being spherical, the outflow may be highly beamed with an opening angle $\leq 1 / \Gamma$. The characteristic covering fraction is $\psi_T = 1 / (2 \Gamma^2)$, and the energy density is

$$u = \frac{L}{2 \pi \Gamma^4 c \delta t^2} = 2000 \text{ erg cm}^{-3} \psi_T^{-1} \Gamma_6^{-4} \delta t_s^{-2} \left( \frac{L}{L_E} \right),$$

deriving $\delta t_s = \Gamma / 10^6$ and $\psi_T = \Psi / \psi_T$. Plank particles can then be accelerated for extreme outflows with $\Gamma > 10^6$ (Thompson & Lacki 2011).

5.1.1. Escape from the accelerator or decay inside it?

The accelerator region is a shell with thickness of $\delta r' = \Gamma c \delta t$. The minimum possible dynamical time is just the light-crossing time of the shell, $t_{\text{dyn}} = \Gamma \delta t = 1000 \text{ s} \Gamma \delta t_s$. This time is also the typical time it takes a CR to escape the accelerator.

### Table 2

Accelerator power to reach $\sigma_{\text{Planck}}$

| Beams | Limiting interaction\(^ a \) | Process | $\sigma_{\text{Planck}}(\text{mb})$ | $\mathcal{O}(\sigma s)$ | $\mathcal{Y}$\(^ b \) | $\mathcal{Y}_{/H}$ | $\mathcal{Y}_{/\text{Myr}}$ | Notes |
|-------|------------------|---------|-----------------|-----------------|-----------------|----------------|----------------|-------|
| $pp$  | $p + p \rightarrow \pi + \text{anything}$ | $\sim 2000$ | $\ln^2 s$ (?) | $1 \times 10^{60}$ | 70 | $3 \times 10^{36}$ | $7 \times 10^4$ | $4 \times 10^{42}$ | $1 \times 10^6$ | (c) |
| $p\gamma$ | $p + \gamma \rightarrow \pi + \text{anything}$ | $\sim 7$ | $\ln^2 s$ (?) | $4 \times 10^{35}$ | 0.2 | $9 \times 10^{35}$ | $200$ | $1 \times 10^{40}$ | $3 \times 10^6$ | (d) |
| $\mu^\pm \mu^\pm$ | $\mu^+ + \mu^- \rightarrow \mu^+ + \mu^- + \text{anything}$ | $1000$ | $\ln^2 s$ | $6 \times 10^{45}$ | $30$ | $1 \times 10^{38}$ | $3 \times 10^4$ | $2 \times 10^{31}$ | $5 \times 10^7$ | (e) |
| $\gamma \gamma$ | $\gamma + \gamma \rightarrow \mu^+ + \mu^- + \text{anything}$ | $0.06045$ | $1$ | $4 \times 10^{30}$ | $2 \times 10^{-4}$ | $9 \times 10^{32}$ | $0.2$ | $1 \times 10^{37}$ | $3000$ | (g) |
| $\nu \bar{\nu}$ | $\nu + \bar{\nu} \rightarrow \ell + \bar{\nu}$ | $2.2 \times 10^{-7}$ | $1$ | $1 \times 10^{49}$ | $7 \times 10^9$ | $3 \times 10^{28}$ | $8 \times 10^{-6}$ | $4 \times 10^{32}$ | 0.1 | (i) |
| $\nu \bar{\nu}$ | $\nu + \bar{\nu} \rightarrow \ell + \bar{\nu}$ | $7.0 \times 10^{-8}$ | $1$ | $4 \times 10^{35}$ | $2 \times 10^{-9}$ | $9 \times 10^{27}$ | $2 \times 10^{-6}$ | $1 \times 10^{32}$ | 0.03 | (i) |

\(^ a \) The cross sections are evaluated at $s = E_{\text{Planck}}^2$.

\(^ b \) The energy needed to create a Schwarzschild radius. The minimum run time is then the light crossing time, and the smallest an accelerator could be $K_2$ over that duration. The size and variability timescale as well as the possibility of a slightly more rapidly growing term, are disputed. See, for example, Azimov (2011); Block & Halzen (2011); Fagundes et al. (2013); Anisovich et al. (2013a,b). From Brown et al. (1973).

\(^ c \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.

\(^ d \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.

\(^ e \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.

\(^ f \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.

\(^ g \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.

\(^ h \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.

\(^ i \) The cross sections are estimated from the full planckian interactions from Table 2. A hadronic Planck accelerator would need to be $K_3$ for a run time of 1 Myr, and a neutrino collider would still be $K_2$ over that duration.
and the maximum time it takes a CR to complete one Larmor orbit (Thompson & Lacki 2011). Particle accelerators naturally create a slew of unstable particles, including muons, mesons, high energy baryons, and gauge bosons (Tables 3 and 4). If their lifetime is short enough, they decay before they can escape the accelerator. Since the unstable particles are time dilated, their decay time in the accelerator frame is

\[ t'_{\text{decay}} = \frac{E}{\Gamma m'_{\text{c}}} \times 10^4 \text{s} \]

Here, \( \tau \) is the rest lifetime of the particle and \( E' \) is the particle (kinetic) energy in the accelerator frame. A particle observed to have energy \( E \) in the engine frame has \( E' = E/\Gamma \). For compactness, I use the auxiliary variables \( \tau_{\text{ms}} = [\tau/(1 \text{ ns})] \), \( E_{28} = [E/(10^{38} \text{ eV})] \), and \( m_{\text{GeV}} = [m/(1 \text{ GeV})] \).

We see most unstable particles decay before they can escape. The minimum energy threshold for escape to win is

\[ E'_{\text{min}} = \frac{\Gamma^2 \delta t m_{\text{c}}^2}{\tau} = 1 \text{ YeV} \frac{\Gamma^2 m_{\text{GeV}} \delta t_{\text{ms}}}{1 \text{ ns}} \]

5.1.2. Synchrotron cooling

The same electromagnetic fields that accelerate charged particles also induce synchrotron losses. As long as the magnetic field isn’t too high (in the Thomson regime), the cooling time \( t'_{\text{synch}} \) decreases with energy as charged particles radiate faster. The more rapid synchrotron cooling is in fact one of the main obstacles to building a Planck accelerator (Thompson & Lacki 2011). Indeed, synchrotron losses are one of the main limiting factors in current accelerators as well. The pitch-angle averaged synchrotron cooling time of a relativistic charged particle in the accelerator is

\[ t'_{\text{synch}} = \frac{9}{32 \pi} \frac{(m'_{\text{c}})^2 \delta t_{\text{ms}}}{Z^4 e^4 c^4 E^4 B^4} \]

in the Thomson regime. Technically, the emission mechanism could be “diffusive synchrotron” or “jitter” radiation if the magnetic field is very tangled on small scales, but the energy loss time is still the same (e.g., Kelner et al. 2013). If we assume that the magnetic energy density is \( u_B = \epsilon_B u' \), then the synchrotron cooling time is

\[ t'_{\text{synch}} = \frac{22}{Z^4 E_{28}^4 \epsilon_B} \left( \frac{L}{L_E} \right)^{-1} \]  \hspace{1cm} (11)

For a beamed flow, the cooling time is

\[ t'_{\text{synch}} = \frac{11}{Z^4 E_{28}^4 \epsilon_B} \left( \frac{L}{L_E} \right)^{-1} \]  \hspace{1cm} (12)

Comparing to the dynamical time of the outflow, it is clear that the synchrotron cooling time is much shorter unless \( \Gamma = \delta t \) is very big. Stable charged particles cannot escape the accelerator region unless their engine-frame energy is less than

\[ E_{\text{esc}} = \frac{9}{8} \frac{\Psi \epsilon^2 \delta t_{\text{ms}} (m'_{\text{c}})^2}{Z^4 e^4 B^4 L^4} = 220 \text{ PeV} \frac{\Psi m^2_{\text{GeV}} \delta t_{\text{ms}}}{Z^4 e^4 B^4} \left( \frac{L}{L_E} \right)^{-1} \]

\[ = 110 \text{ ZeV} \frac{\Psi m^2_{\text{GeV}} \delta t_{\text{ms}}}{Z^4 e^4 B^4} \left( \frac{L}{L_E} \right)^{-1} \]  \hspace{1cm} (13)

Figure 11 shows the energies and output \( \Gamma \) where synchrotron cooling is important for electrons and protons in a fiducial maximal accelerator. YeV electrons cool essentially instantly by synchrotron radiation, especially if the flow is beamed. This demonstrates how hard it is to directly accelerate electrons to these energies, although they can be created as secondaries from other particles. The situation is somewhat better for protons. Planck energy protons can escape before cooling through synchrotron if \( \Gamma \gtrsim 3 \times 10^4 \) (3 x 10^7) for an isotropic (highly beamed) accelerator.

For unstable charged particles, synchrotron cooling is faster...
The classical condition is broken when $B'$ is greater than

$$B'_{\text{QED}} = \frac{m^2 c^3}{\hbar Z e'}$$

or equivalently when the particle energy is bigger than

$$E_{\text{synch}} = \frac{\sqrt{\pi} \Psi c (mc^2)^3 \delta t^4}{\hbar Z e' L}$$

These conditions are displayed in Figure 2 for some commonly produced particles in accelerators. The synchrotron losses are particularly bad for muons, which is important since they are a common decay product of hadrons and one of the main sources of neutrinos. Mesons like pions are marginally better, since they have a shorter lifetime and larger masses (Table 3). The heavier charmed ($D^\pm$, $D_s^\pm$) and beautiful mesons ($B^\pm$) and baryons ($\Lambda^\pm_c$), as well as the $\tau$ leptons, face less stringent losses. These “merely” require outflows with $\Gamma \lesssim 10^4$ $(10^5)$ to decay before cooling at Planck energy inside isotropic (extremely beamed) accelerators, emitting a hard component of prompt neutrinos (Enberg et al. 2009). Even the $W$ bosons face synchrotron losses at Planck energy, unless the outflow is highly relativistic.

The QED Limit – All of the above applies for particles in the Thomson regime. This means that (1) the particle’s motion is essentially classical, and the energy of emitted photons is much smaller than the particle’s kinetic energy, and (2) the particle can be treated as a point charge (e.g., Rvibicki & Lightman 1979). In the Weizsäcker-Williams approach, synchrotron radiation is basically the scattering of a virtual photon by the Inverse Compton effect (Lieu & Axford 1993). Much of the same physics that applies in scattering of photons also applies to synchrotron emission.

The classical condition is broken when $B'$ is greater than $B'_{\text{QED}} = \frac{m^2 c^3}{\hbar Z e'}$, or equivalently when the particle energy is bigger than

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Figure 2. Synchrotron cooling regimes for various common unstable particles. The shading and line styles are the same as in Figure 1. In addition, the red solid lines indicate $E_{\text{esc}}$: below the line, the boosted particle lifetime is shorter than the escape time. I assume that $\delta t = \epsilon B = \epsilon_{\text{rad}} = \sigma_{\text{mb}} = L_{E} = 1$.

Production of mesons (pions, kaons, and eta) from protons emitting synchrotron (Tokuhisa & Kajino 1999; Herpay & Patkós 2008; Kajino et al. 2014) and curvature (Berezinsky et al. 1995; Herpay et al. 2008; Fregolente & Saa 2012) radiation. But no full QED/QCD calculation has been done, so the energy losses of hadrons are essentially unknown for $E \gg E_{\text{QED}}^{\text{synch}}$.

Equation 15 is still the limit if the photon resolves the hadron when its wavelength in the hadron frame is shorter than the hadron’s Compton wavelength. The actual magnitudes of charge radii of hadrons are poorly known and discrepant, but are roughly a fraction of a fm, or $\sim 1$ GeV in natural units (e.g., Amendolia et al. 1984, 1986; Eschrich et al. 2001; Hwang 2002; Pohl et al. 2010). Since the hadrons I consider have masses of order 0.1 – 10 GeV, I assume this is an adequate approximation.

The particle energies and outflow $\Gamma$ outside the Thomson regime are cross-hatched in Figure 1 for electrons and protons and Figure 2 for unstable particles. It is conceivable that the synchrotron losses are slow enough in these regions that particles could escape or decay before cooling, but this awaits full QED and QCD calculations. As we shall see, other losses probably intervene in these regions anyway.
5.1.3. Baryonic cooling

Aside from electromagnetic fields, an outflow can contain matter within it. Thermal protons and neutrons within the accelerator region represent another obstacle for any particle trying to escape. Escaping or decaying hadrons can be cooled by direct collisions with thermal nucleons, whether or not they are charged. CR electrons and positrons lose their energy catastrophically to bremsstrahlung emission when they encounter baryons. Likewise, opacity to γ-rays arises from thermal baryons due to pair production. The typical grammage for all of these processes is several tens of g cm⁻², which I will take to be $\Sigma_{\text{loss}} = 50 \Sigma_{\text{g}}$ g cm⁻².

The density of baryons is found from the energy density (equation 6). Let $\epsilon_{\text{bar}} = \epsilon_{\text{bar}} \epsilon$ be the energy density in baryons. Then the baryonic density is

$$n_{\text{bar}}' = \frac{\epsilon_{\text{bar}} L}{4\pi \Psi \Gamma \delta \tau} = 7.1 \times 10^{11} \text{ cm}^{-3} \frac{\epsilon_{\text{bar}}}{\Psi \Gamma \delta \tau} \left( \frac{L}{L_E} \right).$$

The grammage through the outflow’s shell is $\Sigma' = n_{\text{bar}}' m_\Sigma \delta \tau'$. By setting $\Sigma' = \Sigma_{\text{loss}}$, I define a minimum $\Gamma$, below which particles are cooled by baryons:

$$\Gamma_{\text{esc}} = \left[ \frac{\epsilon_{\text{bar}} L}{4\pi \Psi \Sigma_{\text{loss}} \delta \tau} \right]^{1/5} = 940 \left[ \frac{\epsilon_{\text{bar}}}{\Psi \delta \tau, \Sigma_{\text{g}}} \left( \frac{L}{L_E} \right) \right]^{1/5} = 1.1 \times 10^5 \left[ \frac{\epsilon_{\text{bar}}}{\Psi \delta \tau, \Sigma_{\text{g}}} \left( \frac{L}{L_E} \right) \right]^{1/5}. \quad (16)$$

Baryonic cooling traps stable particles of all energies in slow outflows, but generally synchrotron losses are more important.

Unstable particles that would decay before escaping (see discussion in section 5.1.1) traverse only $\Sigma' = c / \tau' n_{\text{bar}} m_p = E c \tau n_{\text{bar}} m_p (\Gamma / \Gamma_m)$. The minimum $\Gamma$ in these cases are:

$$\Gamma_{\text{esc}} = \left[ \frac{\epsilon_{\text{bar}} L \tau}{4\pi \Psi \Sigma_{\text{loss}} \delta \tau (m_e c^2)} \right]^{1/7} = 3600 \left[ \frac{\epsilon_{\text{bar}} E_{\gamma} \tau_{\text{ne}}}{\Psi \delta \tau, \Sigma_{\text{g}}} \left( \frac{L}{L_E} \right) \right]^{1/7} = 1.1 \times 10^5 \left[ \frac{\epsilon_{\text{bar}} E_{\gamma} \tau_{\text{ne}}}{\Psi \delta \tau, \Sigma_{\text{g}}} \left( \frac{L}{L_E} \right) \right]^{1/5}. \quad (17)$$

This again confirms that synchrotron losses are probably more important.

Muons, $\tau$ leptons, and neutrinos are far more penetrative than other particles. At Planck energies, their stopping length is $\Sigma \gtrsim 10^6$ g cm⁻² (Iyer Dutta et al. 2001). Because of the small dependence of the critical $\Gamma$ on $\Sigma$, the estimates are not changed much for them.

5.1.4. Cooling by thermal radiation

The intense activity in relativistic outflows can also generate a lot of heat. The heat manifests mainly as photons. The radiation energy density is some fraction $\epsilon_{\text{rad}}$ of the total energy density. If the heat radiation is thermal, it has a charac-teristic temperature

$$T' = \left[ \frac{\epsilon_{\text{rad}} L}{4\pi \Psi \epsilon_{\text{rad}} \delta \tau} \right]^{1/4} = 6.1 \times 10^7 \text{ K} \left[ \frac{\epsilon_{\text{rad}}}{\Psi \delta \tau, \Gamma_{\text{esc}}} \left( \frac{L}{L_E} \right) \right]^{1/4} = 2.3 \times 10^4 \text{ K} \left[ \frac{\epsilon_{\text{rad}}}{\Psi \Gamma_{\text{esc}}} \left( \frac{L}{L_E} \right) \right]^{1/4}. \quad (18)$$

Blackbody radiation has a number density $n_\gamma' = \frac{n_e' / (\pi^4 k_B T' / (30 \zeta(3)))}{\approx 2 / (2.7 k_B T')}$ (Kolb & Turner 1990), so the photon density in the fireball is

$$n_\gamma' = \frac{30 \zeta(3)}{\pi^4 k_B} \left( \frac{\epsilon_{\text{rad}} L}{4\pi \Psi \Gamma_{\text{esc}} \delta \tau} \right)^{3/4}. \quad (19)$$

This is generally $10^{14.5} - 10^{18.5}$ cm⁻³ for the parameters I have been using. For leptons, the interaction with radiation results in Inverse Compton scattering, which behaves similarly to synchrotron cooling if the magnetic energy density is replaced by the radiation energy density. Since the photon energy is a few eV and the particle energy is YeV scale, leptonic Inverse Compton is deep in the QED regime.

Hadrons interact with the photons with some cross section $\sigma = 10^{-2} \sigma_{\text{mb}}$ cm². The cross section for hadrons interacting with photons is of order a mbarn (Olive & Particle Data Group 2014). Generally, $\kappa \approx 0.1$ for these hadronic processes, so a hadron could survive multiple collisions with photons, but I ignore this factor since it extends the interaction time by only a factor of a few. Stable particles escape if the optical depth $\Theta_{\text{esc}} = n_\gamma' \sigma \delta \tau'$ is less than 1. I find that the optical depth is

$$\Theta_{\text{esc}} = 1.4 \times 10^5 \left( \frac{\epsilon_{\text{rad}} L}{L_E} \right)^{3/4} \Gamma_{\gamma}^{-3/2} \delta t_{\gamma}^{-1/2} \psi^{-3/4} \sigma_{\text{mb}}. \quad (20)$$

The fireball is clear enough to let particles escape only if $\Gamma$ is bigger than

$$\Gamma_{\gamma} = \left( \frac{30 \zeta(3)}{\pi^4 k_B} \right)^4 \left( \frac{\epsilon_{\text{rad}} L}{4\pi \Psi} \right) \left( \frac{\epsilon_{\text{rad}} L}{4\pi \Psi} \right)^3 c^{-5} \delta \tau^{-2} \Gamma_{\text{esc}}^{1/4} = 3.0 \times 10^4 \left( \frac{\epsilon_{\text{rad}} L}{4\pi \Psi} \right)^{1/4} \left( \frac{L}{L_E} \right)^{3/14} \Gamma_{\gamma} = 8.6 \times 10^7 \left( \frac{\epsilon_{\text{rad}} L}{4\pi \Psi} \right)^{1/8} \left( \frac{L}{L_E} \right)^{3/8}. \quad (21)$$

The regions left of the gold lines in Figure 9 are opaque. The optical depth unstable hadrons experience over their lifetime is $\Theta_{\text{esc}} = n_\gamma' \sigma E / (\Gamma m_c c^2)$. In order for photon scattering to be negligible, the Lorentz factor should be bigger.
than
\[
\Gamma_{\text{decay}} = \left[ \frac{30\zeta(3)\sigma}{\pi^4k_B} \frac{E\tau}{mc^2} \right]^4 \frac{\Gamma_{\text{rad}}\Omega_{\text{SN}}}{4\pi\Psi} \frac{3}{c^5} \delta t^{-2} \right]^{1/22} = 4.6 \times 10^4 \frac{\sigma_{\text{mb}}\Gamma_{\text{rad}}}{\Psi} \frac{E_{28}^2}{m_{\text{GeV}}}^{1/22} \frac{1}{L} \frac{3}{22}
\]
\[
= 3.0 \times 10^6 \frac{\sigma_{\text{mb}}\Gamma_{\text{rad}}}{\Psi} \frac{E_{28}^2}{m_{\text{GeV}}}^{1/16} \frac{1}{L} \frac{3}{16}.
\]

The thermal radiation probably is the biggest obstacle for the escape of protons and neutrons from the fireball, unless its luminosity is low or the particle energy is high. That is simply because there is so many photons. In the highly beamed case, it may be necessary for the fireball outflow to reach \( \Gamma \gtrsim 10^8 \) for long-lived hadrons to escape. The regimes of effective photon cooling of mesons and heavy baryons (gold lines in Figure 2) are fairly similar to the synchrotron cooling regimes (black lines), especially for Planck energy particles.

5.1.5. Summary

The calculations in the previous subsections all basically highlight the compactness problem. When an astrophysical explosion contains a lot of energy, the fireball is so dense that it is opaque unless it is highly relativistic (Cavaliere & Rees 1978). If Planck accelerators need luminosities comparable to the Einstein luminosity, the compactness problem is especially severe. Although outflows with \( \Gamma \) of several hundred are actually achieved in GRBs (Lithwick & Sari 2001; Abdo et al. 2009), even that may not be enough to release Planck energy particles, with \( \Gamma \gtrsim 10^3 \) required for the fireball to be sufficiently transparent for radiation to escape.

Although the internal shocks model I use is applicable to astrophysical transients, it assumes that the accelerator is very chaotic and inefficient. Basically, a fireball model describes a bomb being set off, with the ETs presumably just watching the particles that it produces. An accelerator built by ET may be much “cleaner”, with much less energy density. After all, the ETs themselves need to observe the collisions. Arguably both approaches have precedent in our own history. Traditional particle accelerators like the LHC are very clean, so that the events can be analyzed easily. But in the 1950s and 1960s, nuclear weapons were detonated in space to test the effects of energetic particles in the Earth’ magnetosphere (these tests thankfully have ceased).

5.2. Conditions between the accelerator and Earth

Even if it does escape the accelerator, YeV radiation still does not necessarily reach us. In order to do that, it must (at least) escape from the accelerator’s host galaxy (if any), traverse intergalactic space, and then travel within the Milky Way to Earth. Particles can be blocked from reaching Earth if they decay before making it here, if they cool radiatively, or if they interact with photons.

The magnetic fields and the extragalactic radiation fields in the IGM and the Milky Way are the biggest obstacles for YeV particles. I take the Milky Way to have a magnetic field of \( B = 6 \mu \text{G} \), which extends over 10 kpc (Strong et al. 2000; Beck 2001). Presumably our Galaxy is typical of galaxies hosting ETs. But note that at \( z \approx 1–2 \), there was a population of luminous infrared main-sequence galaxies that contained most star formation. These probably had magnetic fields larger than the Milky Way’s, perhaps of order \( \sim 100 \mu \text{G} \) (Lacki & Thompson 2010), judging from the fact that they lie on the far-infrared–radio correlation observed for star-forming galaxies (e.g., Mao et al. 2011). On the other hand, ETs may live in quiescent, red galaxies, which probably have somewhat weaker magnetic fields than the Milky Way (Moss & Shukurov 1996).

Unlike CRs at ZeV and lower energies, particles at YeV energies are not deflected significantly by magnetic fields. Their gyroradii are

\[
R_L = 180 Z^{-1} \text{ Mpc} \left( \frac{E}{1 \text{ YeV}} \right) \left( \frac{B}{6 \mu \text{G}} \right)^{-1}.
\]

Even if the Galaxy’s magnetic field is perfectly regular, it deflects a YeV particle by at most \( Z/180 \mu \text{G} \). The intergalactic magnetic field is very poorly known, but I scale to \( B_{\text{IGM}} = 1 \mu \text{G} \) with coherence lengths of \( \ell_{\text{IGM}} = 1 \text{ Mpc} \). The average deflection of a YeV CR is \( \sqrt{D_{\text{IGM}}(Z_{\text{IGM}}/R_L)} = \sqrt{D_{\text{IGM}} Z_{\text{IGM}}/E} = 6/5 \times 10^7 Z_{\text{Gpc}}^{-1/2} (\ell/\text{Mpc})^{-1/2} (B_{\text{IGM}}/\mu \text{G}) (E/\text{YeV})^{-1} \). Thus, even charged particles should point directly back at their sources to high precision. The limiting factor on the angular resolution is the ability of experiments to reconstruct the trajectory of the CR.

By far, most extragalactic photons are part of the Cosmic Microwave Background (CMB), with a number density \( n_{\text{CMB}} = 411 \text{ cm}^{-3} \) (Olive &Particle Data Group 2014). The extragalactic background light (EBL) from galaxies at ultraviolet to infrared wavelengths has a much smaller energy density than the CMB (e.g., Franceschini et al. 2008), and each photon contains much more energy. Nor is the radio background a problem, whether in the Galaxy or outside of it. The Galactic radio brightness temperature peaks at \( \sim 10^7 \text{ K} \) around 2 MHz (Brown 1973; Fleishman & Tokarev 1995; Manning & Dulk 2001), implying a photon number density of \( \sim 1 \text{ cm}^{-3} \). Strictly speaking, the extragalactic radio background is very poorly constrained at frequencies below 1 MHz, with brightness temperatures of \( \gtrsim 10^{12} \text{ K} \) possible (Lacki 2010). The expected radio background at these frequencies, however, is thought to be quite small (Protheroe & Biermann 1996).

5.2.1. Could we directly detect beam nucleons?

Planck accelerators are likely to produce a lot of pollution when beam particles collide with each other, according to the arguments in Section 4. But maybe not all beam particles are consumed by these collisions, with some escaping the accelerator. Perhaps the acceleration process is highly random, with only some of the CRs being “harvested” for collisions. Another possibility is that accelerated particles are released during experimental runs. In our colliders, beam particles are directed into a beam dump when the experiment is over. The beam carries a lot of energy, which poses engineering challenges (Schmidt et al. 2000). With a Planck accelerator, the amount of energy in the beams is literally cosmic, so a beam dump may not be practical. Instead, the beam particles might simply be released into interstellar space. As noted in Section 5.1.4, the collider must be clear of thermal photons in order for beam nucleons to escape.

Nucleons experience the Greisen-Zatsepin-Kuz’min (GZK; Greisen 1966; Zatsepin & Kuz’min 1966) effect as they propagate over intergalactic distances, in which photodispro-
actions occur with CMB photons. The effect occurs at energies above 40 EeV, and has a typical energy loss scale of \( \sim 20 \) Mpc. This GZK “horizon” for UHECRs is well known \cite{Bhattacharjee2000}. The reactions produce pions, which decay into neutrinos, \( e^\pm \), and \( \gamma\)-rays. The resulting population of GZK neutrinos are themselves sought in many experiments (section 5.2.4).

In addition, protons at YeV energies experience synchrotron cooling in any intergalactic magnetic fields. Nano gauss magnetic fields cool protons over a timescale \( t_{\text{sync}} = 140 \) Tyr \((E/\text{YeV})^{-1}(B/\text{nG})^{-2}\), which conceivably matters for Planck energy protons. If the intergalactic magnetic fields are picogauss or lower, as often thought, then synchrotron cooling poses no problem.

Protons traversing the Galactic magnetic field have a cooling time of \( t_{\text{sync}} = 4 \) Myr \((E/\text{YeV})^{-1}(B/\text{B}_{\text{MW}})^{-2}\). Thus, synchrotron cooling prevents protons with energies above \( \sim 100 \) YeV from reaching Earth without severe energy losses. Instead, most of the CR proton energy would be \( \sim 1 \) PeV from reaching Earth without severe energy losses. Instead, most of the CR proton energy would be \( \sim 1 \) PeV. If the intergalactic magnetic fields are \( \sim 10^{-8} \) G, the protons traverse the Galactic magnetic field in \( \sim 100 \) yrs. This cooling time is \( \sim 10^5 \) Gpc distance.

Protons traversing the Galaxy have a cooling time of \( t_{\text{sync}} = 30 \) Gyr \((E/\text{YeV})^{-1}(B/\text{B}_{\text{MW}})^{-2}\). Thus, protons traversing the Galactic magnetic field have a cooling time of \( t_{\text{sync}} = 4 \) Myr \((E/\text{YeV})^{-1}(B/\text{B}_{\text{MW}})^{-2}\). Cooling in any intergalactic magnetic fields. Nanogauss energy protons. If the intergalactic magnetic fields are picogauss or lower, as often thought, then synchrotron cooling poses no problem.

While it takes a column of \( \sim 80 \) g cm\(^{-2}\) to stop a GeV proton in hydrogen, the inelastic nucleon-nucleon cross section increases with energy. When a YeV nucleon hits a proton at rest, \( \sqrt{s} = 30 \) PeV \((E/\text{YeV})^{1/2}\), and the approximate inelastic hadron-hadron collision cross section is \( \sim 300 \) mb \cite{Olive2014}. In other words, it takes \( \sim 5 \) g cm\(^{-2}\) to stop a YeV proton in hydrogen, a column attained only in the most extreme starburst regions and in Compton-thick AGNs. The Milky Way’s gas poses no obstacle for these CRs.

5.2.2. Unstable hadrons and leptons

Hadronic colliders make a vast number of mesons, which are another possible type of radiation from Planck accelerators. Mesons in flight frequently decay into muons, with a lifetime of a few microseconds. The lifetimes of these unstable particles are time dilated so that \( \gamma\tau = 0.2 \) pc \((E/\text{YeV})(\tau/10 \) ns) on average (Tables 3 and 4). The longest-lived neutral mesons, particularly the \( K^0\), might propagate a distance of a few kiloparsecs from their sources.

Charged mesons and muons experience synchrotron losses before they can decay, however, limiting their ranges to

\[
ct_{\text{max}} = \left[ \frac{9}{4} \frac{m_e^2 c^3 \tau c}{Z e^2 B^2} \right]^{1/2} \approx 0.1 \text{ kpc} \frac{m_{\text{GeV}} \tau_{\text{ms}}}{B \text{ (6 \mu G)}} \cdot
\]

The maximum range is attained when the particle has an energy

\[
E_{\text{farthest}} = 1 \text{ YeV} \frac{m_{\text{GeV}} \tau_{\text{ms}}}{\tau_{\text{ms}} \text{ (6 \mu G)}} \cdot
\]

Muons with an energy of 1 YeV have a range of 200 pc in the Galaxy. Charged pions only make it 30 pc (at 20 YeV), while kaons may be able to go 150 pc at 600 YeV. The heavier hadrons listed in Table 3 all have ranges of less than 100 pc, reached only when their energy is many YeV.

Given the apparent lack of cosmic engineering in the Solar neighborhood, it is unlikely that these particles can be directly detected at Earth.

5.2.3. Electromagnetic cascades from \( e^\pm \) and \( \gamma\)-rays

Although there are many ways of generating \( \gamma\)-rays and \( e^\pm \) in colliders, including annihilation, pair production, and meson decay, they also strongly couple to the radiation and magnetic backgrounds of the IGM and the Galaxy. Photons convert into \( e^\pm \) when they annihilate with background photons, while \( e^\pm \) convert into photons when they cool by inverse Compton scattering, synchrotron, or triplet pair production \((\text{TPP};\text{ MASTICHADIS}1999;\text{ ANGULEV} et al.1999)\). During each conversion, the particle keeps most but not all of its energy. Thus, over many conversions, the particle’s energy cascades into lower energy forms. Conventionally, it is thought that any energy in ultra high energy (UHE) photons or \( e^\pm \) cascades down into MeV–GeV \( \gamma\)-rays over Gpc distances. The unresolved \( \gamma\)-ray background then constrains the amount of power being generated in UHE \( \gamma\)-rays and \( e^\pm \) in the Universe \cite{Coppi1997, Murase2012}.

There are some subtleties, though, when the primary \( \gamma\)-ray or \( e^\pm \) energy is YeV or above. At these energies, \( \gamma\)-rays annihilate with CMB photons through the double pair production process, with an interaction length of \( 120 \) Mpc \cite{Bhattacharjee2000}. YeV–XeV \( e^\pm \) interact with the CMB through TPP. But while the TPP cross section is large, its inelasticity is extremely small, \( \approx 10^{-7} \), so the energy attenuation length of \( e^\pm \) is \( > 10 \) Gpc for these particles \cite{Bhattacharjee2001, Protheroe1996}.

Cascades can still proceed at these energies through synchrotron cooling, if there are magnetic fields anywhere along the line-of-sight to the source. Picogauss magnetic fields cause electrons to lose energy in just \( t_{\text{sync}} = 13 \) yr \((E/\text{YeV})^{-1}(B/\text{nG})^{-2}\). While the volume-filling intergalactic magnetic field might be as small as \( 10^{-19} \) G \cite{Neronov2010, Dermor2011}, the necessary picogauss to nanogauss magnetic fields may be present in large-scale structures like galaxy filaments \cite{Dolag2005}. An \( e^\pm \) traveling over cosmic distances is likely to hit one of these filaments and then cool. There are probably also femtogauss magnetic fields generated by thermal fluctuations throughout the IGM \cite{Schlickeiser2012, Yoon2014}. The synchrotron \( \gamma\)-rays emitted by the \( e^\pm \) have characteristic energies of 0.5 keV \((E/\text{YeV})^2(B/\text{nG})\).

Any YeV photon that gets near the Milky Way is destroyed by pair production on the magnetic field, \( \gamma + B \rightarrow e^+ + e^- \) \cite{Stecker2003}. The threshold for this process is

\[
E_{\text{thresh}} = \frac{B_{\text{QED}}^2}{B_{\text{MW}}} m_e^2 = 4 \text{ YeV} \left( \frac{B_{\text{MW}}}{6 \text{ \mu G}} \right)^{-1} \cdot
\]

This initiates an electromagnetic cascade. At 1 YeV, the cooling time for \( e^\pm \) is just 10 seconds in the Thomson regime; higher energies enter the QED regime, but the cooling time increases slowly with energy. Thus, the energy in YeV–XeV photons or \( e^\pm \) downgrades into ZeV photons within the Galaxy. These can be directly detected by CR arrays like Auger. So far, no showers from photons have been detected, setting a stringent limit on their flux, much more strin...
gent than the GeV $\gamma$-ray background implies (Rubtsov et al. 2006; Abraham et al. 2009).

Clearly, what happens to YeV–XeV photons and $e^\pm$ needs to be studied in more detail, but they have two possible fates. If they come from distant extragalactic sources and there are femtogauss–picogauss magnetic fields along the line of sight, they cascade into MeV–GeV synchrotron $\gamma$-rays (Coppi & Aharonian 1997; Murase et al. 2012). Otherwise, they cascade into ZeV photons when reaching the Galaxy.

5.2.4. Neutrinos

The most promising way to detect Planck accelerators is almost certainly neutrinos. These may be created in the collider itself through hadronic processes, or through the GZK process as free protons interact with the CMB. It is even possible that the neutrinos are the beam particles themselves. The extreme environments of accelerators might cool particles that produce neutrinos, but a small fraction of the power accelerated can still leak out as “prompt” neutrinos emitted by very quickly decaying hadrons with charm and beauty quarks (Enberg et al. 2009). In fact, neutrinos may be the only high energy particle that escapes the accelerator.

YeV–XeV neutrinos propagate freely through the Universe. Although they can interact with the CMB or the relic neutrino background (Reuel 1993), the Universe has an optical depth of only $\sim 0.001$ (Seckel 1998). They can also cascade in magnetic fields (Kuznetsov & Mikheev 1997), but the threshold for this process is

$$E_{\text{thresh}} = 150 \text{ VeV} \left( \frac{B}{6 \, \mu G} \right).$$

for electron neutrinos and even higher for muon and tau neutrinos (Bhattacharya & Sahu 2009). Even in the Earth’s magnetic field ($B \approx 1$ G), electron neutrinos are not affected unless they have energies of $\gtrsim 100$ XeV. In this case, the attenuation length is $2 \times 10^{15}$ cm ($B$/G)$^2(E$/100 XeV)$^{-1}$ (Erdas & Lissia 2003), much bigger than the Earth itself except at the most extreme energies.

5.3. Should we expect to see anything?

A Planck accelerator is literally a cosmic investment by anyone who builds it. The amount of effort required to build and operate one, and the rarity of Planck events, could push ETs to be as efficient as possible in extracting information from it. They may want to carefully study every outgoing particle made in the collisions. Would they let any escape so that we could detect them? There could also be environmental reasons to stop high energy radiation, to prevent it from harming life elsewhere in the host galaxy.

Our own accelerators frequently measure the energy of outgoing particles with calorimeters, which stop the particles entirely. One could imagine an ET somehow building calorimeters extreme enough to stop even neutrinos, or at least stopping mesons and leptons before they can decay into neutrinos. However, calorimeters are not necessary to characterize the collisions. Outgoing charged particles emit radiation and ionize matter, which can provide a signature of their presence (chapter 33 of Olive & Particle Data Group 2014 is an extensive review of particle detectors).

We have some reasons to hope that ETs would at least allow neutrinos to escape the accelerator:

- It takes an enormous column to stop a YeV–XeV neutrino, about $\sim 10^6$ g cm$^{-2}$, and Planck accelerators are necessarily big. A neutrino-stopping spherical shell with radius given by equation would have a mass of 0.005 $M_\odot$, and that is with $B_{\text{QED}}$ electromagnetic fields in the accelerator. A Planck accelerator can easily be the size of a Solar System, in which case a spherical neutrino-stopping shell requires $\gtrsim 10^5$ $M_\odot$ of material. Of course, these requirements are a lot less burdensome if outgoing neutrinos are beamed.

- The energy dumped in calorimeters does not simply disappear, but is dissipated. This presents a lot of practical challenges. There would be a large thermal background in the detectors from the heat generated, which could make detecting individual particles more difficult. Stopped particles ionize atoms in the detectors as well, causing damage. Finally, the calorimeters would become radioactive under the onslaught of radiation. Aside from the problem of disposing of cosmic amounts of radioactive waste, the high energy emission from radioactive nuclei can mask desired signals.

- A Planck accelerator also potentially serves as a YeV–XeV neutrino factory. The propagation of YeV–XeV neutrinos over interstellar distances sets very tight constraints on new physics like Lorentz Invariance Violation. These effects are expected to appear around the quantum gravity scale. Our own precision tests can only constrain low-order effects with weak energy dependence (e.g., Abdo et al. 2009), but this is not a problem for Planck energy neutrinos. If the neutrinos are released for experiments elsewhere in their host galaxies to detect and characterize, we too could detect them at extragalactic distances. Altruistic ETs might even build their Planck accelerator to be used as a cosmic neutrino factory observed throughout the Universe by other ETs.

- Most speculatively, the ETs might release YeV–XeV neutrinos to advertise their having a Planck accelerator (c.f., Swain 2006). Of course, this assumes that ETs want to communicate. But if ETs generally find cosmic engineering distasteful, a few may build Planck accelerators and then draw attention to it to dissuade others from building their own.

But ultimately if all radiation is trapped by the accelerator, then it will appear to us only as waste heat. Cosmic engineering can generally be detected by its waste heat in the infrared (Dyson 1964; Wright et al. 2014b). An artificial Planck accelerator may be much hotter than a structure like a Dyson sphere, though, since it is not primarily intended to be lived on. On the one hand, it could be a NIR-excess source, looking like an obscured AGN. In fact, it could actually be a NIR-excess source, looking like an obscured AGN. In fact, it could actually be a rare subtype of obscured AGN, especially if it uses black holes (c.f., Kardashev 1995; Thompson & Lacki 2011). But if it takes the form of a relativistic fireball (section 5.1), most of its thermal radiation could emerge as X-rays or soft $\gamma$-rays. These accelerators would be transient, appearing more like a GRB than anything else, and would be missed by our current surveys for cosmic engineering.

6. YeV NEUTRINO LIMITS

Methods for detecting high energy neutrinos are basically the same as for detecting other types of UHECRs. An UHE neutrino has some probability of colliding with the particles
High energy neutrinos are easier to detect than low energy neutrinos because the cross section for weak interactions is greater at larger energies. At most energies, neutrinos are more likely to interact with nuclei than electrons, with a nucleon-neutrino cross section of

\[ \sigma_{\nu N} \approx 2 \times 10^{-30} \text{ cm}^{-2} \left( \frac{E}{\text{YeV}} \right)^{0.363} \]  

from Gandhi et al. (1998). The typical column to stop an UHE neutrino is thus

\[ \Sigma_{\nu N} = \frac{m_p}{\sigma_{\nu N}} \approx 8 \times 10^5 \text{ g cm}^{-2} \left( \frac{E}{\text{YeV}} \right)^{-0.363}, \]  

which is equivalent to about 8 km of ice or water, or about 1 or 2 km of rock.

In addition to a greater interaction probability, higher energy neutrinos do interact produce a louder cascade signal. The neutrino emits a weak boson carrying about \( \sim \) 100 GeV of energy neutrinos that do interact produce a louder cascade signal. I therefore extend the number flux limits to 100 YeV–ZeV energies (Allison et al. 2012). The maximum energies \( E_{\text{best}} \) the bounds are given. Since \( E_{\text{best}} \) varies over many orders of magnitude, I estimate \( A \) for every experiment at 100 YeV by scaling with the neutrino-nucleon cross section energy dependence (Gandhi et al. 1998).

The highest sensitivity yet attained at YeV energies is by the Westerbork Synthesis Radio Telescope (WSRT) NuMoon experiment, which had an etendue of \( \sim 0.01 \text{ sr} \) at 100 YeV. Since NuMoon had an exposure of only 46.7 hours (Scholten et al. 2009, Buitink et al. 2010), though, the limits are actually quite weak at YeV energies (as seen in Figure 3). In the next 10 to 20 years, limits on YeV neutrinos will improve by a factor of 10 or more, when the Low Frequency Array (LOFAR; Buitink et al. 2013) and Square Kilometer Array (SKA; Bray et al. 2014) search the Moon for Cherenkov flashes. The best etendue that ground-based lunar Cherenkov experiments can reach is limited by the surface area of the Moon’s near side, \( a = 0.044 \text{ sr} \). Future experiments can detect radio pulses more deeply buried neutrino showers if they are more sensitive, but their etendue can be at most a factor of 10 greater than NuMoon. The main improvement from LOFAR and SKA at Planck energies is simply that they will observe the Moon for longer periods (perhaps 1000 hours; Bray et al. 2014).

The method for deriving limits on the flux from point sources is similar. The number of events expected in the experiment is then

\[ \langle N_{\text{events}} \rangle = E \frac{dN}{dE} A_{\text{eff}} t_{\text{exp}}, \]  

where \( t_{\text{exp}} \) is the livetime of the experiment. From the lack of events (\( \langle N_{\text{events}} \rangle \ll 1 \)) and the flux limits given by each experiment, I calculate their \( A \). These are listed in Table 5 for the maximum energies \( E_{\text{best}} \) the bounds are given. Since \( E_{\text{best}} \) varies over many orders of magnitude, I estimate \( A \) for every experiment at 100 YeV by scaling with the neutrino-nucleon cross section energy dependence (Gandhi et al. 1998).

The relevant quantity is the effective etendue \( A \) of the experiment. This is the product of the effective field of view \( \Omega_{\text{eff}} \) and the effective area \( a_{\text{eff}} \) of the detector (which is the actual area of the target scaled by the detector efficiency). The number of expected events is then

\[ \langle N_{\text{events}} \rangle = E \frac{dN}{dE} A_{\text{eff}} t_{\text{exp}}, \]  

where \( t_{\text{exp}} \) is the livetime of the experiment. From the lack of events (\( \langle N_{\text{events}} \rangle \ll 1 \)) and the flux limits given by each experiment, I calculate their \( A \). These are listed in Table 5 for the maximum energies \( E_{\text{best}} \) the bounds are given. Since \( E_{\text{best}} \) varies over many orders of magnitude, I estimate \( A \) for every experiment at 100 YeV by scaling with the neutrino-nucleon cross section energy dependence (Gandhi et al. 1998).

The method for deriving limits on the flux from point sources is similar. The number of events expected in the experiment is then

\[ \langle N_{\text{events}} \rangle = E \frac{dN}{dE} A_{\text{eff}} t_{\text{exp}}, \]  

where \( a_{\text{eff}} \) takes into account the position of the source on the sky. The sensitivity of neutrino experiments typically varies with direction, but order of magnitude estimates can be set by assuming a uniform sensitivity across the field of view, so that \( a_{\text{eff}} \approx A/\Omega_{\text{eff}} \). Experiments observing the Moon at low radio frequencies (~ 100 MHz) include WSRT and eventually LOFAR and SKA-Low. It is thought that these experiments are sensitive to neutrinos hitting the Moon’s surface coming from any inclination in the sky (Scholten et al. 2009), in which case \( \Omega_{\text{eff}} \approx \pi \text{ sr} \). Most of the other lunar Cherenkov experiments in Table 5 observe at GHz frequencies, in which case \( \Omega_{\text{eff}} \) is smaller. This makes them more sensitive to point sources than their lower \( A \) implies, if their field of view includes the source. LUNASKA, for example, is more sensitive to Centaurus A than NuMoon (James et al. 2011).
Table 5

| Experiment | Start Date | Epochs | $\Delta t$ (yr) | $t_{\text{exp}}$ (yr) | $E_{\text{best}}$ (YeV) | Best $E \frac{d\phi}{dE}$ limit (km$^2$ sr$^{-1}$ yr$^{-1}$) | $A_{\text{best}}$ (km$^2$ sr) | Extrapolated$A (100 \text{ YeV})$ (km$^2$ sr) |
|------------|------------|--------|----------------|-------------------------|---------------------|---------------------------------|------------------|----------------------------------|
| 1.3. ARA   |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. JEM-EUSO |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. LOFAR  |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. SKA1-Low |           |        |                |                         |                     |                                 |                  |                                  |
| 1.3. ANITA-II |           |        |                |                         |                     |                                 |                  |                                  |
| 1.3. RESUN   |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. SAUND II |           |        |                |                         |                     |                                 |                  |                                  |
| 1.3. FORTE   |            |        |                |                         |                     |                                 |                  |                                  |

Future experiments

| Experiment | Start Date | Epochs | $\Delta t$ (yr) | $t_{\text{exp}}$ (yr) | $E_{\text{best}}$ (YeV) | Best $E \frac{d\phi}{dE}$ limit (km$^2$ sr$^{-1}$ yr$^{-1}$) | $A_{\text{best}}$ (km$^2$ sr) | Extrapolated$A (100 \text{ YeV})$ (km$^2$ sr) |
|------------|------------|--------|----------------|-------------------------|---------------------|---------------------------------|------------------|----------------------------------|
| 1.3. ARA   |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. JEM-EUSO |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. LOFAR  |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. SKA1-Low |           |        |                |                         |                     |                                 |                  |                                  |
| 1.3. ANITA-II |           |        |                |                         |                     |                                 |                  |                                  |
| 1.3. RESUN   |            |        |                |                         |                     |                                 |                  |                                  |
| 1.3. SAUND II |           |        |                |                         |                     |                                 |                  |                                  |
| 1.3. FORTE   |            |        |                |                         |                     |                                 |                  |                                  |

I show the expected background for monoenergetic neutrino sources with various Kardashev ratings and various densities as blue lines in Figure 3. Current experiments imply that K3 neutrino sources are typically spaced $> 100$ Mpc apart, for all neutrino energies of $1 \text{ EeV} \lesssim E \lesssim 3 \text{ YeV}$. Evidently, K3 YeV accelerators are either very rare, totally opaque, highly beamed, or intermittent. In a typical 100 Mpc long cube, there are roughly $2 \times 10^{14} \text{ M}_\odot$ of stars, and 5000 galaxies with stellar masses bigger than $10^{10} \text{ M}_\odot$ Baldry et al. (2012) – about 10% of the number being studied by the G Infrared Search, which also has negative results Wright et al. (2014a). This supports the idea that K3 entities are uncommon.

6.3. Indirect limits on the isotropic steady-state neutrino background

The weak limits on the highest energy neutrinos stem from the tiny expected number fluxes, but this does not mean the energy flux is small. If a significant fraction of the energy in YeV–XeV neutrinos is converted into radiation of lower energy, it would be detectable with current instruments. This indirectly limits the cosmic YeV–XeV neutrino background.

The “natural” strength of the neutrino background is the Waxman-Bahcall bound, $E^2 \frac{d\phi}{dE} = 2.7 \times 10^{-8} \text{ GeV cm}^2 \text{ s}^{-1} \text{ sr}^{-1}$ Waxman & Bahcall (1999); after including all flavors and supposing there is no evolution; light grey shading in Figure 3. It is essentially the flux in observed UHECRs. The reasoning behind this bound is that if high energy neutrinos are accelerated in the same sources as UHECRs, the cosmic luminosity in neutrinos is at most the luminosity accelerated in UHECRs, assuming that the sources are transparent to UHECRs. While its assumptions may not apply to artificial accelerators, it does set an upper limit on the natural background of neutrinos. In fact, the background of PeV neutrinos observed by IceCube roughly lies on the Waxman-Bahcall bound Aartsen et al.
Figure 3. Current limits on the isotropic, steady-state background of YeV neutrinos at Earth. The solid black lines are extant limits, while the grey solid lines are the projected limits from future experiments (references listed in Table 5). I extrapolate the limits from WSRT and LOFAR to XeV energies assuming a constant sensitivity (dashed lines) or as $A \propto E^{-0.36}$ from the neutrino-nucleon cross section dependence (dotted lines). Indirect limits on the cosmic background include constraints on emission of W-bursts (darker grey shading), and the Waxman-Bahcall limit (light grey shading; no evolution). The indirect limits only apply to extragalactic backgrounds. I also show the expected neutrino backgrounds from various extragalactic (blue) and Galactic (red) monoenergetic neutrino sources with various space densities and Kardashev ratings. The point source fluxes are converted into effective isotropic backgrounds using $\Omega_{\text{eff}} = \pi \text{sr}$.

The Waxman-Bahcall bound implies a small density of neutrino sources, no matter the neutrino energy (c.f. equation (33)):

$$n_{\text{source}} \lesssim 4 \times 10^{-8} \text{ Mpc}^{-3} \left( \frac{L_{\text{source}}}{10^{10} \text{ L}_\odot} \right)^{-1} \left( \frac{t_{\text{source}}}{10 \text{ Gyr}} \right)^{-1}. \quad (34)$$

This corresponds to a spacing of $\lesssim 280$ Mpc between K3 neutrino sources (a volume that includes $5 \times 10^{15} \text{ M}_\odot$ of stars, and about $10^5$ galaxies with $> 10^{10} \text{ M}_\odot$ of stars), or $\gtrsim 6$ Mpc between K2.5 neutrino sources. Shock CR acceleration produces particle spectra with a $dN/dE \propto E^{-2}$ or steeper spectra. If this is how an artificial accelerator works, then it is necessarily brighter at the well-constrained PeV–ZeV energies than at YeV–XeV energies. However, the Waxman-Bahcall bound does not apply if neutrino sources are monoenergetic.

A more robust limit can be set by the interaction of YeV–
XeV neutrinos with the CMB. Although the optical depth of the Universe to $\nu\gamma$ interactions is only $\sim 7 \times 10^{-4}$, it is not zero (Seckel 1998). A small but potentially observable fraction of cosmic YeV–XeV neutrinos is converted into W bosons during their propagation to Earth. The W bosons produce hadrons and leptons (a W-burst), with about half of the energy ultimately cascading down into GeV $\gamma$-rays.

The observed GeV $\gamma$-ray background (Aabu et al. 2010; Ackermann et al. 2015) thus limits the cosmic energy injection into neutrinos even at these energies (Coppi & Aharonian 1999; Murase et al. 2012). I calculate this limit on a neutrino background at YeV energies and plot it in Figure 3, including both the CMB and the extragalactic background light presented in (Franceschini et al. 2008). Numerically, the $\nu\gamma$ bound corresponds to an energy flux of $E^2d\phi/dE \approx 2 \times 10^{-5}$ GeV cm$^{-2}$ s$^{-1}$ sr$^{-1}$ at all energies greater than YeV. The resulting limit on the density of K3 neutrino sources is

$$n_{\text{source}} \lesssim 4 \times 10^{-5} \text{Mpc}^{-3} \left(\frac{L_{\text{source}}}{10^{10} \text{L}_{\odot}}\right)^{-1} \left(\frac{t_{\text{source}}}{10 \text{ Gyr}}\right)^{-1},$$

or a spacing of 30 Mpc. While not as strong as the direct limits at energies below 1 YeV, it does confirm that steady, isotropically-radiating YeV–XeV neutrino sources are very rare, occurring in less than 1% of massive galaxies (less than 1 per $\sim 6 \times 10^{12} \text{M}_{\odot}$ of stars). The limit could probably be improved by considering the interaction of YeV–XeV neutrinos with the relic cosmic neutrino background (the Z-burst mechanism; Fargion et al. 1993; Weiss 1999), but this requires knowledge of the neutrino masses (Roulet 1999), so I do not include it.

Note that these limits only apply to the cosmic averaged emission of neutrinos. They do not exclude any given individual source close to or within the Galaxy. Most of the constraining power comes from galaxies that are several Gpc away.

Individual YeV neutrino sources might be visible through the W-burst mechanism in GeV $\gamma$-rays, even if they have no intrinsic electromagnetic emission. Note that more distant sources would have a brighter apparent $\gamma$-ray luminosity, because the optical depth for neutrinos along the sightline is larger (c.f., Essey et al. 2011). In practice, though, the expected flux even for a K3 YeV neutrino source is too small to see with current GeV telescopes like Fermi-LAT.

### 6.4. Limits on YeV neutrino bursts

Planck accelerators may not be steady sources, but could be brief transients as in the fireball scenario (section 5.1). The assumption of steady sources applies if the lifetime of the sources $\Delta t$ is longer than the exposure time $t_{\text{exp}}$ and if there is at least one source in the sky at any given time. For comparison, typical neutrino experiments last for a few weeks to a few years (Table 5).

There are three conditions for detecting the YeV neutrinos from a transient event. The first is that a detection experiment must be active when the transient occurs. The second is that the experiment must be pointed at the transient when it occurs. The third is that the fluence of neutrinos must be big enough that the experiment is likely to detect at least one during the event. Suppose an experiment is sensitive to bursts of a given luminosity out to a distance $D_{\text{max}}$. The number of transients it observes is

$$N_{\text{events}} = \Gamma_{\text{burst}} t_{\text{exp}} \Omega_{\text{eff}} D_{\text{max}}^2 / 3,$$

where $\Gamma_{\text{burst}}$ is the rate of the events per unit volume, $t_{\text{exp}}$ is the exposure/active time of the experiment, and $\Omega_{\text{eff}}$ is the effective field of view of an experiment. This equation applies if the bursts are short, with $\Delta t \ll t_{\text{exp}}$.

Longer-lasting transients are more likely to be active when the experiment is on, but they are harder to detect since only a fraction of the neutrino fluence can be detected. If $\Delta t \gg t_{\text{exp}}$, the number of transients observed can be approximated by replacing $t_{\text{exp}}$ with $\Delta t$ in equation (36) and adjusting $D_{\text{max}}$ to reflect the new sensitivity. The observed fluence is the total fluence multiplied by a factor $t_{\text{exp}} / (\Delta t)$, however.

The number of neutrinos generated can be related to the total integrated energy released during the event. As shown in Table 2, a particle accelerator that can set useful constraints on Planck physics would generally need to convert $10^8$–100 $M_{\odot} c^2$ of energy into YeV–XeV particles. From energy conservation, $10^{40}$ 100 YeV neutrinos can be generated from each $M_{\odot} c^2$ of energy in accelerated particles. The fluence of an isotropic burst is

$$\Phi = \frac{120}{\pi} \left(\frac{\nu}{M_{\odot} c^2}\right) \left(\frac{E}{100 \text{ YeV}}\right)^{-1} \left(\frac{D}{\text{Gpc}}\right)^{-2}.$$  

Current limits on neutrino bursts can therefore be calculated from the effective areas of neutrino experiments. From Table 2, UHE neutrino experiments range from those with huge effective areas but short exposure times (like NuMoon) to those with small effective areas but years of exposure (like RICE). Bursts can be detected at distances of $D_{\text{max}} \lesssim [(\nu c/\Delta \Omega) / (4\pi c^2)]^{1/2}$. In the sensitive/short exposure extreme, a 100 $M_{\odot} c^2$ burst could have been detected by NuMoon out to 9 Gpc ($E_{\nu}/100$ YeV)$^{-1}$, ignoring cosmological effects. A minimal $10^{-8}$ $M_{\odot} c^2$ burst from a neutrino-neutrino accelerator (Table 2) would have been detectable from a distance of 90 kpc ($E_{\nu}/100$ YeV)$^{-1}$, which is far enough to encompass the entire Galaxy and the Magellanic Clouds for 1 YeV $\lesssim E_{\nu} \lesssim 300$ YeV. RICE, the longest lasting YeV neutrino experiment, could only have detected a 100 $M_{\odot} c^2$ burst out to 3 Mpc ($E_{\nu}/100$ YeV)$^{-1}$, only $10^{-10}$ of the volume probed by NuMoon.

In order to limit the rate at which short bursts occur per galaxy, we want to maximize the product of the probed volume ($\Omega_{\text{eff}} D_{\text{max}}^3$) and the exposure time. A useful figure-of-merit is $M = \Omega_{\text{eff}} t_{\text{exp}}$. I computed $M$ for the experiments listed in Table 2 using the extrapolated $A$ at 100 YeV and with $\Omega_{\text{eff}} = \pi$. NuMoon has the largest $M$ by far, $5 \times 10^7$ km$^3$ yr. Of course, the range of NuMoon is so far that cosmological effects matter, so this cannot be taken as exact. Yet, only FORTE comes anywhere near NuMoon, with $M = 7 \times 10^5$ km$^3$ yr. ANITA, the third best, has $M = 8 \times 10^4$ km$^3$ yr; RICE has a mere $M = 400$ km$^3$ yr. Larger effective areas are key.

LOFAR and SKA will have effective areas somewhat better than NuMoon and will run for a week to a month, with $M \approx 10^9$–$10^{10}$ km$^3$ yr. For comparison, if a burst occurred in every $L_*$ galaxy once in the past 10 Gyr, then there would be about 1 burst per year in the Hubble volume. We are still far from ruling out a hadronic Planck accelerator being present in every galaxy, if they are run in short bursts.

### 6.5. Towards 1 yr exposures: Earth’s oceans as an experimental volume
Because YeV–XeV neutrinos are so rare, the Moon simply is not a big enough detector to observe distant colliders. The only other planetary-sized body we can observe UHE neutrinos signals from the ground is the Earth itself. About 70% of Earth’s surface is covered by the oceans, and their average depth is of order a few kilometers. The columns of water in the oceans are big enough to stop YeV–XeV neutrinos, so the oceans are a good candidate experimental volume.

There are at least three potential signals from neutrino cascades in seawater. The first is the radio Cherenkov pulses, as observed on the Moon. But since seawater is highly conductive, radio emission is absorbed rapidly in the oceans, so these are undetectable. The second is the optical Cherenkov flash, sought by the ANITA neutrino telescope. These experiments seek TeV–PeV neutrinos, though, and use small experimental volumes with closely spaced photodetectors (Ageron et al. 2011). They cannot be readily extended to YeV–XeV energies because the absorption length of optical light is a few hundred meters (Águilera et al. 2005). Thus, the photodetectors cannot be more than \( \sim 1 \) km apart. For the cubic kilometer volumes of PeV neutrino telescopes, this is not a problem, but for all of the oceans, one would need billions of photodetectors.

The third option is to listen for the sound of a neutrino cascade (Urick 1983). A 100 YeV neutrino dumps about 3 megaloules of energy (about 1 kg TNT equivalent) into a cubic meter of volume, assuming 20% of its energy goes into a hadronic cascade (Nies & Bertin 2006). This heats the water, causing it to expand, and launches a pressure wave (Askar van Dolsgoheit 1977; Learned 1979; Lehnen et al. 2002). The characteristic frequencies are of order 10 kHz. Acoustic experiments search for these sound pulses; the largest so far is SAUND (Kurahashi et al. 2010). The conversion of heat into sound is very inefficient (Learned 1979; Price 1996), and the sound is beamed (Lehnen et al. 2002), partly accounting for why the etendue of acoustic detectors are so far weak. Present efforts are directed towards finding 100 EeV GZK neutrinos, whereas Planck neutrinos will be proportionally much louder, compensating for the low efficiency.

Suppose that a fraction \( \eta_{\text{sound}} \) of the energy in the neutrino is emitted as a sound pulse with frequency \( \nu_{\text{pulse}} \). The power of the pulse is approximately \( \eta_{\text{sound}} \nu_{\text{pulse}} \). The sound is not emitted isotropically, but is concentrated in a disk (or “pancake”) expanding perpendicularly from the neutrino cascade (Lehnen et al. 2002; Vandenbroucke et al. 2005). The anisotropy can be described by an angular distribution \( dI/d\Omega \), normalized so that \( \int (dI/d\Omega) d\Omega = 1 \). For hadronic showers, the emission disk has an angular thickness of order \( \Delta \theta \approx 1^\circ \) (Lehnen et al. 2002), so \( dI/d\Omega \approx 100 \) within the disk. In addition, seawater absorbs sound with an energy attenuation length \( D_{\text{abs}} \). Then the energy flux at a distance \( D \) in the water and direction \( \mathbf{k} \) is

\[
I = \frac{dI(k)/d\Omega \eta_{\text{sound}} \nu_{\text{pulse}} \rho c_{\text{sound}}}{2\pi D^{2}} \exp(-D/D_{\text{abs}})^{2}.
\]

The pressure difference is related to the energy flux as

\[
\Delta P = \frac{dI(k) \eta_{\text{sound}} \nu_{\text{pulse}} \rho c_{\text{sound}}}{4\pi D^{2}} \exp\left(-\frac{D}{D_{\text{abs}}}\right)^{1/2}.
\]

For a 100 YeV neutrino detected at a distance, the pressure wave has an amplitude of roughly

\[
\Delta P = 1.4 \text{ Pa} \left(\frac{D}{100 \text{ km}}\right)^{-1} \exp\left(-\frac{D}{2D_{\text{abs}}}\right) \times \left(\frac{dI(k)/d\Omega \eta_{\text{sound}}^{2} \nu_{\text{pulse}}^{2}}{100 \text{ YeV}} \frac{E}{10 \text{ kHz}}\right)^{1/2}.
\]

I estimate \( \eta_{\text{sound}} \) from the pressure profiles given in Lehnen et al. (2002). An acoustic array operating at these frequencies has an expected pressure sensitivity of 1 mPa (Lehnen et al. 2002) to 0.1 Pa (Kurahashi et al. 2010).

The first limiting factor for an acoustic array is sound absorption in seawater, since the energy attenuation lengths are a few kilometers or less above 10 kHz (Fisher & Simmons 1977; Francois & Garrison 1982; Ainslie & McCollm 1998). Under some standard assumptions about ocean temperature (\( T = 2^\circ \)C), salinity (\( S = 0.035 \)), acidity (\( \text{pH} = 8 \)), and depth (\( z = 5 \) km), the ocean attenuates sound at a rate

\[
\frac{\alpha_{\text{sound}}}{\text{dB/km}} 
\begin{cases} 
0.13 \nu_{\text{kHz}}^{2} & (0.2 \text{ kHz} \lesssim \nu_{\text{kHz}} \lesssim 0.84) \\
0.089 + 0.0053 \nu_{\text{kHz}}^{2} & (0.84 \lesssim \nu_{\text{kHz}} \lesssim 45) \\
11 + 3.4 \times 10^{-4} \nu_{\text{kHz}}^{2} & (45 \lesssim \nu_{\text{kHz}})
\end{cases}
\]

according to Ainslie & McCollm (1998), where \( \nu_{\text{kHz}} \) is the sound frequency in kHz. This can be translated into an energy attenuation length:

\[
D_{\text{abs}} = 10 \text{ dB}/(\alpha_{\text{sound}} \ln 10) = 4.3 \text{ km} \left(\frac{\alpha_{\text{sound}}}{\text{dB/km}}\right)^{-1}.\]

At 10 kHz, the attenuation length is about 7 kilometers. Thus, the exponential factor in equation (39) is 0.03 for \( D = 50 \) km, \( 8 \times 10^{-4} \) for \( D = 100 \) km, and \( 2 \times 10^{-5} \) for \( D = 150 \) km. Despite the enormous losses, the pulse from a 100 YeV neutrino is so loud that it could be detected \( D_{\text{range}} \approx 45 \) km away by an array sensitive to 0.01 Pa sounds.

The other limitation is the narrowness of the acoustic disk. If a neutrino comes from the zenith, the acoustic disk expands outwards into the sea without interacting with the seafloor and sea surface. But if a neutrino comes from the horizon, most of the sound is directed towards the seafloor and sea surface. A neutrino must have a zenith angle of \( \lesssim z/D_{\text{range}} \) to be regarded as horizontal, where \( z \) is the ocean depth. For \( z \approx 5 \) km, this limits the zenith angle to about \( 6^\circ \), or only 0.3% of the celestial sphere. A completely vertical neutrino illuminates only \( \Delta \theta \) in azimuth out to large distances.

If we were only interested in detecting zenith neutrinos, then we could settle for a few detectors per \( D_{\text{range}} \). To cover the Earth’s oceans would require \( A_{\text{Earth}}/D_{\text{range}}^{2} \approx 2 \times 10^{5} \) detectors. For comparison, the SAUND II array used only 49 hydrophones (Kurahashi et al. 2010). Furthermore, these detectors would actually be strings of hydrophones, to ensure that one intersects the acoustic disk (Vandenbroucke et al. 2005). But since the distance between the strings is big, the hydrophones can be spaced about one per kilometer on the strings.

The effective etendue of an ocean array that detected only zenith neutrinos would actually be less than NuMoon, although a hydrophone array that covered the Earth could observe in all directions and have a longer livetime. To effectively detect horizon neutrinos on-disk, the number of detec-
tor strings must be multiplied by $2\pi/\Delta \theta \approx 400$, for a grand total of $10^8$ strings.

A much more promising route is to consider the off-disk acoustic emission from a shower. Because of the exponential attenuation factor, the range only depends logarithmically on $dI/d\Omega$. According to Lehtinen et al. (2002), the sound fluence of a hadronic shower $5^\circ$ away from the disk is about 0.1% of maximum. Unfortunately, the sound emission at larger angles is not given. As a fiducial estimate, suppose $dI/d\Omega = 10^{-3}$. The pressure differential then remains above 1 Pa out to $\sim 15$ km for a 100 YeV neutrino (0.01 Pa sensitivity), increasing the number of detectors by a factor of $\sim 10$ only. For a Planck neutrino, the range increases to $\sim 30$ km (0.01 Pa sensitivity). Furthermore, the detectors would not have to be strings, since they would be listening for the isotropic sound emission.

It may be worthwhile to study the low frequency component of these pulses with care. Sounds below a kHz can propagate thousands of kilometers, so a very sparse array might be able to detect the cascades of the highest energy neutrinos.

6.6. The far future of YeV neutrino detection: how big can we go?

It would take exposures of Earth-centuries or millennia to rule out K3 YeV accelerators anywhere in the Universe, much less those with the minimal luminosities listed in Table 2. But must we limit ourselves to Earth, if we are going to extrapolate that far into the future?

6.6.1. Jupiter (a bad idea)

Jupiter has a surface area that is $\sim 100$ times bigger than that of Earth. Could we deploy acoustic detectors throughout Jupiter’s atmosphere to listen for YeV neutrinos? I will be fanciful for a moment about our capabilities this far in the future and ignore the logistics. Maybe the detectors can be quickly deployed by launching them on self-replicating machines, creating an ecology of “floaters” once speculated to live in Jupiter (Sagan & Salpeter 1976).

The basic challenge facing any such attempt at instrument creation is that the density of an atmosphere is low while the column density to stop a YeV neutrino is relatively high (equation 29). The outer thousand kilometers of Jupiter’s atmosphere is approximately adiabatic with $P = P_0 (\rho / \rho_0)^{5/3}$ (Seiff et al. 1996). From the equation of hydrostatic stability, $dP/\rho \sim \rho g$, the density in Jupiter’s atmosphere at depth $z$ is

$$\rho = \left[ \frac{2g\rho_0}{5P_0} \right]^{3/2} \rho_0.$$  

The column density of depth $z$ is given by $d\Sigma/dz = \rho$, so that $\Sigma = (2/5)\rho z$. Define as a reference “surface” the 1 bar depth of Jupiter’s atmosphere, where $P_0 = 10^6$ dyn cm$^{-2}$, $\rho_0 = 0.0002$ g cm$^{-3}$, and the temperature is $T_0 = 170$ K (Seiff et al. 1996). The gravitational acceleration in the outer parts of Jupiter is $g = 2500$ cm s$^{-2}$. Neutrinos interact at a typical depth of

$$z_\nu = \frac{5}{2\rho_0 g} \left( \frac{P_0}{\rho} \right)^{3/2} \Sigma_{\nu/5}^{2/5} = 1100 \text{ km} \left( \frac{\Sigma_{\nu/5}}{10^6 \text{ g cm}^{-2}} \right)^{2/5},$$

where the density is $\rho_\nu = 0.022$ g cm$^{-3} (\Sigma_{\nu}/10^6 \text{ g cm}^{-2})^{3/5}$ and the pressure is $P_\nu = 2500$ bar ($\Sigma_{\nu}/10^6 \text{ g cm}^{-2}$). The temperatures at these depths are

$$T_\nu = T_0 \left( \frac{P_\nu}{P_0} \right)^{2/5} = 3900 \text{ K} \left( \frac{\Sigma_{\nu}}{10^6 \text{ g cm}^{-2}} \right)^{2/5}.$$  

Jupiter’s atmosphere may be nonadiabatic at depths greater than 1000 bar, and the temperature could be somewhat lower (Guillot et al. 1994), but in any case, $z_\nu$ is extremely deep.

These physical conditions present enormous challenges to detecting neutrinos. The pressure is a few times that in Earth’s deepest ocean trenches, but the temperature is far beyond even the surface of Venus. No known electronics can withstand $\gg 1000$ K temperatures for any significant time; indeed, most solid materials melt at these depths. Even if the material engineering challenges could be overcome, the high temperatures lead to a huge thermal background of noise that would cover up neutrino signals. Similar considerations apply for the other gas giants, not to mention the Sun.

Gaseous atmospheres are a poor choice for a YeV neutrino detector target because they are too compressible. The pressure in an incompressible liquid ocean made of water builds up rapidly, as its density does not change with depth. The pressure and density of a liquid is basically independent of its temperature. But a gas can support the increased pressure load by becoming hotter without becoming much denser. Because of the much lower gas densities, neutrino absorption occurs very deep in the atmosphere, where it is extremely hot.

6.6.2. A million Kuiper Belt Objects (a much better idea)

The minor bodies of the Solar System collectively have more surface area than the major bodies. Moreover, they are solid, so that $10^6$ g cm$^{-2}$ columns are achieved in all minor bodies with radii $R \geq 10$ km. Detecting neutrinos in these icy objects does not even require landing on them. An orbiting satellite can detect radio Cherenkov pulses from neutrino showers within them, as proposed for Europa (Shoji et al. 2011; Miller et al. 2012) and Enceladus (Shoji et al. 2012).

In terms of mass, the Kuiper Belt is the largest reservoir of minor bodies within 100 AU. Collectively, they have a large surface area. The radius distribution of Kuiper Belt objects (KBOs) is

$$\frac{dN_{\text{KBO}}}{dR} \approx 710 \text{ km}^{-1} \left( \frac{R}{45 \text{ km}} \right)^{-4},$$

for $0.25 \text{ km} \leq R \leq 45 \text{ km}$, as determined by stellar occultations (Schlichting et al. 2012). This integrates to a collective surface area of

$$a_{\text{eff}} = \int_{10 \text{ km}}^{45 \text{ km}} \frac{dN_{\text{KBO}}}{dR} 4\pi R^2 dR = 2.8 \times 10^9 \text{ km}^2 = 5.6 \times 10^9 \text{ km}^2,$$

for KBOs with $10 \text{ km} \leq R \leq 45 \text{ km}$. Smaller objects have an exposure penalty, since the probability for neutrino interaction goes as $R$ when the column depth is less than $\Sigma_{\nu}$: $a_{\text{eff}} \approx \int 4\pi R^2 (R/10 \text{ km}) (dN_{\text{KBO}}/dR) dR \propto \ln R$. I find that there is roughly 16 $\oplus$ in surface area on KBOs with $1 \text{ km} \leq R \leq 10 \text{ km}$, and another 16 $\oplus$ on KBOs with $0.1 \text{ km} \leq R \leq 10 \text{ km}$. Far greater populations may exist in the Oort cloud, although they are much harder to reach.

A future civilization may instrument the larger KBOs to detect YeV neutrinos, attaining an exposure rate about $\sim 10$ times greater than that from the Earth alone. If this becomes practical within a few centuries, the total integrated exposure to YeV neutrinos could be increased rapidly with a few
decades of observation time. The main practical challenge is the sheer number of objects that would need to be observed. There are roughly

$$N_{\text{KBO}} = \int_{0}^{45 \text{ km}} \frac{dN_{\text{KBO}}}{dR} dR \approx 1 \times 10^6$$  \hspace{1cm} (45)$$

KBOs with radii between 10 and 45 km, and about $10^9$ with radii between 1 km and 10 km. Merely finding all of these objects, much less visiting them all, is a formidable challenge.

Unlike the case of Jupiter, no miraculous advances in material science are needed, though.

7. DISCUSSION

7.1. Summary

Do some aliens leave their home planets and go on to cosmic engineering, do they all die before they can try, or do they all decide to stay at home? The Fermi Paradox implies the first possibility is always wrong if ETs exist in any numbers. Otherwise it is evidence that ETs are very rare. One issue with the argument is that most of the proposed reasons for cosmic engineering involve either contacting young, human-like societies across the Universe, or consuming cosmic amounts of energy. Remaking stars and galaxies to get our attention can seem a bit extravagant to get our attention, and consuming everything within reach is a questionable project if ETs care about sustainability and avoiding aggression with other ETs.

But I suggest there’s at least one more motivation, curiosity: some scientific problems can only be answered with cosmic engineering.

The Nightmare of particle physics is the dream of astronomers searching for ETs. Planck accelerators must be bigger than the Sun, for realistic electromagnetic field strengths (Section 3). The power used by a Planck accelerator is cosmic, because the natural cross section of Planck physics is tiny, with only one Planck event occurring per $\sim 10^{45}$ collisions (Section 4). The luminosities necessary for a hadronic collider in particular requires a K3 entity, if the experiment is to be completed within a million years (Table 2). Whether or not they feel like communicating, we can still look for these kinds of energy usages.

The conditions in the accelerator region may be very hostile if it is a dirty fireball with a compactness problem, thermalizing all of the high energy particles. Mesons and muons can be cooled before they decay into neutrinos, although we can still hope to see prompt neutrinos from kaons, charm, and beauty. Alternatively, YeV-XeV radiation generated by the collisions could escape if the collider is clean or highly relativistic (Section 5.1). Cosmic ray nucleons, photons, and $e^\pm$ can be detected from tens of Mpc away if they escape, and neutrinos are not stopped at all (Section 5.2). If nothing else, we can find Planck accelerators by their thermal emission, although it may emerge at much higher energies than the mid-infrared wavelengths of Dyson spheres (Section 5.3).

Our current knowledge of the high energy neutrino backgrounds implies that K3 YeV neutrino sources are typically at least 30 Mpc apart, if they are steady and radiate isotropically (Sections 6.2 and 6.3). Individual K3 accelerators can only be detected out to about 2 Mpc. In addition, there is no K2.5 or brighter 100 YeV neutrino source within about 10 kpc, implying there are no active Planck accelerators in the Galaxy (Section 6.2). LOFAR searches for radio Cherenkov pulses on the Moon can improve YeV neutrino exposures by a factor of ten to a hundred, mainly because of a longer exposure time (Section 6.1). If we could place hydrophone arrays throughout the Earth’s oceans and listen for cascade acoustic signals for a year, we could improve the bounds by another factor of a thousand. There are many practical difficulties, though, particularly the $10^5$ hydrophones required and the beaming of the acoustic signals (Section 6.3). Exposures much beyond an Earth-year might be achieved in the far future by searching with satellites for radio Cherenkov pulses in all $10^6$ Kuiper Belt objects wider than 20 km (Section 6.6).

7.2. Where is their nonthermal emission?

The Universe emits very little nonthermal emission – the lack of observed YeV neutrinos is just one symptom of this fact. If the Universe were filled with diverse cosmic engineers with varying motives, one might figure that there would be artificial emission across all energies and messengers. Some of their projects might generate radio broadcasts; others might generate $\gamma$-ray emission; others still neutrinos or gravitational waves – a sort of Copernican principle for messengers.

But almost all of the luminosity of the Universe is thermal, with $\nu g_{\text{BH}} \approx 3 \times 10^{-14}$ erg cm$^{-3}$ in the infrared to ultraviolet bands (Franceschini et al. 2008), and a similar density in thermal MeV neutrinos emitted during core-collapse supernovae (Beacom 2010). The characteristic energy density of UHECRs and neutrinos, the Waxman-Bahcall limit, is a million times smaller (Waxman & Bahcall 1999). Likewise, the GeV $\gamma$-ray background has a density that is $\sim 10^{-5}$ that of the EBL (Ackermann et al. 2015). While the intensity of the extragalactic GHz radio background is disputed, it is at most $4 \times 10^{-6}$ the density of the EBL (Fixsen et al. 2011) and perhaps a factor of 10 below that. At lower frequencies, the constraints just become even tighter, with $1/\nu_{\text{IC}} < \nu^{-0.3}$ (Fixsen et al. 2011).

While it is true that most of the power used in cosmic engineering must be thermal emission, if even one in a thousand galaxies had a K3 entity that emitted 1% of its power in one of these bands, they would overproduce the observed background radiation. The problem is especially severe at radio wavelengths, where we can now observe star-forming galaxies out to high redshift. Yet most star-forming galaxies appear to lie on the far-infrared–radio correlation, indicating natural synchrotron and free-free radio emission (Yun et al. 2001; Mao et al. 2011). K3 radio emitters would make these galaxies brighter at radio wavelengths and easier to see, but we have not found any artificial extragalactic radio (or $\gamma$-ray or neutrino) sources. Cosmic engineering’s effect on the Universe appears to be utterly minuscule, and the Fermi Paradox remains mystifying.

7.3. Neutrino SETI in context

Although searches for UHECRs and neutrinos of all energies from ETs have been proposed in the past (Subotowicz 1979; Learned 1994; Swain 2006; Silagadze 2008; Learned et al. 2009), there has not been a dedicated effort to do them, unlike in radio and optical. There are crucial differences between neutrino and radio/optical astronomy that are relevant. Radio and optical telescopes use focusing optics. Their fields of view are small, so they must be aimed, except for something like the Fast Fourier Transform Telescope (Reegmark & Zaldarriaga 2009). In addition, individual photons are easy to detect. The backgrounds from natural sources and the Galaxy mean that one must carefully search for desired signals. A radio/optical telescope cannot
be simply turned on and left to search the entire sky for any signal from all sources. Observations need specific goals in these wavebands. If SETI is not an explicit goal of the observers, one must settle for a “commensal” program, where one searches for artificial signals within the field of view of the primary targets, hoping that ETs happen to lie in those directions (Bowyer et al. 1983; Siemion et al. 2014).

In contrast, neutrino telescopes cannot be aimed. While the geometry of the Earth or the Moon can target some of the sky, their fields of view are nonetheless vast. We cannot focus high energy neutrinos; nor can we move cubic kilometers of ice or water or regolith, much less entire planets. Thus, there is no real way to be specific about the targets of a neutrino telescope. The main challenge instead is simply detecting individual neutrinos, just because they are so rare, and figuring out where they come from. Any experiment looking for high energy neutrinos is a SETI experiment, since every operating program is commensal with every other for high energy neutrinos. As experiments across the neutrino spectrum, from TeV to YeV, reach the astrophysically motivated energy neutrinos. As experiments across the neutrino observing program is commensal with every other for high energy neutrinos.

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